First Science with JouFLU

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FIRST SCIENCE WITH JouFLU

by

NICHOLAS JON SCOTT

Under the Direction of Harold A. McAlister

ABSTRACT

Jouvence of FLUOR (JouFLU) is a major overhaul of the FLUOR (Fiber Linked Unit for Optical Recombination) beam combiner built by the Laboratoire d'études spatiales et d'instrumentation en astrophysique (LESIA) and installed at the CHARA Array. These upgrades improve the precision, observing efficiency, throughput, and integration of FLUOR with the CHARA Array as well as introduce new modes of operation to this high-precision instrument for interferometry. Such high precision observations with FLUOR have provided the first unambiguous detections of hot dust around main sequence stars, showing an unexpectedly dense population of (sub)micrometer dust grains close to their sublimation temperature, 1400 K. Competing models exist to explain the persistence of this dust; some of which suggest that dust production is a punctuated and chaotic process fueled by asteroid collisions and comet infall that would show variability on timescales of a few years. By re-observing stars from the exozodiacal disks survey we have searched for variations in the detected disks. We have found evidence that for some stars the amount of circumstellar flux from these previously detected exozodiacal disks,
or exozodis, has varied. The flux from some exozodis has increased, for some the flux has decreased, and for a few the amount has remained constant. These results are intriguing and will be no doubt useful for future modeling of this phenomenon. Furthermore, long-term monitoring is suggested for some of these objects to confirm detections and determine the rate of variation.

INDEX WORDS: Interferometry, Interferometric techniques, high angular resolution, CHARA Array, FLUOR, JouFLU, debris disks, exozodiacal light, instrumentation, fiber interferometry, long baseline interferometry
FIRST SCIENCE WITH JouFLU

by

NICHOLAS JON SCOTT

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Georgia State University

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FIRST SCIENCE WITH JouFLU

by

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December 2015
I'd like to dedicate this work to friends and family who supported my mental state and its absence during its completion. I feel very privileged to get to know and work with some of the most talented people I've ever met. I don't know why anyone would ever do anything else... unless they liked sleep... or sanity...

Finally, I'd like to make special mention of Robin for her unwaivering support, love, and insight. I can not imagine a better companion in all adventures.
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- Paul Nuñez, NASA/JPL

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FLUOR has made use of NASA’s Astrophysics Data System and of the SIMBAD database, operated at CDS (Strasbourg, France). Exozodi survey extension and variability studies funded by NASA planetary science division grant NNN13D460T, under the Origins of Solar Systems (now Exoplanet Research) program.
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<tbody>
<tr>
<td>AAVSO</td>
<td>American Association of Variable Star Observers</td>
</tr>
<tr>
<td>ADU</td>
<td>Analog to Digital Unit</td>
</tr>
<tr>
<td>ALIU</td>
<td>ALIgnment Units</td>
</tr>
<tr>
<td>AO</td>
<td>Adaptive Optics</td>
</tr>
<tr>
<td>AR</td>
<td>Anti-Reflective</td>
</tr>
<tr>
<td>AT</td>
<td>Auxiliary Telescope</td>
</tr>
<tr>
<td>AU</td>
<td>Astronomical Unit</td>
</tr>
<tr>
<td>BCL</td>
<td>Beam Combination Laboratory</td>
</tr>
<tr>
<td>BL</td>
<td>BackLight</td>
</tr>
<tr>
<td>BRT</td>
<td>Beam Reducing Telescope</td>
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<tr>
<td>BSF</td>
<td>Beam Synthesis Facility</td>
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<td>Abbreviation</td>
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<tr>
<td>BS</td>
<td>Beam Splitter</td>
</tr>
<tr>
<td>CALI</td>
<td>the PICNIC-based science camera for JouFLU</td>
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<tr>
<td>CFHT</td>
<td>Canada France Hawaii Telescope</td>
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<tr>
<td>CHAMP</td>
<td>CHARA Michigan phase-tracker</td>
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<tr>
<td>CHARA</td>
<td>Center for High Angular Resolution Astronomy</td>
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<tr>
<td>CMOS</td>
<td>Complementary Metal Oxide Semiconductor</td>
</tr>
<tr>
<td>COAST</td>
<td>Cambridge Optical Aperture Synthesis Telescope</td>
</tr>
<tr>
<td>CoG</td>
<td>Center of Gravity</td>
</tr>
<tr>
<td>DRS</td>
<td>Data Reduction Software</td>
</tr>
<tr>
<td>ELT</td>
<td>Extremely Large Telescope</td>
</tr>
<tr>
<td>ESO</td>
<td>European Southern Observatory</td>
</tr>
<tr>
<td>FET</td>
<td>Field-Effect Transistor</td>
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<tr>
<td>FIR</td>
<td>Far Infrared</td>
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<tr>
<td>FLUOR</td>
<td>Fiber Linked Unit for Optical Recombination</td>
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<tr>
<td>FoV</td>
<td>Field-of-View</td>
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<tr>
<td>FPA</td>
<td>Focal Plane Array</td>
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<tr>
<td>FPN</td>
<td>Fixed Pattern Noise</td>
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<tr>
<td>Acronym</td>
<td>Definition</td>
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<tr>
<td>FPS</td>
<td>Frames Per Second</td>
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<tr>
<td>FTS</td>
<td>Fourier Transform Spectrograph</td>
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<tr>
<td>FWHM</td>
<td>Full Width at Half Maximum</td>
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<tr>
<td>GI2T</td>
<td>Grand Interéromètre à 2 télescopes</td>
</tr>
<tr>
<td>GMT</td>
<td>Giant Magellen Telescope</td>
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<td>GPI</td>
<td>Gemini Planet Imager</td>
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<td>GSU</td>
<td>Georgia State University</td>
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<tr>
<td>GUI</td>
<td>Graphical User Interface</td>
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<tr>
<td>HgCdTe</td>
<td>Mercury Cadmium Telluride</td>
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<tr>
<td>HOSTS</td>
<td>Hunt for Observable Signatures of Terrestrial Systems</td>
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<td>HST</td>
<td>Hubble Space Telescope</td>
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<td>HZ</td>
<td>Habitable Zone</td>
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<td>InGaAs</td>
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<td>IRAS</td>
<td>The Infrared Astronomical Satellite</td>
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<td>The InfraRed camera</td>
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<td>InfraRed Michelson Array</td>
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<td>IRTF</td>
<td>InfraRed Telescope Facility</td>
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<td>Infrared</td>
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<tr>
<td>ISI</td>
<td>Infrared Spatial Interferometer</td>
</tr>
<tr>
<td>ISO</td>
<td>The Infrared Space Observatory</td>
</tr>
<tr>
<td>JouFLU</td>
<td>Jouvence de FLUOR</td>
</tr>
<tr>
<td>JPL</td>
<td>Jet Propulsion Laboratory</td>
</tr>
<tr>
<td>JWST</td>
<td>James Webb Space Telescope</td>
</tr>
<tr>
<td>KIN</td>
<td>Keck Interferometry Nuller</td>
</tr>
<tr>
<td>KI</td>
<td>Keck Interferometer</td>
</tr>
<tr>
<td>LBTI</td>
<td>Large Binocular Telescope Interferometer</td>
</tr>
<tr>
<td>LDC</td>
<td>Longitudinal Dispersion Corrector</td>
</tr>
<tr>
<td>LD</td>
<td>Limb-Darkened</td>
</tr>
<tr>
<td>LED</td>
<td>Light Emitting Diode</td>
</tr>
<tr>
<td>LEECH</td>
<td>LBTI Exoplanet Exozodi Common Hunt</td>
</tr>
<tr>
<td>LESIA</td>
<td>Laboratoire d’études spatiales et d’instrumentation en astrophysique</td>
</tr>
<tr>
<td>LHB</td>
<td>Late Heavy Bombardment</td>
</tr>
<tr>
<td>Abbreviation</td>
<td>Full Form</td>
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<tr>
<td>--------------</td>
<td>-----------</td>
</tr>
<tr>
<td>LoS</td>
<td>line-of-sight</td>
</tr>
<tr>
<td>LVF</td>
<td>Le Verre Fluoré</td>
</tr>
<tr>
<td>LXD</td>
<td>Long Wavelength Cross-dispersed Mode</td>
</tr>
<tr>
<td>mas</td>
<td>milliarcsecond</td>
</tr>
<tr>
<td>MATISSE</td>
<td>Multi AperTure mid-Infrared SpectroScopic Experiment</td>
</tr>
<tr>
<td>MIRA-I.2</td>
<td>Mitaka optical and InfraRed Array project</td>
</tr>
<tr>
<td>MIRC</td>
<td>Michigan InfraRed Combiner</td>
</tr>
<tr>
<td>MIR</td>
<td>Mid Infrared</td>
</tr>
<tr>
<td>MMR</td>
<td>Mean Motion Resonances</td>
</tr>
<tr>
<td>MROI</td>
<td>Magdalena Ridge Observatory Interferometer</td>
</tr>
<tr>
<td>NAOJ</td>
<td>National Astronomical Observatory of Japan</td>
</tr>
<tr>
<td>NA</td>
<td>Numerical Aperture</td>
</tr>
<tr>
<td>NICMOS</td>
<td>the NICMOS-based science camera for JouFLU</td>
</tr>
<tr>
<td>NIRO</td>
<td>Near InfraRed Observer</td>
</tr>
<tr>
<td>NIR</td>
<td>Near Infrared</td>
</tr>
<tr>
<td>Nloops</td>
<td>the number of times the chip is read</td>
</tr>
<tr>
<td>NOAO</td>
<td>National Optical Astronomy Observatory</td>
</tr>
<tr>
<td>Abbreviation</td>
<td>Full Form</td>
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<tr>
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<tr>
<td>NPOI</td>
<td>Navy Precision Optical Interferometer</td>
</tr>
<tr>
<td>Nreads</td>
<td>the number of times a pixel is read</td>
</tr>
<tr>
<td>NRL</td>
<td>Naval Research Laboratory</td>
</tr>
<tr>
<td>NSF</td>
<td>National Science Foundation</td>
</tr>
<tr>
<td>OAP</td>
<td>Off-Axis Parabolic mirror</td>
</tr>
<tr>
<td>‘OHANA</td>
<td>Optical Hawaiian Array for Nano-radian Astronomy</td>
</tr>
<tr>
<td>OPD Scan</td>
<td>OPD Scanning stage</td>
</tr>
<tr>
<td>OPD Stat</td>
<td>OPD Static stage</td>
</tr>
<tr>
<td>OPD</td>
<td>Optical Path Difference</td>
</tr>
<tr>
<td>OPLE</td>
<td>Optical Path Length Equalization</td>
</tr>
<tr>
<td>OUTPUT</td>
<td>OUTPUT stage</td>
</tr>
<tr>
<td>PAH</td>
<td>Polycyclic Aromatic Hydrocarbon</td>
</tr>
<tr>
<td>Pautron</td>
<td>the electronics that control the readout of the CALI camera</td>
</tr>
<tr>
<td>PA</td>
<td>Position Angle</td>
</tr>
<tr>
<td>pc</td>
<td>parsec</td>
</tr>
<tr>
<td>PFN</td>
<td>Palomar Fiber Nuller</td>
</tr>
<tr>
<td>PIONIER</td>
<td>Precision Integrated-Optics Near-infrared Imaging ExpeRiment</td>
</tr>
<tr>
<td>Abbreviation</td>
<td>Description</td>
</tr>
<tr>
<td>--------------</td>
<td>-------------</td>
</tr>
<tr>
<td>PI</td>
<td>Principal Investigator</td>
</tr>
<tr>
<td>PoPs</td>
<td>Pipes of Pan</td>
</tr>
<tr>
<td>PRNU</td>
<td>Photo Response Non-Uniformity</td>
</tr>
<tr>
<td>PR</td>
<td>Poynting-Robertson</td>
</tr>
<tr>
<td>PSF</td>
<td>Point Spread Function</td>
</tr>
<tr>
<td>PTC</td>
<td>Photon Transfer Curve</td>
</tr>
<tr>
<td>PTI</td>
<td>Palomar Testbed Interferometer</td>
</tr>
<tr>
<td>QE</td>
<td>Quantum efficiency</td>
</tr>
<tr>
<td>ROI</td>
<td>Region(s) of Interest</td>
</tr>
<tr>
<td>RON</td>
<td>Readout noise</td>
</tr>
<tr>
<td>RV</td>
<td>Radial Velocity</td>
</tr>
<tr>
<td>SD</td>
<td>Standard Deviation</td>
</tr>
<tr>
<td>SED</td>
<td>Spectral Energy Distribution</td>
</tr>
<tr>
<td>SIMBAD</td>
<td>Set of Identifications, Measurements and Bibliography for Astronomical Data</td>
</tr>
<tr>
<td>SIM</td>
<td>Space Interferometry Mission</td>
</tr>
<tr>
<td>SNR</td>
<td>Signal-to-Noise Ratio</td>
</tr>
<tr>
<td>SpeX</td>
<td>0.7-5.3 micron medium-resolution spectrograph and imager</td>
</tr>
<tr>
<td>Abbreviation</td>
<td>Description</td>
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<tr>
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</tr>
<tr>
<td>SPHERE</td>
<td>Spectro-Polarimetric High-contrast Exoplanet REsearch</td>
</tr>
<tr>
<td>SST</td>
<td>Spitzer Space Telescope</td>
</tr>
<tr>
<td>sub-mm</td>
<td>sub-millimeter</td>
</tr>
<tr>
<td>SUSI</td>
<td>Sydney University Stellar Interferometer</td>
</tr>
<tr>
<td>SXD</td>
<td>Short Wavelength Cross-dispersed Mode</td>
</tr>
<tr>
<td>TMT</td>
<td>Thirty Meter Telescope</td>
</tr>
<tr>
<td>UD</td>
<td>Uniform Disk</td>
</tr>
<tr>
<td>USNO</td>
<td>United States Naval Observatory</td>
</tr>
<tr>
<td>UT</td>
<td>Unit Telescope</td>
</tr>
<tr>
<td>VAMPIRES</td>
<td>Visible Aperture-Masking Polarimetric Interferometer for Resolving Exoplanetary Signatures</td>
</tr>
<tr>
<td>Viscam</td>
<td>Visible band alignment camera</td>
</tr>
<tr>
<td>VLTI</td>
<td>Very Large Telescope Interferometer</td>
</tr>
<tr>
<td>VLT</td>
<td>Very Large Telescope</td>
</tr>
<tr>
<td>WDS</td>
<td>Washington Double Star</td>
</tr>
<tr>
<td>WISE</td>
<td>Wide-field Infrared Survey Explorer</td>
</tr>
<tr>
<td>WL</td>
<td>CHARA laboratory white-light source</td>
</tr>
</tbody>
</table>
XPS  Newport XPS Motion Controller
ZIMPOL  Zurich Imaging POLarimeter
ZnSe  Zinc Selenium
“An expert is a person who has made all the mistakes that can be made in a very narrow field.”

— Niels Bohr
Part I

Introduction
This chapter serves to present the reader with a primer on the main aspects of a telescope and in particular angular resolution. A brief history is given on the relevant discoveries in optics and the study of the nature of light. This leads to the modern era of interferometry and its development as an astronomical tool. The basic principles and applications of interferometry to modern astronomy are discussed.

1.1 Use of the Telescope

It goes without saying that telescopes are crucial to modern astronomy. Almost every paradigm changing event can be traced to the advent of a new or refined observational technology. It is important to remember there are three fundamental aspects to the function of the telescope:

**Sensitivity** — Telescopes collect and concentrate light making dim sources brighter. The total amount of collecting area of a telescope’s objective lens or primary mirror
determines its sensitivity to faint sources. This is further influenced by the number of optical elements and their transmission or reflection characteristics. In the case of digital light collection techniques, one must also consider the quantum efficiency of the detector. This sensitivity or “light gathering power” is of particular importance to spectroscopy and photometry.

**Magnification** — Telescopes can alter the apparent size of objects, the earliest uses of the telescope were to bring the far near, making distant objects appear larger. The degree of magnification is determined by the ratio of curvature between telescope optics, typically the objective and the eyepiece. However, the amount is strongly limited by the angular resolution. A high magnification image, if it lacks sufficient angular resolution, is all but worthless.

**Angular Resolution** — This is the limit to the size of details it is possible to distinguish for a given source. It may be determined, for example, by the minimum separation needed to distinguish between two point sources. The angular size of an object is dependent upon its physical size and distance. The angular resolution of a telescope is a function of the diameter of the telescope’s opening aperture. A high angular resolution telescope is capable of resolving small details relative to the object’s distance. High angular resolution is a requirement for imaging stellar environments.

The ability to image or even just directly measure the size of an astronomical body gives us access to numerous physical properties including its distance. If its distance is known we may derive its physical size, temperature, and luminosity. Knowledge of a star’s magnitude or
luminosity and its temperature or spectral classification allows it to be placed on the Hertzsprung-Russell (HR) diagram (Hertzsprung 1909; Russell 1914), which is useful for understanding the age of a star and how it will evolve through time. Imaging, in its simplest form, is the measurement of the projected size or diameter of an object along a particular line in the plane of the sky, perpendicular to the line-of-sight (LoS). Recording multiple measurements of diameter may reveal asymmetries like those found in rapidly rotating stars. Only high angular resolution enables these kinds of direct observations. Even the closest stars, apart from the Sun, appear minuscule due to the vast distances involved and even the largest stars have small angular diameters. To image these stars would require telescopes of such size that they are technically and economically unfeasible. Interferometry is an inevitable solution to this problem.

1.2 Interferometry

Interferometry is, in essence, a technique of combining light from separate apertures while maintaining its phase. The interference pattern of light resulting from its wave properties is analogous to water or sound waves. If two stones are dropped in a still pool of water each will generate ripples radiating outward. When the waves originating from each stone cross they will interfere. Where the crests or troughs meet the amplitude of the wave will increase, and where crest meets trough the waves will cancel. This is termed constructive and destructive interference. This pattern of interference is also referred to as fringes. If encoded spatially, the interference pattern is bright and dark bands from constructive and destructive
interference. Alternatively, fringes may be encoded temporally and appear as increases and decreases in the intensity of light.

For astronomy, long-baseline interferometry enables the use of multiple telescopes in order to achieve the resolving power of a much larger telescope. By separating telescopes but maintaining the phase of the light incident upon them, one can get the benefits in angular resolution from a much larger single mirror. In this way it is as if one could carefully sample points of a single giant mirror and combine the light from each sampled point; thereby gaining the angular resolution provided by the separation of the individual points.

If just two points are sampled, the angular size of the observed object can be determined; and with multiple points, sampling various spatial scales, an image of the object may be produced. In this way, the large single mirror of a telescope could be thought of an integrated interferometer, made of an infinite number of discrete points with an infinite number of baselines between them. The greater the separation between the points, the higher is its resolution. By building a large interferometer one can gain the high angular resolution to image stars by sacrificing the light gathering power of a difficult-to-construct, extremely large filled aperture.

The Raleigh criterion for the minimum separation of sources resolvable by a telescope is defined by $\sin \theta_{\text{min}} = 1.22 \frac{\lambda}{d}$, where $\theta_{\text{min}}$ is the minimum angle at which detail can be resolved, $\lambda$ is the wavelength of light which is being observed, and $d$ is the diameter of the telescope. This relation is commonly referred to as angular resolution or the diffraction limit and can rarely be approached due to the effects of atmospheric “seeing” or distortion caused by
turbulence in the atmosphere. The formula is derived from the condition that the first
diffraction minimum of one source coincides with the maximum of another. Under this
circumstance, the two objects are at the absolute minimum separation for them to be
resolved. Note that in this and subsequent formulas the small angle approximation \( \sin \theta \approx \theta \) holds valid.

The minimum angular resolution element for an interferometer is given by \( \theta_{\text{min}} = \frac{\lambda}{2b} \) which is
defined as the minimum separation necessary to distinguish between two sources, \( \theta_{\text{min}} \). In
this equation, \( b \) represents the baseline or separation between two apertures. In addition to
the relative ease of creating very long baselines compared to filled aperture telescopes, there
is an inherent factor of 2.44 advantage in angular resolution by using an interferometer.

1.3 History of Interferometry

The relatively short history of long baseline optical interferometry is borne upon the much
longer history of discoveries in optics leading to acceptance of a wave theory of light. The
first astronomical interferometers would not come until late in the 19th century and the
development of large-scale high precision instrumentation.

1.3.1 Newton’s Rings

The first description of the phenomenon of interference comes from Robert Boyle and Robert
Hooke (Born & Wolf 1999), working independently. Boyle described the colors apparent in
thin films and noted that the color was dependent upon the thickness of the film. In his book
*Micrographia* Hooke described the colors present in thin sheets of Muscovy-glass (mica) and was able to reproduce the effect using lenses (Hooke 1665). Newton analyzed these effects by pressing a convex lens above a flat glass plate (Newton, Isaac 1704). This produces a pattern of bright and dark concentric rings. These rings occur in both transmission and reflection and, when generated with white light, display a pattern of colors. Newton was able to derive a formula that predicts the radius of the ring for a given wavelength:

\[ r_n = \sqrt{R \lambda \left(N - \frac{1}{2}\right)}, \]  

where \( N \) is the ring number, \( \lambda \) is the wavelength, and \( R \) is the radius of curvature of the lens.

### 1.3.2 The Wave Nature of Light

The contributions of Boyle, Hooke, and Newton to the understanding of the colors of thin films as well as those of Grimaldi (Grimaldi 1665) on diffraction lend support to a theory that presents light as some sort of wave. This concept of a wave nature of light was first published by Christian Huygens in 1690 (Huygens 1690). Despite this evidence, the predominant view held by scientists of the time was one put forth by Newton of a corpuscular nature of light. As a result of conflicting personal accounts and rivalries, it is essentially lost to history and rivalry just how much each individual contributed to the explanation of the colors of thin films; but, it is this work that first began to describe, if not explain, the wave nature of light and the principle of interference.

The next large step toward a cohesive theory of the wave nature of light came from Thomas
Young in 1801. Young devised the now famous double-slit experiment (Figure 1.1) in which a narrow-band light source far behind two narrow but closely-placed slits projects light through the two slits and onto a screen. A portion of the light that passes through one slit overlaps with part of the light from the other slit. If light is composed of particles one would expect a distribution of light cast on the screen behind each slit. If corpuscular theory is correct and light is a particle, then where the light from the two slits overlapped, Young should have seen a region of increased brightness behind each slit. In 1803, Young performed his experiment and observed instead a striped pattern of constructive and destructive interference. This effect can only be satisfactorily explained by a wave theory of light.

Not only was corpuscular theory overturned but Young was able to calculate the wavelength of light from the spacing between the light and dark interference bands. In 1807, Young published this work in *A Course of Lectures on Natural Philosophy and the Mechanical Arts* (Young 1807). From this work, \( \tan \theta = \frac{y}{D} \) relates the position of the fringes with angular spacing, \( \theta \approx \frac{\lambda}{d} \), where \( d \) is the separation of the two slits, \( D \) is the distance between the slits and the screen, \( y \) is the distance from the center of the screen to the interference band, and \( \theta \) is the angle from the center of the slits to the interference band maximum. For different orders it can be shown that: \( d \sin \theta = m \lambda \). The equation,

\[
I(x) = I_0 \cos^2 \left[ \frac{\pi xd}{\lambda D} \right],
\]

(1.2)
gives the intensity of light, where \( I_0 \) is the maximum intensity on the screen. Figure 1.2 shows monochromatic fringes of three different wavelengths and the polychromatic fringe resulting
from their superposition.

Figure 1.1: A figure from Young’s 1807 “A course of lectures on natural philosophy and the mechanical arts” showing his double slit experiment.

Work on interference and polarization by Arago and Fresnel (Arago & Fresnel 1819) in 1817 advanced the wave theory of light, demonstrating that the wave must be transverse as opposed to longitudinal like sound. The wave theory of light would remain dominant until the quantum revolution led by Max Planck in 1900 and by Einstein’s explanation of the photoelectric effect in 1905, for which he won the Nobel prize, refined the theory of light into the more enigmatic wave-particle duality we hold today.
Figure 1.2: Example monochromatic fringes and their superposition resulting in a polychromatic fringe packet.

1.3.3 Fourier

In developing his work “The Analytical Theory of Heat” in 1822, Joseph Fourier proposed that any function could be represented as the weighted sum of a series of sines (Fourier 1822). Fourier analysis is now an integral part of the modern world, and the Fourier transform freely allows the processing of information between the frequency or spatial domains. The Fourier transform and its inverse are used extensively in the analysis of interferometric data:

\[
X(f) = \int_{-\infty}^{\infty} x(t)e^{-i2\pi ft} \, dt \quad \text{(1.3)}
\]

\[
x(t) = \int_{-\infty}^{\infty} X(f)e^{i2\pi ft} \, df. \quad \text{(1.4)}
\]
1.4 The Beginnings of Astronomical Interferometry

After Young’s work, a solid basis for interference and its relation to the wave properties of light has been established and mathematical tools developed that can interpret complex interference patterns. But, at this point in the narrative, it is not yet known how to relate the fringe pattern with the physical properties of its source. Furthermore, much of the complexity of astronomical interferometry comes not from the theory but instead from the reality that the construction and operation of a large interferometer requires a level of technological skill not available until the post-industrial era.

1.4.1 Fizeau, Pease, & Michelson

Joseph Fizeau first proposed stellar interferometry in 1868. In his report, he suggested that there is a relation between the size of a light source and the smearing of the interference fringes it produces and that this could serve to put an upper limit on a star’s dimension (Labeyrie et al. 2006). This was later demonstrated in 1873 by Édouard Stéphan who masked the aperture of a single mirror to leave only a pair of widely separated exposed portions of the mirror. Unfortunately, this and subsequent work were only able to place an upper limit on the diameters of the observable stars at 0.16”. This is a factor of 100 times greater than the diameters of many stars observable by interferometers of today.

In his 1890 and 1891 papers, Albert Michelson described methods to observe astronomical sources, such as the Sun and double stars, with aperture-masking interferometry (Michelson
He also defined a primary observable of interferometric fringes, their contrast or visibility, now universally adopted. A simple description of interferometric visibility is given by

\[ V = \frac{\text{fringe amplitude}}{\text{average intensity}} = \frac{I_{\text{max}} - I_{\text{min}}}{I_{\text{max}} + I_{\text{min}}}. \]  

(1.5)

The diffraction-limit of a single aperture telescope is often reduced by the effects of turbulence in the atmosphere, rendering a >1-m aperture telescope no more effective than a 10-cm telescope. Michelson noted, however, that visibility is mostly unaffected by the turbulent atmosphere. In these pioneering works Michelson lays out the dependence of fringe visibility upon the separation of the apertures and even incorporated the effects of limb-darkening upon the measurements of stellar diameter.

Stellar angular diameters are so small that it was not until Michelson and Pease mounted the 20-foot beam interferometer on the Mount Wilson 100-inch Hooker telescope in 1920 that the first angular diameter for a star could be measured (Michelson & Pease 1921). With this instrument (Figure 1.3) they found a uniform disk diameter for \( \alpha \) Orionis of 47 milliarcsecond (mas) with an uncertainty of 10\%.

Work proceeded in constructing an even larger device to specialize in stellar interferometry. The 50-ft stellar interferometer was completed in 1929 (Pease 1930). The new interferometer gave only moderate return for Francis Pease alone; the large structure suffered from engineering and operation difficulties including flexure and vibration of its long truss. It became apparent that the technology was not yet available to reach longer baselines.
Figure 1.3: The 20-foot beam interferometer atop the 100-inch Hooker telescope on Mt. Wilson. From Michelson’s 1921 “Measurements of the Diameter of α Orionis with the Interferometer”.

1.4.2 Principles: van Cittert-Zernike Theorem

A monochromatic electric field can be represented by

\[ E(\mathbf{r}, t) = \Re \left\{ A e^{-i(k \mathbf{r} - kct)} \right\}, \quad (1.6) \]

where \( A \) is the amplitude of the wave and \( k = \frac{2\pi}{\lambda} \). By taking the real part of the electric field in one dimension, this may be written as:

\[ E(t) = A \cos (kx - kct + \phi), \quad (1.7) \]
here $x$ is the optical path length and $\phi$ is the phase of the wavefront. A detector measures the time averaged modulus for $N$ beams,

$$S(k) = \frac{1}{T} \int_0^T \left( \sum_{i=1}^{N} E(t) \right) dt$$  \hspace{1cm} (1.8)

Inserting Equation 1.7 into Equation 1.8 and assuming two-beam combination this becomes

$$S(k) = \frac{1}{T} \int_0^T \left[ A_1 \cos (kx_1 - kct + \phi_1) + A_2 \cos (kx_2 - kct + \phi_2) \right]^2 dt,$$  \hspace{1cm} (1.9)

and expanding this yields

$$S(k) = \frac{1}{T} \int_0^T \left[ A_1^2 \cos^2 (kx_1 - kct + \phi_1) + 2A_1A_2 \cos (kx_1 - kct + \phi_1) \cos (kx_2 - kct + \phi_2) + A_2^2 \cos^2 (kx_2 - kct + \phi_2) \right] dt.$$  \hspace{1cm} (1.10)

A basic trigonometric relation derived from the unit circle,

$$\cos(\beta - \alpha) = \cos \alpha \cos \beta + \sin \alpha \sin \beta,$$

allows this the cross term to be rewritten

$$\cos (kx_1 - kct + \phi_1) \cos (kx_2 - kct + \phi_2) =$$

$$\left[ \cos (kx_1 + \phi_1) \cos (kct) + \sin (kx_1 + \phi_1) \sin (kct) \right] \left[ \cos (kx_2 + \phi_2) \cos (kct) + \sin (kx_2 + \phi_2) \sin (kct) \right].$$  \hspace{1cm} (1.11)

Plugging this into Equation 1.10 and expanding gives

$$S(k) = \frac{1}{T} \int_0^T \left[ A_1^2 \cos^2 (kx_1 - kct + \phi_1) + 2A_1A_2 \left( \cos (kx_1 + \phi_1) \cos (kct) + \sin (kx_1 + \phi_1) \sin (kct) \right) + \right.$$

$$\left. (\cos (kx_2 + \phi_2) \cos (kct) + \sin (kx_2 + \phi_2) \sin (kct)) + A_2^2 \cos^2 (kx_2 - kct + \phi_2) \right] dt.$$  \hspace{1cm} (1.12)
The integrals: \( \frac{1}{2\pi} \int_0^{2\pi} \cos^2 x \, dx = \frac{1}{2\pi} \int_0^{2\pi} \sin^2 x \, dx = \frac{1}{2} \) and \( \frac{1}{2\pi} \int_0^{2\pi} \cos x \sin x \, dx = 0 \) allow us to greatly simplify the equation.

\[
S(k) = \frac{1}{2} A_1^2 + \frac{1}{2} A_2^2 + A_1 A_2 \left[ \cos (kx_1 + \phi_1) \cos (kx_2 + \phi_2) + \sin (kx_1 + \phi_1) \sin (kx_2 + \phi_2) \right]
\]  

(1.13)

Substituting \( I = A^2 \) and condensing the trigonometry gives

\[
S(k) = \frac{1}{2} I_1 + \frac{1}{2} I_2 + \sqrt{I_1 I_2} \cos (k(x_1 - x_2) + (\phi_1 - \phi_2)).
\]  

(1.14)

The visibility amplitudes are normalized so that the results are between 0 and 1 and the signal is divided by the mean intensity.

\[
S(k) = 1 + 2\sqrt{\frac{I_1 I_2}{I_1 + I_2}} \cos (k(x_1 - x_2) + (\phi_1 - \phi_2)).
\]  

(1.15)

Equation 1.15 is known as the monochromatic fringe equation and introduces the transfer function,

\[
T = \frac{2\sqrt{I_1 I_2}}{I_1 + I_2}.
\]  

(1.16)

The full N-beam derivation of this is found in ten Brummelaar (2014).

In reality, light from a physical source is never purely monochromatic. Light from a real source has some bandwidth, \( \Delta \lambda \), and its fringes are finite and only exist near the zero Optical Path Difference (OPD) position.

To describe the wave field produced by a polychromatic source it is helpful to introduce the concept of correlation and coherence. In interferometry, we are essentially exploiting the nature of how light interferes with itself and then measuring the cross-correlation. The
cross-correlation of two signals $E_1$ and $E_2$ is defined as

$$\Gamma(\tau) = \int_{-\infty}^{+\infty} E_1(t) \times E_2^*(t + \tau) \, dt,$$

where $E_2^*$ is the complex conjugate of the electric field.

Born & Wolf (1999) present a full derivation of discussion of correlation functions of light beams. For our purposes here, we will introduce the mutual coherence function for our detected electric field, which is a specific form of the cross-correlation function,

$$\Gamma_{12}(\tau) = \left< S_1(t + \tau) S_2^*(t) \right>.$$  \hspace{1cm} (1.18)

The mutual coherence function can be normalized to the autocorrelation functions, yielding the complex degree of coherence,

$$\gamma_{12}(\tau) = \frac{\Gamma_{12}(\tau)}{\sqrt{I_1 I_2}}.$$  \hspace{1cm} (1.19)

From this we can obtain the general interference formulation,

$$I = I_1 + I_2 + 2\sqrt{I_1 I_2} \gamma_{12} e^{k(s_2 - s_1)},$$  \hspace{1cm} (1.20)

where $s$ is the path length. Taking only the real part of the complex degree of coherence, this can be written as

$$I = I_1 + I_2 + 2\sqrt{I_1 I_2} |\gamma_{12}| \cos(2\pi\sigma \Delta x + \Phi_{12}),$$  \hspace{1cm} (1.21)

where $\sigma$ is the wavenumber and $\sigma = 1/\lambda = k/2\pi$. If $0 < \gamma_{12} < 1$ then the waves are partially coherent. It is this degree of coherence, along with the phase term, which we measure with
an interferometer. Bringing this together with Equation 1.15 and Equation 1.16, gives us

\[ S(k) = 1 + T_{12} |\gamma_{12}| \cos [2\pi \sigma \Delta x + \Delta \phi + \Phi_{12}]. \] (1.22)

Because light is a wave and is capable of exhibiting interference, it must then possess the property of coherence. Waves are perfectly coherent if they have the same frequency and phase. The degree to which two waves, or portions of the same wave, match can be measured by its degree of correlation. For a partially coherent source, the cross-correlation is between 0 and 1. A wave may maintain a certain degree of coherence over a distance of its propagation. This is the coherence length, \( L \),

\[ L = \frac{\lambda^2_0}{\Delta \lambda} = R \times \lambda_0, \] (1.23)

where \( R = \frac{\lambda}{\Delta \lambda} \) is the spectral resolving power, \( \lambda_0 \) is the central wavelength, and \( \Delta \lambda \) is the bandwidth. This is related to the coherence time by

\[ t_c = \frac{L}{c}, \] (1.24)

where \( c \) is the speed of light. A polychromatic wave, thanks to Fourier, can be expanded as a sum of monochromatic waves.

\[ E(r, t) = \int_0^\infty A(\lambda, t) e^{-i2\pi(\sigma - \frac{\sigma_0}{2})} d\lambda. \] (1.25)

If we restrict the possible range of wavelengths to \( \lambda_0 \pm \Delta \lambda/2 \), or in terms of wave number \( \sigma_0 \pm \Delta \sigma/2 \), where wavenumber is \( \sigma = 2\pi/\lambda \), we develop the quasi-monochromatic case with a finite bandwidth. Let’s assume the bandpass is a top-hat function centered on the central
wavelength. Integrating Equation 1.22 across the bandwidth gives

\[ S(k) = \frac{1}{\Delta \sigma} \int_{\sigma_0-\Delta \sigma}^{\sigma_0+\Delta \sigma} \left( 1 + T_{12} |\gamma_{12}| \cos [2\pi \sigma \Delta x + \Delta \phi + \Phi_{12}] \right) d\sigma. \] (1.26)

Evaluating the integral leads to

\[ S(k) = \frac{1}{\Delta \sigma} \left. \sigma^{\sigma_0+\Delta \sigma} \left( T_{12} |\gamma_{12}| \frac{\sin 2\pi \sigma \Delta x + \Delta \phi + \Phi_{12}}{2\pi \Delta x} \right) \right|_{\sigma_0-\Delta \sigma}^{\sigma_0+\Delta \sigma} \] (1.27)

and using the trigonometric identity \( \sin (a + b) - \sin (a - b) = \)
\[ \sin (a) \cos (b) + \sin (b) \cos (a) - \sin (a) \cos (b) + \sin (b) \cos (a) = 2 \sin (b) \cos (a) \] with \( a = 2\pi \sigma_0 \Delta x + \Delta \phi + \Phi_{12} \) and \( b = \pi \Delta \sigma \Delta x \) to evaluate the limit conditions results in

\[ S(k) = 1 + T_{12} |\gamma_{12}| \frac{\sin (\pi \Delta \sigma \Delta x) \cos (2\pi \sigma_0 \Delta x + \Delta \phi + \Phi_{12})}{\pi \Delta \sigma \Delta x} \] (1.28)

which can be further simplified to

\[ S(k) = 1 + T_{12} |\gamma_{12}| \sin (\pi \Delta \sigma \Delta x) \cos (2\pi \sigma_0 \Delta x + \Delta \phi + \Phi_{12}) \] (1.29)

This is the fringe equation for the quasi-monochromatic case. The fringe envelope is the Fourier transform of the optical bandpass. In this case, the bandpass was a top-hat function and the resulting envelope is a sinc function. As one restricts the bandwidth, the monochromatic case is approached and the fringe envelope, or packet, widens; on the contrary, with a broad bandwidth, a narrower fringe packet is produced (See Figure 1.2).

Today's work in interferometry relies upon the concepts of the coherence function developed by van Cittert in 1934 (van Cittert 1934) and Zernike in 1938 (Zernike 1938). Recalling Young's double-slit experiment or any setup where light from a distant source passes through
a mask to project upon a screen, if multiple sources are present then a superposition of fringes will be observed on the screen. If a source is large enough, a blurring of the fringes will be seen. The larger the source, the greater the blurring. While the fringes are distinct, the wavefront of the light can be considered coherent. The distance over which light is spatially coherent is the coherence length (Equation 1.23) and depends upon the wavelength and the bandwidth. If the source size (or slit separation) is increased beyond this coherence distance fringes will no longer be seen.

A thought experiment to illustrate spatial coherence is to consider the analogy of a pond where a large object drops into the still water. Near to the object, the water’s surface is chaotic but as the waves travel to the far shore they are orderly ripples or it could be stated; they are spatially coherent.

If the field emitted by one source is sampled at two points such that the light paths are within the coherence length, the self-interference will not necessarily be zero. Distinct astronomical sources are spatially incoherent; however, point sources are self-coherent. In between these two cases there exists a range where partial coherence is displayed. The relationship between partial coherence and the object is the basis of the van Cittert-Zernike theorem.

To put this another way: A spherical electric field wave from a light source radiates outward at the speed, c. A photon does not exist until detected, instead there is the probability wave and a certain chance that the photon will interact and collapse the wave. It is these waves that can interfere if they are coherent. It is also possible for these waves to interfere with themselves. In stellar interferometry, photons originate from a distance source. Due to the
vast distance involved, the source is effectively a point source. The photons may originate from different regions of the surface of a star and are not coherent with each other. But the probability waves of each photon pass through multiple apertures and the photons in the form of this wave interfere with themselves. Upon striking a detector, the wave function collapses into a photon and this results in the fringe pattern that we can measure.

The van Cittert-Zernike theorem formally relates the fringe contrast to the angular distribution of the source on the sky. Proof of this theorem is available in Born & Wolf (1999). The van Cittert-Zernike theorem states that for a monochromatic incoherent source, the Fourier transform of the complex spatial coherence function yields the angular intensity distribution of the source. This relation is described in terms of the complex visibility, $\gamma$, the intensity distribution of the source, $S(\alpha')$, and the spatial frequency in the Fourier plane, $u = \frac{\beta}{\lambda}$, where $u$ and $v$ are components of the baseline vector projected on the sky:

$$\gamma(u, v) = \gamma(u, v) = \int \int S(\alpha') e^{2\pi i \alpha' \cdot \vec{u}} d\alpha' = |\gamma| e^{i\phi}. \quad (1.30)$$

More simply as the mutual coherence is related the Fourier transform of the target intensity map by

$$\gamma(u, v) = F(I(\alpha, \beta)). \quad (1.31)$$

Five assumptions are made with van Cittert-Zernike theorem. First, the source is assumed to be distant and the in the far field condition. Second, the source is assumed to be small in angle but extended in two-dimensions. Astronomical sources are, of course, three-dimensional but in the far field are observed as two-dimensional projections on the sky.
As discussed earlier, the quasi-monochromatic case is assumed, where the light is filtered through a finite bandpass. Next, it is assumed that the source is spatially incoherent, which is the case for most astronomical sources. Finally, space is assumed to be a homogeneous medium, so that light from each region of the source is not differentially refracted in comparison to light from other regions of the source. This holds true to all but the smallest degree, until the light enters Earth’s atmosphere, where variations due to turbulence impart distortions on the wavefront. This results in the need for careful calibration in order to accurately interpret interferometric observations.

From this theorem, it can be shown that an incoherent source observed from a great distance may display spatial coherence. This means it is possible to measure fringes from astronomical sources, and with a multiple aperture interferometer each baseline gives one component of the Fourier transform of the source. The visibility is the modulus of the degree of coherence, $|\gamma|$, and the phase is the argument, $\arg(\gamma)$.

This demonstrates the basic relationship between coherence and the visibility of fringes. The visibility modulus for a uniform disk as a source is given by

$$V = \left| \frac{2J_1 \left( \frac{\pi B\theta}{\lambda} \right)}{\left( \frac{\pi B\theta}{\lambda} \right)} \right|$$

where $J$ is the Bessel function of the first kind and first order, and $B$ is the separation between two apertures. This function is the Fourier transform of the uniform disk. Figure 1.4 shows the visibility curves for stars of various angular diameter.

So far we have assumed the simplest case of a two-beam interferometer and focused on the
Figure 1.4: Visibility curves for stars of 1, 2, 4, 8, and 16 mas in K band. A source is said to be unresolved where V is close to 1, resolved if V = 0, and over-resolved after the first null.

Visibility term. However, from van Cittert-Zernike theorem there is also the phase term, $\Phi$, which is a measurable quantity. However, much of the phase information is corrupted by the effects of the turbulent atmosphere. If a patch of turbulent air crosses into the path of a single telescope of an interferometer, this will change the amount of delay and induce a shift in the interference pattern. Large shifts can cause phase wrapping, making it impossible to unambiguously measure phase. Imaging of complex or asymmetric objects relies on complex phase information. Thankfully, there exists a way to recover at least some of this information. The shift in delay over one telescope has a corresponding shift in the opposite direction of the
other telescope. So if one were to form a triangle of three telescopes one could express,

\[
\begin{align*}
\phi_{12} &= \Phi_{12} + \epsilon_1 - \epsilon_2 \\
\phi_{23} &= \Phi_{23} + \epsilon_2 - \epsilon_3 \\
\phi_{31} &= \Phi_{31} + \epsilon_3 - \epsilon_1 \\
CP &= \phi_{12} + \phi_{23} + \phi_{31}.
\end{align*}
\]

Here, \( \epsilon \) is the unknown phase error; and \( \phi \) and \( \Phi \) represent, respectively, the measured and real phase of a baseline pair. The sum of these phases is the closure phase, \( CP \), which is an observable quantity where individual phase delays introduced by the atmospheric turbulence over a specific telescope cancel out. This technique was first introduced for radio interferometry by Roger Jennison in 1958 (Jennison 1958).

1.4.3 Early Long Baseline Optical Interferometry

After the limited success of the 50-ft interferometer on Mt. Wilson the field of optical astronomical interferometry stagnated for many years. It was not until 1956 that Robert Hanbury Brown and Richard Twiss developed the intensity interferometer which they tested on Sirius (Hanbury Brown & Twiss 1956). They measured correlated intensity to determine stellar diameters. Prior to this the only developments in astronomical interferometry had come from the radio wavebands. A new intensity interferometer was constructed at Narrabri Observatory in the 1960's which measured the diameters of 32 stars (Hanbury Brown et al. 1974). This was followed by Antoine Labeyrie constructing a 12-m baseline amplitude interferometer at Nice Observatory (Labeyrie 1975). This was the first practical demonstration
of a long-baseline optical Michelson interferometer. William Tango and Richard Twiss discuss in detail the challenges of designing a long baseline Michelson stellar interferometer in Tango & Twiss (1980).

1.4.4 The Case for Long-Baseline Astronomical Interferometry

![Figure 1.5](image)

Figure 1.5: This figure compares the angular resolution of single aperture telescopes to multiple aperture long baseline optical interferometers.

The advantages of interferometry to perform high angular resolution measurements are numerous. For bright sources it is possible to resolve spectroscopic binaries, measure limb diameters, and even image stellar features such as starspots. It is interesting to note that if one follows the precedent that aperture doubles every 40 years, it would be the year 2150 before we could build a 330-m telescope (van Belle et al. 2004) but the operation of several hundred meter baseline interferometers is happening today. Even kilometer baseline ground-based interferometers are technically possible if funded. Van Belle notes that the cost
of modern telescopes scales with aperture size to the 2.5 power. In his report, he suggests that there exists a tipping point for large single aperture (100 m class) telescopes, above which space-based telescopes are more cost effective. Ground-based interferometry can now achieve and exceed angular resolutions from the overwhelmingly-large class of ground-based telescopes now being envisioned. Figure 1.5 compares the angular resolution of existing and proposed telescopes and interferometers while Figure 1.6 compares the cost of filled-aperture telescopes, interferometers, and space-based telescopes. The future generation large filled-aperture telescopes will likely be oversubscribed for work on
extragalactic and cosmological applications, making interferometers the ideal choice for stellar and astrophysical investigation. Furthermore, there are numerous opportunities for space-based interferometers or hypertelescopes, which are an hemispherical array of mirrors that focus light onto a common point, and ground-based kilometer and sub-kilometer baseline interferometer designs. Ultimately, to reach the ever more demanding observational goals in high resolution astronomy, interferometry is an inevitable outcome for the future of astronomy.
“Space is big. You just won’t believe how vastly, hugely, mind-bogglingly big it is. I mean, you may think it’s a long way down the road to the chemist’s, but that’s just peanuts to space.”

— Douglas Adams
CHARA and current interferometry

Georgia State University’s Center for High Angular Resolution Astronomy (CHARA) currently operates the most powerful long-baseline optical interferometer in the world at Mt. Wilson Observatory, California. This optical and infrared interferometer consists of six one-meter telescopes in a Y-configuration (see Figure 2.1) and has baselines available between 34-m and 331-m. It has a limiting resolution of 200 micro-arcseconds in the visible wave band. This would be equivalent to seeing details the size of a coin from 16,000-km away. The light from each telescope is conveyed via evacuated tubes to a central Beam Synthesis Facility (BSF). In this 100-m long building the individual light beams pass through a complex system of motorized carts and optics to ensure that each beam travels the exact same distance, in order to maintain zero Optical Path Difference (OPD) throughout the night as the Earth rotates (see Figure 2.2). When light from the same source travels along two or more equidistant paths, it is termed zero OPD. Once the path lengths are equalized, the beams are
passed to a beam combining laboratory housing numerous instruments that prepare the incoming beams of light for interferometric combination and obtain scientific measurements of fringe visibility and phase. A complete description of the CHARA Array is available in ten Brummelaar et al. (2005).

Figure 2.1: Layout of the CHARA Array’s six telescopes, light pipes, and Beam Combining Lab within the context of the other facilities on Mt. Wilson (left). The illustration on the right shows the size of a mirror of equivalent resolving power to the CHARA Array. Also visible within the outline of the Beam Combining Lab are the ”Pipes of Pan” (PoPs). Mirrors inserted at these points allow for various fixed intervals of large delay.

2.1 Background

Long baseline optical interferometry in astronomy is a unique and powerful technique with over a century of history. In the recent decades it has emerged as a rapidly advancing field. Facilities such as the CHARA Array (ten Brummelaar et al. 2005), Navy Precision Optical Interferometer (NPOI) (Armstrong et al. 1998), Sydney University Stellar Interferometer (SUSI) (Davis et al. 1999a), and Very Large Telescope Interferometer (VLTI) (Glindemann et al. 2000) utilize this technology for astronomical investigations (refer to
section 2.3 for more detail). Interferometry’s ability to resolve the finest details possible from
the ground enables astronomers to study stars as more than points of light. Starspots, stellar
shapes, debris disks, circumstellar environments, and compact binary systems among many
other key points of interest may be studied in detail.

![Figure 2.2: Layout of basic long baseline interferometer.](image)

$B \sin \theta$ is the optical path delay which must be compensated for in order to achieve fringes.

2.2 Details of the Array

The CHARA Array consists of six one meter altitude-azimuth telescopes arranged in a
Y-configuration along the North-East, South, and North-West directions. Two telescopes lie
along each arm of the “Y”, a distal one designated “1” and proximal one, “2”. The telescopes
are afocal beam compressors with a compression ratio of 0.125. The location of these
telemoscopes provides 15 non-redundant baselines from 34 to 331-m. Starlight is conveyed
from the telescopes to the Beam Synthesis Facility (BSF) via evacuated pipes. A 12.5-cm collimated beam travels through the 20-cm diameter pipes at 0.1% atmospheric pressure. This minimizes the affects of ground-level turbulence generated by uneven terrain and differential heating.

Once inside the BSF, the light encounters the vacuum “turning boxes”, which orient the beam paths into six parallel lines and into the Optical Path Length Equalization (OPLE) facility. There the light encounters the Pipes of Pan (PoPs), a switchable optical delay system so named because they are a series of parallel pipes much like a Pan flute. Within these pipes are fixed stations that house mirrors which can be introduced into the beam. The PoPs are used to introduce large amounts of fixed delay compensation: 0, 36.6, 73.2, 109.7, or 143.1-m of delay can be introduced on any beam or combination of beams. This provides for delay based on which region of the sky the star is in relative to the projected baseline orientation.

After reflecting off of the PoPs mirror, the starlight is fed through a periscope and leaves vacuum. The light travels parallel but above the vacuum lines inside the lab until incident upon the OPLE carts. These carts provide the continuously variable delay necessary as the telescope tracks the star across the sky. The carts travel along a pair of precisely aligned 45-m long steel rails.

The cart itself is magnetically coupled to its drive and carries a parabolic mirror at the back and a secondary mirror mounted on a piezo stack at the front of the cart to form a cat’s-eye retro-reflector. Cat’s eye retro-reflectors have the advantage of a wide viewing angle
compared to corner cubes. The starlight bounces off the parabola to the secondary and back again, leaving the cart parallel to its incident beam but offset horizontally. Also running parallel but outside the science beam is the metrology laser beam. This is a 1.3-\(\mu\)m laser which tracks the cart position to \(\pm 2\)-nm.

Delay compensation is achieved in stages by servoing the position of the cart on the rails and controlling the magnetic voice-coils. One coil connects the isolated drive section of the cart, the other coil is attached to the parabolic mirror on the cart. Fine delay compensation is achieved by varying the position of the secondary with a piezo stack. After leaving the cart the starlight passes through the Beam Reducing Telescopes (BRTs) which compress the 12.5-cm beam to a 1.9-cm beam. This beam is then split into its visible and NIR components (\(\lambda > 1\mu\text{m}\)) by the beam separation stage carrying a dichroic and a flat mirror. This allows starlight from any telescope to be placed in any of the six beams.

At this point we now have six visible and six NIR beams leaving the BSF after having undergone equal reflections to maintain polarization. All of the beams are at this point delay compensated, and any telescope can feed starlight into any beam. Longitudinal Dispersion Correctors (LDCs) correct for differences in air path between the beams. These beams all feed into the Beam Combination Laboratory (BCL) where the visible beams pass through another dichroic beamsplitter which directs a portion of the visible beam to the active “tip/tilt” system that is in a control loop with the telescope secondary mirrors. The remaining portion of the visible beam may continue to one of the visible waveband beam combiners. The infrared beams continue unimpeded to one of the Infrared (IR) beam combiners. Table 2.1
compares the available beam combiners at the CHARA array.

Figure 2.3: The CHARA delay lines.

Figure 2.4: The CHARA OPLE Carts.
Figure 2.5: The CHARA Beam Samplers.
Table 2.1: A list of the available beam combiners and instruments in operation at the CHARA Array. The central wavelengths for the wavebands, V, R, I, J, H, K, are approximately: 0.54, 0.64, 0.8, 1.2, 1.6, and 2.2 \( \mu \text{m} \), respectively. The separation of bands for VEGA is 30nm.

<table>
<thead>
<tr>
<th>Mode</th>
<th>Telescopes</th>
<th>Waveband</th>
<th>Limiting Mag</th>
<th>Typical</th>
<th>Best</th>
<th>Spectral Resolution</th>
<th>Type</th>
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<td>5.0</td>
<td>30000</td>
<td>Open air</td>
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<td>6.5</td>
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<td>6000</td>
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<td>H:K</td>
<td>4.5 (3.0)</td>
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<td>Fiber-based imager</td>
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<td>V:R:I</td>
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<td>Fiber</td>
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Table 2.2: Major features of recent long baseline optical interferometers. Parenthetical numbers denote design capabilities which may not be fully operational yet.

<table>
<thead>
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<th>Interferometer</th>
<th>Waveband</th>
<th>Number of elements</th>
<th>Aperture (m)</th>
<th>Limiting magnitude</th>
<th>Baseline Min(m)</th>
<th>Max(m)</th>
<th>V2 Accuracy</th>
<th>Status</th>
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<td>R</td>
<td>I</td>
<td>J</td>
<td>H</td>
<td>K</td>
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<td>V</td>
<td>R</td>
<td>I</td>
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<td>H</td>
<td>K</td>
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<td>8.2+1.8</td>
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<td>I</td>
<td>J</td>
<td>H</td>
<td>K</td>
<td>1(10)</td>
<td>1.4</td>
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<td>H</td>
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<td>10</td>
<td>10.3</td>
<td>85</td>
</tr>
<tr>
<td>Palomar Testbed Interferometer</td>
<td>J</td>
<td>H</td>
<td>K</td>
<td>3</td>
<td>0.4</td>
<td>7</td>
<td>86</td>
<td>110</td>
</tr>
</tbody>
</table>
2.3 Retired Interferometers

Numerous optical interferometers have contributed to the field of astronomy over the past twenty years. A few of the major large arrays, past and present, are listed in Table 2.2. Many early interferometers were constructed with a short lifespan or as proof-of-concept facilities. These interferometers paved the way for the current generation of long-baseline optical interferometers. Nevertheless these facilities served as valuable testbeds for current beam combination techniques and many contributed groundbreaking scientific results.

2.3.1 COAST

The Cambridge Optical Aperture Synthesis Telescope (COAST) is made up of four siderostats operated by the University of Cambridge. Its largest commissioned baseline was 67-m but it was built to operate at up to 100-m. COAST was the first facility to produce an image from interferometric data. It is no longer operating as an interferometer, but is a testbed for delay lines and other subsystems for other interferometers such as Magdalena Ridge Observatory Interferometer (MROI) and Very Large Telescope Interferometer (VLTI).

2.3.2 GI2T

The Grand Interéromètre à 2 télescopes (GI2T) is located at the Observatoire de la Côte d’Azur in the south of France, near Grasse (Mourard et al. 1994). Preceded by earlier smaller aperture interferometers but following the same idea to construct an optical
interferometer array presented by Labeyrie (Labeyrie 1976), GI2T is a pair of 1.5-m Cassegrain-Coudé telescopes that feed into a central laboratory. This Michelson-type interferometer utilized many novel and unique solutions to the challenges of optical interferometry. One example of this is the use of the now iconic spherical concrete telescope housings. It was the first interferometer to have high spectral resolution and the first to combine apertures larger than the characteristic seeing size, $r_0$.

2.3.3 IRMA

InfraRed Michelson Array (IRMA) was two 20-cm alt-az siderostats feeding fixed telescopes from a single N-S baseline (Dyck et al. 1993). The siderostats could be stationed at from a 2.5-m minimum separation to 19.5-m apart at 0.5-m intervals. This interferometer was the first to operate in the NIR. It successfully demonstrated repeatable visibility measures approaching 1% precision and stellar diameter measurements with a precision of 2%.

2.3.4 IOTA

The Infrared Optical Telescope Array (IOTA) was a Michelson stellar interferometer with three 45-cm siderostats in an L-shaped array. It resulted from a collaboration of the Smithsonian Astrophysical Observatory, Harvard University, the University of Massachusetts, the University of Wyoming, and MIT/Lincoln Laboratory. IOTA saw first fringes in 1993 and operated on Kitt Peak National Observatory until 2006. It was the first facility to use a fiber-based beam combiner and produced numerous papers and theses.
2.3.5 PTI

Caltech and NASA’s Jet Propulsion Laboratory (JPL) worked together to construct the Palomar Testbed Interferometer (PTI), which operated from 1995 to 2008. It served as a testbed for new interferometric technologies also producing over 50 refereed papers. It demonstrated new interferometric techniques, such as dual beam astrometry. Many of these techniques were applied to the Keck Interferometer (KI) and to subsystems of the CHARA Array.

2.3.6 KI

The KI was part of NASA’s Exoplanet Exploration Program. It joins the twin 10-m Keck telescopes at the summit of Mauna Kea on Hawaii’s “big island” into a single 85-m baseline interferometer. The project began in 1996 and since 2003 has been highly successful in studying stellar and galactic phenomena including gas and dust disks, young stars, and planet formation. It was the first very large telescope interferometer and it was the first to utilize Adaptive Optics (AO). The project was canceled in 2010 after the Space Interferometry Mission (SIM) was canceled and the planned outrigger telescopes for KI were not installed.

2.4 Currently Operating Interferometers

The following interferometers represent the current state-of-the-art in the field of interferometry. These interferometers, along with the CHARA Array, combine science with
computer, mechanical, and optical engineering to contribute to the astronomical community and showcase interferometry as a mature scientific discipline.

2.4.1 LBTI

The Large Binocular Telescope Interferometer (LBTI), located on Mt. Graham, is unique among the interferometers on this list due to its two large 8.4-m single-aperture mirrors sharing a common mount (Hinz et al. 2004). This has the advantage of giving the interferometer a very faint limiting magnitude and a 23-m baseline. LBTI achieved first fringes in 2010 and has proceeded to incorporate adaptive optics (Hinz et al. 2012). The project is funded in part by NASA as a pathfinder for future missions intent on imaging exo-Earths. LBTI has begun a crucial role in the study of exozodiacal dust.

2.4.2 NPOI

The Navy Precision Optical Interferometer (NPOI) is located on Anderson Mesa near Flagstaff, AZ and is part of the United States Naval Observatory (USNO) and operated by Lowell Observatory. Primary use of NPOI is through the Naval Research Laboratory (NRL). Astrometry, or the accurate measurement of the positions of celestial objects, is a primary purpose for this facility. The U.S. Navy uses this astrometric data for traditional navigation. The site consists of four fixed astrometric stations and six portable 0.5-m siderostats (Armstrong et al. 1998). The facility has 10 stations with baselines available from 9 to 179-m. Ultimately NPOI aims to reach a 432-m maximum baseline. Recently NPOI has begun
upgrades to include four 1.8-m telescopes (Armstrong et al. 2014).

2.4.3 SUSI

The Sydney University Stellar Interferometer (SUSI) is a two-element interferometer that consists of a fixed N-S linear array of eleven 14-cm effective aperture siderostats (Davis et al. 1999b). Based on an initial prototype that had a baseline of 11.-m. It is designed to operate with baselines ranging from 5 to 640-m; however only baselines up to 160-m have been implemented. Much of the design and software architecture of SUSI influenced the CHARA Array.

2.4.4 VLTI

The European Southern Observatory (ESO) operates the VLTI in Paranal. VLTI coherently combines light from the four Very Large Telescope (VLT) Unit Telescopes (UTs) or the four movable Auxiliary Telescopes (ATs). The UTs are 8.2-m telescopes and the ATs are 1.8-m telescopes. The ATs can be placed at any of up to 30 stations. The interferometer hosts multiple beam combining instruments; VINCI, the VLT Interferometer Commissioning Instrument (Kervella et al. 2000), the first generation AMBER (Petrov et al. 2007), MIDI Leinert et al. (2003), and PIONIER (Le Bouquin et al. 2011), and the second generation GRAVITY (Eisenhauer et al. 2008) and MATISSE (Lagarde et al. 2012).
2.4.5 **ISI**

The Infrared Spatial Interferometer (ISI) is a Mid Infrared (MIR) interferometer operated by the University of California at Berkeley that utilizes heterodyne detection (Hale et al. 2000). It is made up of three movable 1.65-m siderostats located on Mt. Wilson in proximity to CHARA. The heterodyne detection method does not require the transport of the beams to a BSF and is more closely related to radio interferometry.

2.4.6 **MIRA-I.2**

Mitaka optical and InfraRed Array project (MIRA-I.2) is a 30-m optical interferometer located at the National Astronomical Observatory of Japan (NAOJ) (Ohishi et al. 2008). It achieved first fringes in 2002 and began upgrades in 2008.

2.5 **Future Interferometers**

The short-term funding climate will always be uncertain; however, new advances in stellar theory and astrophysical modeling drive the demand for ever higher resolution and imaging in astronomy. The only currently feasible method to reach the necessary resolution is long-baseline optical interferometry. Currently operating interferometers like the CHARA Array, NPOI, VLTI, and many others have pushed optical interferometry from an esoteric fringe science into the mainstream. Optical interferometric data are routinely published and cited at an impressive rate considering the relatively modest facilities and small community.
The next generation of interferometers will hopefully continue to push technological limits and broaden the application of this technique.

2.5.1 MROI

MROI is an ambitious project led by New Mexico Tech. It will ultimately be a ten-element interferometer operating 1.4-m telescopes in wavelength regimes from 0.6 to 2.4-µm. The telescopes and their enclosures will be movable between multiple stations for flexible configuration along baselines from 8.7 to 340-m. MROI’s design incorporates several new techniques and numerous incremental improvements to well-proven technologies developed at other facilities. Groundbreaking on the MROI telescopes began in 2011 and is awaiting funding for completion.

2.5.2 ‘OHANA

The Optical Hawaiian Array for Nano-radian Astronomy (‘OHANA) project plans to link seven existing 3 to 10-m telescopes on Mauna Kea with single-mode optical fibers. Once completed ‘OHANA will have baselines ranging from 80 to 800-m and reach angular resolutions of a few 0.1-mas (Perrin et al. 2003).
“Science is a way of trying not to fool yourself. The first principle is that you must not fool yourself, and you are the easiest person to fool.”

— Richard Feynman
Part II

Instrumentation
The FLUOR is a high-precision visibility instrument for interferometry. Built by the Laboratoire d'études spatiales et d'instrumentation en astrophysique (LESIA) of the Observatoire de Paris, FLUOR produces some of the most precise measurements ever made in stellar interferometry. Originally set up on Kitt Peak, Arizona in 1992, FLUOR was moved to the IOTA on Mt. Hopkins in 1995. FLUOR has been operating at the CHARA Array since 2002.

FLUOR is a two-way infrared interferometric beam combiner operating in the K’ band ($\lambda = 2.20\mu m$, $\Delta \lambda = 0.40$ FWHM). It utilizes the spatial filtering properties of optical fibers to produce visibility measurements with a precision of approximately $\pm 0.3\%$ (Coudé du Foresto et al. 1997). For bright sources, the statistical precision of FLUOR is only limited by the piston mode of atmospheric turbulence which introduces apparent changes in the zero OPD fringe position or fringe jitter. Such high precision allows FLUOR to make measurements of scientifically interesting features such as stellar radii that are accurate to the order of one percent or less.
3.1 Atmospheric Effects

As with all ground-based optical astronomy, interferometric observations suffer due to the atmosphere. Existing tip/tilt correction and future Adaptive Optics (AO) systems at the CHARA Array can help to mitigate this problem, but they cannot remove all effects of atmospheric turbulence on the wavefront.

Differences in temperature, humidity, and pressure, as well as convection cells, winds, and small scale turbulence change the index of refraction of air. These inhomogeneities distort the wavefront and degrade the image of an astronomical source. The effect of random aberrations introduced by passage through the atmosphere is termed “seeing”.

There is a characteristic size of the turbulent cells with large cells spawning smaller eddies. The large and small scales are known as outer and inner scale turbulence. For the atmosphere the outer scale, where energy is added to the system, is on the order of tens of meters, while the inner scale, where the viscosity of the medium dissipates energy, is on the order of millimeters.

In 1941, Andreï Kolmogorov proposed a theory of turbulence wherein inner-scale turbulence is locally homogeneous and isotropic, while outer-scale turbulence is inhomogeneous and anisotropic (Kolmogorov 1941). Energy is transferred from outer to inner scales, where it is dissipated. In addition to the characteristic scale of the turbulence, he proposed two parameters for determining the strength of inner-scale turbulence: the rate of energy generation per unit mass, $\epsilon$, and the kinematic viscosity, $\nu$. $\epsilon$ and has the units $m^2 s^{-3}$, and $\nu$
has the units $m^2s^{-1}$. From dimensional analysis, the Kolmogorov length scale, $\eta$, can be derived,

$$\eta = \nu^{\frac{3}{4}} \epsilon^{-\frac{1}{4}}$$  \hspace{1cm} (3.1)$$

and the corresponding velocity scale:

$$v = (\nu \epsilon)^{\frac{1}{4}}.$$  \hspace{1cm} (3.2)$$

Similarly, the energy spectrum for turbulence can be arrived at by dimensional analysis. The energy input into the system is kinetic so,

$$E \propto v^2 = \frac{L^2}{t^2},$$  \hspace{1cm} (3.3)$$

where $L$ is the characteristic length. The rate at which the energy dissipates is

$$\dot{E} = \frac{E}{t} \propto \frac{L^2}{t^3}.$$  \hspace{1cm} (3.4)$$

If we define $k = 1/L$ as the wavenumber, and $\Delta k = k_1 - k_2$ as the interval then $E(k_1) \Delta k$ should have the dimension of the $E$, so the energy spectrum must have the dimension

$$E(k) \propto \frac{L^3}{t^2}.$$  \hspace{1cm} (3.5)$$

Assuming the energy spectrum, $E(k)$, $k$, and $\epsilon$ are linked we arrive at

$$E(k) = C \epsilon^x \cdot k^y,$$  \hspace{1cm} (3.6)$$

where $C$ is a dimensionless constant. Taking the dimensions we have:

$$\frac{L^3}{t^2} = \left( \frac{L^2}{t^3} \right)^x \cdot \left( \frac{1}{L} \right)^y.$$  \hspace{1cm} (3.7)$$
Separating $L$ and $t$ and solving the system for $x$ and $y$ gives

$$E(k) = C \varepsilon^{3/2} \cdot k^{-3/2},$$

(3.8)

a famous result of the Kolmogorov model and is supported by empirical data.

In Tatarski's model (Tatarski 1961) the fluctuations in phase introduced due to turbulence can be approximated by a Gaussian distribution and the turbulence can be described statistically by the so-called structure function:

$$D_\phi(r) = \langle |\phi(r) - \phi(r + r')|^2 \rangle .$$

(3.9)

Here $D$ is the average variance in phase between two points of the wavefront. In the case of Kolmogorov turbulence the structure function becomes

$$D(r) = C_n^2 r^{2/3},$$

(3.10)

where $D$ and $C$ are in terms of the index of refraction.

A measure of the effect of atmospheric turbulence of light is the Fried parameter, $r_0$ (Fried 1966b). The Fried parameter can be thought of as the dimension of the turbulent pockets in the atmosphere such that they have constant phase, meaning that the resolution limit equals $\lambda/r_0$.

The variance in phase induced by the atmosphere can be put in terms of the Tatarski structure function (Tatarski 1961) and the Fried parameter (Fried 1965),

$$D_\phi(r) = 6.88 \left( \frac{r}{r_0} \right)^{2/3},$$

(3.11)
where the Fried parameter (Fried 1966a) and can be defined as

\[ r_0 = 0.185 \lambda^{6/5} \cos^{3/5}(\zeta) \left[ \int C_n^2(h) \, dh \right]^{2/5} \]  

(3.12)

where \( h \) is the height above ground and \( \zeta \) is the zenith angle.

According to Fried (1965); Noll (1976), the Fried parameter can also be related to the variance of the wavefront phase, \( \sigma \), and the telescope aperture, \( d \), by

\[ \sigma^2 = 1.03 \left( \frac{d}{r_0} \right)^{2/3}. \]  

(3.13)

If a telescope aperture is smaller than \( r_0 \), it is diffraction-limited; telescopes larger than \( r_0 \) are seeing-limited. Typical seeing values at 0.5 \( \mu \)m for Mount Wilson are \( r_0 = 10 \) cm. This scales with wavelength by \( r_0 \propto \lambda^{6/5} \). Due to the differential paths taken through the atmosphere, phase information is lost to a two-telescope interferometer. To truly recover phase information at visible and NIR wavelengths requires three telescopes and phase closure measurements.

Atmospheric turbulence-induced differential paths for incoming starlight results in a phase shift of the wavefront. If these paths are rapidly changing then the OPD will shift rapidly. This shift of the wavefront is termed "atmospheric piston". Piston introduces a random delay to the OPD and causes the central fringe position to change. If the fringes move faster than the system can compensate, then data quality is degraded.
3.2 The Concept of Fiber Coupling in FLUOR

As shown in subsection 1.4.2 the general interferometric intensity for a monochromatic source is given by:

\[ I = I_1 + I_2 + 2\sqrt{I_1 I_2} |\gamma| \cos (2\pi \sigma x + \phi_{\text{obj}} + \phi_{\text{atm}}), \]

(3.14)

where \( I_1 \) and \( I_2 \) are the intensities of two light beams, \( \gamma \) is the complex coherence function, \( \phi_{\text{obj}} \) is the phase of the light from the object, \( \phi_{\text{atm}} \) is the phase introduced by the atmosphere, and \( \sigma \) is the wave number. The monochromatic fringe equation is shown here for simplicity, but the quasi-monochromatic form is also valid.

The measured complex coherence function, \( \gamma \), is the product of the object’s true visibility, the instrument transfer function, and an atmospheric turbulence term; so \( \gamma = T_{\text{instr}} T_{\text{atm}} V_{\text{true}} \).

In a traditional interferometer, the transfer function is determined by interleaving observations of the target star with those of an unresolved calibrator star. This leaves the atmospheric term which is a random variable that prevents the collection of high precision visibilities on the object. FLUOR deals with this term by utilizing single-mode fibers to spatially filter the individual beams. The fibers transform the phase variation into intensity fluctuations, which can be monitored by splitting off the input fibers into separate photometric channels. In FLUOR this is done by channeling each beam into an optical fiber Y-coupler, one output of the Y is one photometric channel, the other feeds an X-coupler. The end result of this is two photometric channels and two interferometric channels. Data from these photometric
channels are then incorporated into the data reduction process to produce a corrected interferogram.

The intensity at the input and output of the X-coupler are linked by a transmission matrix:

\[
\begin{pmatrix} I_1 \\ I_2 \end{pmatrix} = \begin{pmatrix} \kappa_{11} & \kappa_{12} \\ \kappa_{21} & \kappa_{22} \end{pmatrix} \begin{pmatrix} P_1 \\ P_2 \end{pmatrix},
\]

(3.15)

where here \( P \) is the measured photometric signal from each beam.

The \( \kappa \) matrix is a proportionality factor that incorporates the coupling and transmission efficiencies of the system; it can be extracted from the data without the need for prior knowledge of the individual transmission coefficients. For data reduction, both interferometric channels are reduced independently, as is their difference. When the recombination is incoherent the wide band interferometric signal can be expressed as

\[ I = \kappa_1 P_1 + \kappa_2 P_2. \]

(3.16)

See Coudé du Foresto et al. (1997) for a full discussion and derivation.

### 3.3 History

The method of using single-mode optical fibers for spatial filtering of the incoming beams can lead to very accurate measurements of interferometric visibility. This is the foundational principle behind the FLUOR instrument (See Figure 3.1).
FLUOR was first demonstrated at Kitt Peak Observatory with the 80-cm auxiliary telescopes of the McMath Solar Telescope. There, it recorded object visibilities with statistical errors of < 1% on a dozen stars. Later, the instrument was moved to IOTA on Mt. Hopkins. IOTA offered baselines of 7 to 38 m over the single 5.5 m baseline available at the McMath
telescope. FLUOR operated as a key instrument at IOTA from 1996-2001 before moving to the CHARA Array. Throughout its history, FLUOR has undergone numerous modifications to further improve its operating efficiency and precision; reaching visibility amplitude measurement precision of ±0.3% for bright sources (Coudé du Foresto et al. 2003).

3.4 The Instrument

![Diagram of the FLUOR bench layout](image)

Figure 3.2: The bench layout of FLUOR as it was setup initially at the CHARA Array. Two beams enter and are injected into fibers after one beam first reflects off the dither mirror. Two photometric channels are picked off and two interferometric channels are combined inside the fibers in “MONA”. The output of the four fibers is then imaged on the NICMOS detector.

On the FLUOR bench, each of the two beams first encounter an OPD stage that has a payload consisting of a set of two mirrors at 45° to the beam path (see Figure 3.2). Following these stages, each beam hits a fold mirror which directs the light to a two-axis tilting mirror that is actuated by stepper motors. The mirror directs the light onto a gold-coated f/1.3 30° OAP. The OAPs focus the light onto a fiber injection stage capable of XYZ translation. The
optical fibers are terminated in high precision E2000 fiber connectors. These connectors allow the fibers to be unplugged and replugged without the loss of alignment.

Figure 3.3: The MONA fiber beam combiner consists of two Y-fiber couplers and one X-fiber coupler outputting one photometric channel for each beam and two interferometric channels total. Fiber polarization can be adjusted by changing the amount of bend to the fiber in the two loops.

FLUOR interference fringes are produced by temporal OPD modulation using a mechanical stage that carries a mirror. The beam combination occurs within optical-fiber couplers. The fiber couplers are located inside a closed box system called “MONA” built by the company Le Verre Fluoré (LVF). The single-mode fibers of MONA have a core diameter of 8.5 µm and a Numerical Aperture (NA) = 0.17. NA is a dimensionless quantity that characterizes the range of angles for which light will enter a fiber, \[ NA = n \sin \theta = \sqrt{n_{\text{core}}^2 - n_{\text{cladding}}^2}. \]

Two injection stages feed light into two Y-fiber couplers which in turn feed one of their outputs into a X-fiber coupler (See Figure 3.3). Interferometric combination occurs in the X-fiber coupler. MONA outputs two photometric channels (PA and PB) from the Y-fiber couplers and two interferometric channels (I1 and I2) from the X-coupler. All four fibers join into a single fiber bundle in a 125µm square pattern. This bundle connects to another fiber translation.

\(^1\)E2000 is a trademark of Diamond company and complies with IEC 61 754-15 and TIA/EIA 604-16
stage and identical OAP. After the OAP, the light enters an objective lens that images the fiber bundle onto the NICMOS-based science camera for JouFLU (NICMOS), which is read out as four pixels. The difference is taken between the two interferometric pixels to increase the SNR while the photometric channels are recorded simultaneously for data calibration during the reduction stage (Coudé du Foresto et al. 1997).

### 3.4.1 Coupling to Single-mode Fibers

Single-mode fibers filter the wavefronts of incident starlight by passing only the fundamental mode ($LP_{01}$). This property of single-mode fibers serving as an apodization window is the defining feature of FLUOR; it is this, along with the monitoring of photometric channels, that makes FLUOR’s high precision possible (Coudé du Foresto 1994). The benefits of spatial filtering with fibers come at a cost. Photons are lost during fiber injection; sensitivity is sacrificed for precision. Just how much is lost depends upon the coupling efficiency, $\rho$, that comes from matching the starlight to the fiber heads. The coupling efficiency of starlight into single-mode fibers is discussed in Shaklan & Roddier (1988).

Single-mode fiber injection generally follows a Gaussian distribution while the intensity of starlight, over sufficient integration time to reduce speckle behavior, at the focus of a telescope more closely resembles a Bessel (or Airy) function (See Figure 3.4). This mismatch results in a maximum possible coupling efficiency of 82% for an unobstructed telescope. This is reduced by another 4% when Fresnel reflection is taken into account. This calculated maximum possible efficiency does not take into account the upstream beam train
nor the transmission properties of the fibers themselves.

This loss is further modified by the presence of atmospheric turbulence. As stated earlier, the fibers filter out atmospheric effects upon the phase of the wavefront; however, before reaching the fiber head, the atmosphere imparts a speckle pattern upon the starlight which greatly impacts the fiber coupling.

The normalized frequency, $V$, which characterizes the properties of an optical fiber is given by:

$$ V = \frac{2\pi}{\lambda_0} a NA. \quad (3.17) $$

For JouFLU the operating waveband is centered on $K'$, the $NA = 0.17$, and the fiber core radius, $a$, is $4.25 \mu m$. This gives the fibers a normalized frequency of $V = 2.1$. If $V$ is less than 2.4 then only a single mode of wavelength less than $\lambda_{cutoff}$ is guided by the fiber (Gloge 1971a,b).

### 3.4.1.1 Coupling Efficiency

The coupling efficiency, $\rho$, of optical fibers depends upon several factors intrinsic to the fiber interface with the beam: the physical parameters and alignment of the injection optics, the parameters of the optical fibers, and the beam f-ratio. These essentially determine the Airy function and the Gaussian that it overlaps. The resulting overlap integral sets the coupling efficiency (Neumann 1988; Pal 1992):

$$ \eta = \frac{\left[ \int_{0}^{r_{\text{fiber}}} E_b(r) E_g(r) r dr \right]^2}{\int_{0}^{r_{\text{fiber}}} E_b^2(r) r dr \int_{0}^{r_{\text{fiber}}} E_g^2(r) r dr}. \quad (3.18) $$
Here, \( E_g \) and \( E_b \) represent the Gaussian and the Bessel functions respectively.

After normalizing this becomes the coupling efficiency:

\[
\rho = \left| \int_0^{r_{\text{fiber}}} E_b \tilde{E}_g^* r dr \right|^2, \tag{3.19}
\]

where \( \tilde{E} \) is the complex conjugate of the electric field.

![Figure 3.4: The electric field across the plane of the fiber head. The acceptance mode of the fiber is modeled as a Gaussian (blue, dashed) and the telescope beam (black, solid) is an Airy function for \( f/d = 2.67 \). The solid and dashed lines are for an unobstructed beam and the dotted lines are the Gaussian and Airy functions for a beam with a 0.25 central obscuration.](image)

The coupling efficiency of optical fibers on telescopes can also be highly affected by the telescope central obstruction and atmospheric turbulence. For the CHARA telescopes, the primary mirror is obscured by the secondary in the amount of 1.6% of the primary mirror area. The beam diameter at the JouFLU bench is 19 mm. The OAP that injects light in to the fiber has a focal length of 50.8 mm. So for FLUOR, \( f/d = 2.67 \). The 78% maximum theoretical coupling efficiency is reduced only negligibly to due to the central obscuration in
In the presence of turbulence $\rho$ decreases with $d/r_0$. The best improvement for this by tip/tilt occurs when $d/r_0 = 4$. The coupling efficiency is reduced by typical atmospheric turbulence to 0.49.

The following cases from Shaklan & Roddier (1988) relate the impact of atmospheric turbulence upon coupling efficiency:

\[
\begin{align*}
    r_0 &= 2d, \rho \approx 0.74; \\
    r_0 &= d, \rho \approx 0.65; \\
    r_0 &= \frac{d}{2}, \rho \approx 0.45; \\
    r_0 &= \frac{d}{4}, \rho \approx 0.2.
\end{align*}
\]

At the CHARA Array, typical $r_0$ values are recorded as 10 cm or better in the visible band. This is $r_0 \approx 50$ cm in the K band and the corresponding $\rho$ is $\approx 0.45$. The falloff in coupling efficiency when going from good to poor seeing is rapid. FLUOR is able to record fringes at $r_0$ values $> 5$ cm in the visible; below that level data are generally of very poor quality.

The total power coupled into the fiber is proportional to $I(\pi/4)(d)^2\rho_{\text{max}}$ with the maximum power occurring when $d/r_0 = 4$. The best performance in terms of reduced visibility error occurs when $d/r_0 < 1$ (Guyon 2002). At CHARA this corresponds to $r_0 > 20$-cm in the visible.

### 3.4.1.2 Field of View

There are many factors which affect the Field-of-View (FoV) of an interferometer. If the beam combination occurs in the image plane, the isoplanatic patch limits field of view. This is the characteristic region over which the turbulence is the same. It can be defined by the
isoplanatic angle,

\[ \theta = 0.314 \cos(\zeta) \left( \frac{r_0}{H} \right), \]  

(3.20)

where \( \zeta \) is the zenith angle and \( H \) is the mean effective turbulence height (Hardy 1998),

\[ H = \left( \frac{\int dh C_n^2(h)h^5}{\int dh C_n^2(h)} \right)^{\frac{1}{5}}. \]  

(3.21)

Here, \( h \) is the height in the atmosphere.

In the pupil plane, a source off axis will have a phase difference. Once the difference is equal to the coherence length, there will not be fringes. The further away a source, the greater the difference in phase. A source \( \lambda/2B \) away will have a \( \pi \) phase difference. So, the FoV is baseline dependent, with longer baselines having a more restricted FoV. This is the dominant restriction to the FoV at long baselines, with the FoV limited by the fringe spacing times the coherency envelope (Thompson et al. 1986; Monnier 2003). The FoV can then be shown as

\[ \theta = \frac{\lambda}{B} \frac{\lambda}{\Delta \lambda} = \frac{\lambda^2}{B \Delta \lambda}, \]  

(3.22)

where \( \lambda \) is the wavelength, \( B \) is the baseline, and \( \Delta \lambda \) is the bandwidth.

A consequence of using fibers on an interferometer is a FoV restricted to a single telescope Airy disk. The coupling efficiency drops severely for a single fiber per aperture interferometer when a source is an angular distance of \( \lambda/d \) from the optical axis. For FLUOR this means the
full FoV is $2 \cdot 0.45''$:

$$\frac{\lambda}{d_{\text{max}}} = \frac{2.2 \times 10^{-6}m}{1m} = 2.2 \times 10^{-6} \cdot 206265 = 0.45''$$

$$\approx 0.9'' \text{ FWHM},$$

assuming Gaussian statistics (Guyon 2002). Much of the research done with FLUOR is on extended sources, such as the exozodiacal disks survey, and for these a shorter baseline is used, making the fiber injection a major factor in determining the FoV.

### 3.4.2 Piston Considerations

Piston is the only atmospheric disturbance not filtered out by the single-mode fibers. The effect of piston on the interferogram is to introduce random delay. This causes “jitter” in the fringe position. If the jitter is greater than the amplitude of the fringe scan or faster than the scan rate, then data is lost. Piston noise causes a loss of spectral information and of visibility phase.

One way around this problem is to implement an active fringe tracker. This tracks the position of zero OPD in another waveband and uses this feedback loop to introduce an offset in delay to counter the piston effect. Such an instrument exists, the CHARA Michigan phase-tracker (CHAMP) (Berger et al. 2008). It would be possible to utilize CHAMP to track fringes in the H-band, while operating FLUOR in the K-band. The initial steps for this capability have begun, the necessary dichroics to split the light between the two instruments
have been purchased, however more work is required to utilize CHAMP.

Guy Perrin has developed a statistical method to remove low frequency differential piston from interferometric data (Perrin 1997).

Currently, without a fringe tracker, FLUOR is limited to operation at 500 Hz. For most seeing conditions, this is adequate to maintain the fringe packet within the data scan recorded.

3.5 Uses and Key Results

FLUOR has contributed data to numerous publications in multiple fields of astrophysics. A few examples of topics from papers in various fields are presented here.

- **Stellar astrophysics** Angular diameters (Perrin et al. 2004), $T_{\text{eff}}$ (Perrin et al. 1998), limb darkening (Aufdenberg et al. 2008), hydrodynamical modeling (Chiavassa et al. 2010)

- **Rapid rotators** (Aufdenberg et al. 2006) rotation geometry, gravity darkening (Aufdenberg et al. 2007)

- **Asteroseismology** (Kervella et al. 2008; Mazumdar et al. 2009; Bruntt et al. 2010)

  P-mode oscillations

- **Circumstellar environment** extended envelopes (Mérand et al. 2007)

- **Binaries** Masses, orbital parallax, tidal effects, mass transfer, low mass or faint companions (Aufdenberg et al. 2009)

- **Be stars** (Touhami et al. 2007)
• **Debris disks and exozodiacaal dust** (Absil et al. 2006) and exozodiacal dust (di Folco et al. 2007)

• **Young star circumstellar environments** (Akeson et al. 2005)

• **Cepheid variables** P-L calibration (Kervella et al. 2004a) and Baade-Wesselink distance measures (Kervella et al. 1999), diameters and distance (Kervella et al. 2001) and pulsation modes (Gallenne et al. 2012)

• **Mira variables** pulsation properties (Mennesson et al. 2002)

• **High precision measurement of extended sources** (di Folco et al. 2007)

• **High dynamic range sources** (Absil et al. 2008a; Berger et al. 2008) Contrast ratios of $10^2$ to $10^6$ (5-15 magnitudes).
“Astronomy, as nothing else can do, teaches men humility.”

—— Arthur C. Clarke
4.1 The Need for Upgrade

In recent years, it has become clear that several improvements could increase the efficacy of FLUOR in terms of its efficiency, throughput, and integration with the CHARA Array. The JouFLU project, loosely translated as “rejuvenation” of FLUOR, has improved much of the optical bench setup of FLUOR to provide greater opto-mechanical stability. This includes new motorized mounts for the mirrors that feed light into the optical fibers, new higher-precision motorized stages that control the OPD to generate fringes, an OPD Scanning stage (OPD Scan), an OPD Static stage (OPD Stat), an infrared pupil-plane camera system, a visible light alignment camera, and improvements to the NICMOS science camera. OPD Scan is used for dynamic control of the path, while OPD Stat provides a static offset. Figure 4.1 shows some of the extent of the changes to the optical bench setup. In addition to the hardware upgrades, a concurrent replacement of the control software system has produced an entirely new software system that is compliant with the CHARA operating environment, enabling JouFLU
to be maintained at the forefront of improvements to combiners at the Array. As another combiner receives relevant new software features or tools, JouFLU may now benefit from them too with a minimum of effort. From these software changes, FLUOR now has remote operation capability, potential for greater science data throughput, and higher statistical precision.

Figure 4.1: left FLUOR shortly after its move to the CHARA Array. right JouFLU in its current condition in the CHARA lab.

4.2 Installation

After design and qualification testing at Laboratoire d'études spatiales et d'instrumentation en astrophysique (LESIA) during 2011, the JouFLU components were shipped to Mount Wilson for integration at the array. The JouFLU parts arrived in late January of 2012. The installation was completed by the end of February, by Emilie Lhomé with the assistance of Nicholas Scott (Lhomé et al. 2012). During the last two weeks of March 2012, JouFLU was commissioned and got first on-sky fringes.
The process of installing JouFLU began with the complete removal of all components on the FLUOR optical bench. Once the bench was cleared and cleaned, the JouFLU components were positioned and all cables routed. Concurrently with the hardware installation, numerous software changes were introduced and tested. Stage control and camera timing control systems were implemented. Once the final software replacements and optically alignments were completed laboratory fringes were obtained.
Figure 4.2: This timeline shows the progression of the JouFLU upgrades from initial software improvements to the recording of science data.
4.3 Components

Figure 4.3: JouFLU in the configuration used when observing. The alignment stages are out of the path of the beam, and the OUTPUT stage is moved to the open position. Also shown for completeness are the optics of the alignment portion of the bench, see Figure 4.22 for their use. The FTS beamsplitter is removed, see Figure 4.29.

The Jouvence of FLUOR series of upgrades touched nearly every component of the FLUOR bench. Numerous mechanical stages have been added as part of the upgrade, necessitating new controllers and entirely new software. While most of the improvements from the JouFLU project center on new mechanical stages, optical elements were added to improve the instrument and to provide for new alignment techniques. Multiple optical components were added, adjusted, or changed. Some general descriptions of the effect of these changes will be listed here, with details parsed into sections based on the function of the component. Following this, motion control and alignment instruments, which affect the entire instrument, will be explained.
4.4 Input

For the purposes of this document we define “input” as everything upstream of the fibers themselves.

4.4.1 M0 implementation

The CHARA Array delivers two beams to the JouFLU optical bench. These are ordered in pairs and may be either beams 3 & 4 or beams 5 & 6. Beam 3 or 5 is termed beam A and beam 4 or 6 is beam B. The beams first reach the JouFLU optical bench with a known offset in delay space of $A - B = -284$ mm. This delay comes from the Visbeams combiner earlier in the CHARA system and is a result of the 3 inch separation between the beams and the $30^\circ - 60^\circ - 90^\circ$ triangular geometry of the combiner.

The first optical element the beams reach is stage M0 located at the corner of the adjacent MIRC table. This stage consists of an elevated optical breadboard with four magnetic bases located on its underside. These magnetic bases are the positions for two removable fold mirrors that hang downward so that they intercept either beam A or Beam B. The placement of the M0 mirrors provides additional OPD to beam A, $B3 - B4 = 140 + 76$ mm. The beams are directed at $90^\circ$ to the JouFLU optical table with beam A on the inward side and beam B closest to the table edge. The beam axes are horizontally separated by 140 mm.

In addition to the mirror placements the M0 stage can instead hold dichroics to allow use of JouFLU with the CHAMP for fringe tracking. These dichroics bend the K-band (2.2 $\mu$m)
portion of a beam 90 degrees while transmitting the H-band (1.5-1.8 µm) light.

Table 4.1: Physical and Spectral specifications for the M0 dichroics.

<table>
<thead>
<tr>
<th>Physical Specifications</th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>Size:</td>
<td>50.8 +0/-0.2 mm</td>
</tr>
<tr>
<td>Thickness:</td>
<td>10.0 mm nominal</td>
</tr>
<tr>
<td>Edge Treatment:</td>
<td>Ground Edge</td>
</tr>
<tr>
<td>Parallelism:</td>
<td>30 seconds or better</td>
</tr>
<tr>
<td>Flatness:</td>
<td>1/20 wave @ 633 nm before coating</td>
</tr>
</tbody>
</table>

| Spectral Characteristics:     |       |
| AOI:                          | 45 deg |
| Cut-on:                       | 1850 nm nominal |
| Transmission:                 | ≥ 85% 1490 – 1780 nm |
| Reflection:                   | ≥ 95% 2030 – 2370 nm |

In order to compensate for the differential polarization phase delay discussed in section 4.11, Lithium Niobate plates will be added to the CHARA optical path by the end of 2015. These plates have the dimensions: 40 x 30 x 4 mm and are Anti-Reflective (AR) coated on both faces. There are plans to purchase separate plates with coatings optimized to cover H or K wavebands. These plates are mounted on rotation stages that allow the projected thickness of the plate in the optical beam to be varied, resulting in an adjustable amount of polarization phase delay for each beam.

4.4.2 Fiber Injection Tip/Tilt Stages

After entering the JouFLU bench and passing through the OPD Scan and OPD Stat stages, the beams encounter a final fold mirror before hitting Zaber model T-MM actuated two-axis tip/tilt stages. These stages each hold a single flat mirror as payload that directs the light from each beam onto an OAP. To ensure the optimal injection of light into the optical fibers by the OAPs, the Zaber stages perform a raster scan. During this procedure, the I1 and I2 pixels on
the science camera (See Figure A.11) are summed for each position of the Zaber as it performs a line search about a square region. The result of this is effectively an image of the fiber core itself.

The previous mirror mounts that directed light onto the OAPs and into the optical fibers were actuated by stepper motors. These stepper motors were not sufficiently precise to produce repeatable alignment for fiber injection. These mirror mounts have been replaced by Zaber stages. The minimum step size of these stages is 310 mas or $\approx \frac{1}{6}$ of the 8.5 $\mu$m core diameter of the fibers. In practice, a step size of $\approx 1.5$ $\mu$m is used to enhance repeatability of raster scans. This is a large improvement over the previous stepper motors that had a step size roughly equal to the fiber core diameter. The higher-precision injection allows for accurate raster scanning to maximize the amount of light that reaches MONA. The Zaber stages are also much faster than the previous stepper motors, allowing for more rapid and larger raster scans (See Figure 4.4 for an example raster scan).

The raster scan is typically performed as an 11 x 11 search for coarse alignment to find light on the fiber. This is then followed by a finer 5 x 5 search (See Figure 4.4). Once the spot is found, the Zaber may be centered on the maximum value or the Center of Gravity (CoG). Optimal positions may be stored and saved from one observation to the next and default positions for these stages may be set depending on the configuration of the instrument.

Additionally the size of each step in the scan may be defined as follows:
\[ d\theta = 0.00008528^\circ \]

\[
\text{min step size on OAP} = 2 \times \text{focal length of OAP} \times \sin \frac{d\theta}{2}
\]

\[
\text{Zaber motion at fiber (\(\mu\)m)} = \text{zaber microstep size} \times 2 \times \text{min step size on OAP}.
\]

The factor of two in the final step comes from the reflection of the mirror mounted on the Zaber. The default Zaber microstep value is 40 but this may be reduced to 20 for great seeing conditions. Twenty microsteps gives a motion of 3 \(\mu\)m at the fiber, 40 microsteps gives 6 \(\mu\)m motion at the fiber. The minimum useful step size is 10 zaber steps, this gives a physical step of 1.5 microns or \(1/5.67\) of the fiber diameter.

The quality of the raster scan is indicative of atmospheric seeing conditions, the telescope focus, and the OAP alignment and focus.

Figure 4.4: left The resulting image from a 5x5 raster scan of the CHARA white light source. right The results of a Gaussian fit to the raster scan.
Table 4.2: Physical specifications for the OAPs.

<table>
<thead>
<tr>
<th>Physical Specifications</th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>Off-Set Angle</td>
<td>30°</td>
</tr>
<tr>
<td>Diameter (mm)</td>
<td>50.8</td>
</tr>
<tr>
<td>Diameter Tolerance (mm)</td>
<td>+0.00/-0.38</td>
</tr>
<tr>
<td>Focal Length Tolerance (%)</td>
<td>±1</td>
</tr>
<tr>
<td>Surface Figure, RMS</td>
<td>λ/4</td>
</tr>
<tr>
<td>Parent Focal Length PFL (mm)</td>
<td>50.8</td>
</tr>
<tr>
<td>Effective Focal Length EFL (mm)</td>
<td>54.45</td>
</tr>
<tr>
<td>Surface Roughness (Angstroms)</td>
<td>&lt; 175 RMS</td>
</tr>
<tr>
<td>Substrate</td>
<td>Aluminum 6061-T6</td>
</tr>
<tr>
<td>Coating</td>
<td>Protected Gold</td>
</tr>
<tr>
<td>Y Offset (mm)</td>
<td>27.22</td>
</tr>
<tr>
<td>Wavelength Range (µm)</td>
<td>0.7 - 2</td>
</tr>
</tbody>
</table>

4.4.3 Off-axis Parabolas

Each Zaber motorized mount directs light onto a two-inch diameter, gold-coated f/1 OAP.

This OAP focuses the starlight onto the single-mode fiber core that is the input for MONA.

The efficiency of the JouFLU system is highly dependent upon the quality of this injection.

For the JouFLU upgrade, the original FLUOR OAPs have been replaced and re-aligned with new OAPs that meet or exceed the original specifications.

4.5 Combination

Combination includes the interferometric beam combination and the OPD control to enable interference.

4.5.1 MONA

At the heart of FLUOR and JouFLU is the beam combination itself. This is done inside a component called MONA. Two fluoride glass (CaF$_2$) single-mode fibers lead from their
respective input injection stages, bring light into MONA. These fibers have a 8.5-µm core and 125-µm cladding with a LP_{11} cutoff mode = 1.95-µm.

MONA is a triple coupler system: 2 Y-couplers (truncated X-couplers) designed to split the incoming light 80/20, with 20% going to the photometric channel. A single X-coupler splits the remaining light 50/50 to two interferometric outputs. The actual flux for each channel measured in the laboratory with the CHARA laboratory white-light source (WL) is:

\[
\begin{align*}
\text{p1} & : 4200/0.90 = 4667 \\
\text{i1} & : 2100/0.80 = 2625 \\
\text{i2} & : 4100/0.75 = 5467 \\
\text{p2} & : 900/0.72 = 1250
\end{align*}
\]

This is the number of counts in a single target pixel divided by the fraction of light in the target pixel compared to the surrounding pixels, which gives the total counts for the spot on the detector. The counts for each pixel and the fractional flux are displayed by the JouFLU server. The fractional flux, f, is calculated by \( \frac{f_{\text{target}}}{f_{\text{spot}}} \), where \( f_{\text{target}} \) is the flux in the target pixel and \( f_{\text{spot}} \) is sum of the counts in the target pixel and the 8 pixels adjacent to it. Channel I2 has roughly twice the flux of I1; and P1 has approximately 3.7 times the flux of P2.

From Equation 1.21, one can build the corrected interferogram for using the photometric signals, \( P_1 \) and \( P_2 \):

\[
I_{\text{corr},1} = \frac{I_1 - P_1 - P_2}{2\sqrt{P_1 P_2}} = \mu \cos (2\pi \sigma \Delta x + \Phi),
\]

(4.1)

where \( \mu \) is the coherence factor.
Including the previously defined $\kappa$ terms (Equation 3.15) this becomes

$$I_{\text{corr},1} = \frac{I_1 - \kappa_{11} P_1 - \kappa_{12} P_2}{2\sqrt{\kappa_{11}\kappa_{12} P_1 P_2}},$$

(4.2)

for one interferometric channel. Repeating this for the other channel and inserting both into the general monochromatic interference law leads to

$$I_{\text{corr}} = \frac{I_1 - \kappa_{11} P_1 - \kappa_{12} P_2}{2\sqrt{\kappa_{11}\kappa_{12} P_1 P_2}} + \frac{I_2 - \kappa_{21} P_1 - \kappa_{22} P_2}{2\sqrt{\kappa_{21}\kappa_{22} P_1 P_2}} + \sqrt{\left(\frac{I_1 - \kappa_{11} P_1 - \kappa_{12} P_2}{2\sqrt{\kappa_{11}\kappa_{12} P_1 P_2}}\right) \left(\frac{I_2 - \kappa_{21} P_1 - \kappa_{22} P_2}{2\sqrt{\kappa_{21}\kappa_{22} P_1 P_2}}\right)} |\gamma_{12}| \cos (2\pi \sigma \Delta x + \Phi_{12}).$$

(4.3)

The real-time monitoring of the photometric channels enables extremely high precision visibility amplitude measurements (Perrin 2003).

Figure 4.5: The internal layout of the MONA fiber-based beam combiner on FLUOR and JouFLU. It consists of two Y-fiber couplers and a single X-coupler. One output of each Y-coupler exits MONA, the other arm passes through the polarization controlling loops before entering the X-coupler. Note: here the photometric outputs, PA and PB, are equivalent to P1 and P2, respectively.

Adjustment of the polarization state is done through the use of “Mickey ears”, loops of the fibers that can be rotated to induce stress birefringence in the fiber (Lefevre 1980). These
loops are controlled via knobs on the side of MONA. When the knobs are at 0, the fiber is least constrained. The Mickey ears are on a 10:1 gear reduction, so ten turns on the knob is one turn of the fiber. The dial indicator for the knob has a maximum value of slightly over 600, which corresponds to 6 turns of the fiber.

The fibers are connectorized with E2000 UPC connectors. Four female connections are on the rear side of MONA. Each feeds one of the four channels to a fiber bundle. The fiber bundle brings the four fibers together so they share a common output ferrule. The cores are aligned in a square with 125 $\mu$m per side. This output is mounted on an output stage similar to the ones at the input side. Another OAP of the same type collimates the beam and directs it to a stationary flat mirror. An objective lens then images the beam onto the detector where each channel is read as a single pixel on the chip.

4.5.2 Optical Path Delay

JouFLU combines two beams pair-wise from CHARA; so JouFLU can take either beams 3 & 4 or beams 5 & 6. To simplify discussion, beam 3 or 4 is referred to as beam A and either beam 5 or 6 is beam B. The CHARA beams have a differential offset prior to reaching the FLUOR table, such that $A - B = -284$ mm. Placement of the M0 mirrors for JouFLU alters the optical path, $A - B = 140$ mm + 76 mm. Internal to the JouFLU optical table, the placement of the stages and optics changed the path $A - B = 170 + 98$ mm.

It is important to match the amount of dispersion as much as possible as a mismatch results in a reduced visibility measurement. For JouFLU, the primary issue is chromatic dispersion
and is the result of inhomogeneities within the fibers. Chromatic dispersion affects the phase of the Fourier transform of the fringe. Large amounts of chromatic dispersion can arise in relatively short lengths of optical fiber (Coudé du Foresto et al. 2001). To compensate for this, when MONA was assembled the differential chromatic dispersion was measured and the lengths of the fiber arms of the combiner were adjusted. As a result of these fiber length adjustments to equalize dispersion in the two arms of the combiner, there is a differential offset internal to MONA of, \( A - B = -200 \text{ mm} \). The final path length result is:

\[
A - B = -284 + 140 + 76 + 170 + 98 - 200 = 0 \text{ OPD}.
\]

A change in OPD is needed for the FTS mode due to the reflection of the input beam from the beam splitter to the FTS mirror of 140mm. So for FTS mode,

\[
A - B = 170 + 98 - 200 - 140 = -72 \text{ mm}.
\]

This is compensated for by moving the OPD Stat stage by 36mm.

### 4.5.2.1 OPD Scan and OPD Stat

There are two OPD stages: OPD Scan, a dynamic scanning stage that modulates the OPD and generates fringes within MONA and OPD Stat, an adjustable static stage to correct residual OPD. These stages each carry a pair of mirrors in a dihedral arrangement. Movement of one of the OPD stages along the axis of the beam results in a change in the path length of twice the amount the stage moved.

The scanning stage meets rigid requirements as to linear velocity stability over its full range of travel. This stage was tested for such stability while in Meudon and achieves \( \approx 1\% \) error in its
velocity at 110 $\mu$m/s (See Figure 4.6). In addition, further custom tuning was performed by a Newport technician. The stage is actuated by a linear DC motor and has 50 mm of travel. The greater range of travel for this new stage greatly surpasses the 200 $\mu$m of the FLUOR piezoelectric dither mirror. The increased range is necessary for the use of the FTS mode. During normal observation mode and while collecting fringes at 100 Hz, the scanning stage travels at 105 $\mu$m/s (half the optical path velocity due to double pass) over a range of 150 $\mu$m. For FTS mode, the stage travel range must be 10 times this. The exact velocity of the OPD Scan stage is determined by the NICMOS camera readout frequency. Fringes are temporally modulated and scanned at a rate of five samples per fringe (2.5 times Nyquist). This rate was chosen based on experienced learned with the CLASSIC beam combiner; five samples per fringe produces data that can be well calibrated; more than five samples per fringe does not improve the data quality. So, the OPD Scan stage velocity is determined by

$$\text{Vel}_{\text{scan}} = \frac{R_{\text{camera}} \cdot \lambda_0}{2 \cdot N_{\text{sample}}}$$

(4.4)

where, in practice, $R_{\text{camera}} = 500$ Hz, $N_{\text{sample}} = 5$, $\lambda_0$ is the central wavelength, and the factor of 2 is due to the double pass of the beam on the stage.

While collecting fringes at 100 Hz NICMOS reads out at 500 Hz (2 ms). The Newport XPS Motion Controller (XPS) has been programmed to send a signal to the JouFLU control computer to report when the OPD Scan stage is moving at a constant velocity. However, this method of triggering the camera was found to be unnecessary and goes unused in order to
decrease the overhead on the JouFLU computer. Instead, the motion of the OPD Scan stage and the readout of the camera are synchronized by a fixed time delay. The acceleration period of OPD Scan is constant, and has been measured, for a given rate so the data recording sequence is triggered only after that delay. This ensures that the fringes are only recorded under the constant velocity situation and not when the stage is accelerating.

The second stage, OPD Stat, is a static stage and does not move during data acquisition. OPD Stat does not have the strict velocity requirements of OPD Scan, but it does need longer

![Figure 4.6: This plot shows the results of testing the OPD Scan stage’s motion. Plotted is the stage position (dashed line) and the stage velocity for a scan of 250-µm at 110-µm/s (solid line). This shows the rapid acceleration period and velocity stability of the stage.](image)

The parameters during OPD Scan component qualification testing:

- **scan length** = 0.500 mm
- **absolute value of maximum velocity** = 0.116 mm/s
- **mean velocity** = 0.109 mm/s
- **set velocity** = 0.110 mm/s
- **mean velocity rms error** = 1.33%.

The second stage, OPD Stat, is a static stage and does not move during data acquisition.
travel. **OPD Stat** corrects for the offset created by introduction of the **FTS** beam splitter.

Approximately 4 cm of stage difference are introduced in **FTS** mode due to the separation of the two beam paths (see **Figure 4.3**).

### 4.6 Output & Detection

After combination in fibers, there are various elements that direct the light from MONA after the fiber bundle output to the science camera. It is here in the collimated beam that the instrument can be configured for normal observations, spectral dispersion of the beam, or to direct a source back upstream for alignment purposes (see **Figure 4.7** and **Figure 4.8**).

![Figure 4.7: The OUTPUT stage in the spectrally dispersed data collection mode.](image)

The light from the output stage is incident upon a science detector. From the output fiber bundle, four pixels are illuminated representing the two interferometric signal channels and the two photometric channels. These channels are recorded by four single pixels on the detector. If observing in spectrally dispersed mode, instead of single pixels a region consisting of multiple sub-channels is detected. For JouFLU, two different science cameras
Figure 4.8: The OUTPUT and ALIU stages in the alignment configuration. The retro-injected LED can be imaged on the viscam detector to conjugate the beams by adjusting the alignment of the Zaber tip/tilt stages to overlay the image of the red LED with the CHARA laboratory WL source.

have been implemented. The original NICMOS-based camera was replaced by the
PICNIC-based science camera for JouFLU (CALI) as part of the upgrade program. Later, it
was determined that a return to the original NICMOS camera was advantageous.

4.6.1 Camera CALI

CALI was the originally implemented science camera for JouFLU. It is a 256x256 array
PICNIC Complementary Metal Oxide Semiconductor (CMOS) Mercury Cadmium
Telluride (HgCdTe) detector operating in the NIR. As specified it would have been able to
read any number of arbitrary sized Region(s) of Interests (ROIs) at a maximum readout rate
of 10 kHz for four single pixel windows.
Table 4.3: Detector specifications for CALI.

<table>
<thead>
<tr>
<th>Detector Specifications</th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>Technology</td>
<td>PICNIC CMOS HgCdTe</td>
</tr>
<tr>
<td>Spectral range</td>
<td>1.65 - 2.5 $\mu$m</td>
</tr>
<tr>
<td>Array format</td>
<td>4 - 128x128 quadrants</td>
</tr>
<tr>
<td>Pixel pitch</td>
<td>40 $\mu$m</td>
</tr>
<tr>
<td>Pixel readout rate</td>
<td>250 kHz</td>
</tr>
<tr>
<td>Frame rate</td>
<td>14 Hz</td>
</tr>
<tr>
<td>Pixel operability</td>
<td>99.94%</td>
</tr>
<tr>
<td>ADC</td>
<td>16 bits</td>
</tr>
<tr>
<td>Out voltage</td>
<td>0 - 0.5 V</td>
</tr>
<tr>
<td>Quantum efficiency (QE)</td>
<td>70% @ 2 $\mu$m</td>
</tr>
<tr>
<td>Readout noise (RON)</td>
<td>18.4 e-</td>
</tr>
<tr>
<td>Dark current</td>
<td>0.084 e-/s</td>
</tr>
<tr>
<td>Gain</td>
<td>307 e-/mV</td>
</tr>
<tr>
<td>Well capacity</td>
<td>&gt;150,000 e-</td>
</tr>
</tbody>
</table>

4.6.2 Issues with CALI

Upon receipt and installation of CALI, it became clear that the camera is not optimized to work at high speed (1 kHz) due to a reset anomaly; resetting the chip, when running at high rates, results in a large spike of read noise. Read noise is caused by thermal electrons that result from the movement of charge across the detector. It is uncertain why this camera displayed excessive read noise when operating in this mode, but it is thought to be a problem with the camera electronics. Also, the required integration time of 2 ms (for 500 Hz) readout lies outside the linear response of the chip. As a result of this, the calibration of the pixels within the reset anomaly is not obvious. The values given for the flux inside the reset anomaly zone are very difficult to compare with the number of photons. The reset anomaly affects all pixels and lasts between 50 and 100 ms. Functionally as a result of this reset issue, along with additional read timing inefficiencies, a significant loss of sensitivity was noticed when CALI was integrated into the JouFLU system. As a result, the on-sky performance was
limited to only a handful of the brightest targets. Attempts were made by engineers in Meudon to reprogram a new camera readout control card, but subsequent tests showed little improvement to the sensitivity of the camera.

4.6.2.1 Timing tests

The first tests to be performed with the new card were of the timing in the form of a chronogram. The oscilloscope was connected to the the electronics that control the readout of the CALI camera (Pautron) output and various settings of the number of times the chip is read (Nloops), the number of times a pixel is read (Nreads), and exposure duration were measured. The maximum Nloops and Nreads possible with a typical 2 ms (500 Hz) readout of 4 pixels, “scan” Region(s) of Interest (ROI), of the chip and a 100 ms “movie” mode (1 1 18 16 ROI) was determined. These were measured under a plan discussed in Meudon of using the shortest possible exposure (exposure number 0 which corresponds to 147 µs) and increasing the number of loops to gain integration time. For the 2-ms scan ROI, the maximum number of loops was 23. For the 100-ms movie mode, the maximum number of loops was 42. For any case, the maximum number of reads allowable by the camera is 4. All test plots show the results of tests using the 2 ms, 4 pixel “scan” region of interest, which is the expected mode for use on sky.
Figure 4.9: No light on left, light on right. Mean counts plotted against the variance as exposure was increased from 0 to the maximum of 7 while the CHARA incandescent lights were held constant. High counts were found even in the absence of light incident upon the camera. An accurate measure of the read noise (y-intercept) and gain (slope) was unable to be determined. This is partly due to the instability of the camera in this readout configuration and as well as to the problem of high counts and sometimes large negative counts.

Figure 4.10: No light on left, light on right. Mean counts plotted against the variance as the number of loops were increased from 2 to the maximum of 23 while the CHARA incandescent lights were held constant. High counts were found even in the absence of light incident upon the camera. An accurate measure of the read noise (y-intercept) and gain (slope) was unable to be determined. This is partly due to the instability of the camera in this readout configuration and as well as to the problem of high counts and sometimes large negative counts.

4.6.2.2 Mean counts vs. Variance plots

Figure 4.9, Figure 4.10, and Figure 4.11 show the results of tests on the camera with the light off and camera aperture covered and with the CHARA lab overhead incandescent lights on and the camera aperture open.
Tests show that the camera operating in four-pixel readout mode produces a large, unstable read noise as well as poor dynamic range. Some of this behavior was predicted by the engineers in Meudon who developed the Pautron camera control card. However, the degree of these problems and on-sky performance of the camera was uncertain until these tests. The conclusion is that the camera is not ready for use on sky at this time, and an alternative option of adapting the previous NICMOS camera for use was implemented.

It is worth noting that in the course of these tests the performance of not only the new electronics but also the old card was poor. Both cards had very high noise, large numbers of counts in the absence of light, and numerous hot or poorly responsive pixels. The performance of the old card seems significantly degraded since it was last used on sky in May 2012. The conclusion is that there is some additional problem with the camera affecting both electronics systems.

Troubling issues when operating at 500 Hz:
High counts with or without light
High noise
Many hot or stuck pixels
Large negative counts
High reset noise
Reset Anomaly
Poor sensitivity
Degraded performance

A persistent problem with the camera is the forced choice between acceptable stability and a level of sensitivity usable on sky. The parameters that result in an acceptable amount of noise sacrifice so much integration time that the sensitivity of the camera is poor. It is possible that this is a result of the necessary readout mode where we read individual pixels instead of the entire chip. A resolution to this would require more work on the electronics card and the camera/electronics control by the electronics engineers.

It was determined in January 2013 that the camera was not ready to perform on sky under the currently planned mode. The new electronics do not represent a great enough improvement in the camera’s overall sensitivity or effectiveness. In order to not sacrifice an observing season it was decided to switch back to the NICMOS camera. This change was completed within two weeks of work on the mechanics and software with the cooperation of Theo ten Brummelaar, Lazlo Sturmann, and Judit Sturmann. Concurrent with this, a solution to the low readout bandwidth of NICMOS was implemented. The serial readout electronics were replaced with a faster Ethernet connection similar to the Near InfraRed Observer (NIRO) at CHARA (ten Brummelaar et al. 2013). The addition of this connection should provide a
significant improvement in the efficiency of the NICMOS camera over its previous performance on FLUOR before it was replaced by CALI. This increase in readout speed enables the NICMOS camera to support a spectrally dispersed mode with a maximum of 5 spectral channels at 500 Hz.

4.6.3 Camera NICMOS

![Image](image.png)

Figure 4.12: “movie” mode. This display gives live visual feedback from the NICMOS camera. The amount of flux in each pixel from the fiber bundle is displayed along with alignment cues. The green boxes are the four actively read pixels. The red crosses demarcate the centroid of the light in the region near each pixel. The green cross and red box in the center represent the mean centroid of all four spots. This enables accurate alignment of the camera with respect to the output stage. The three grey pixels are known bad pixels and are forced to a single value.

The four outputs (2 interferometric and 2 photometric) of the MONA combiner are imaged onto four pixels of a NICMOS3 Focal Plane Array (FPA) HgCdTe detector, housed in a camera
(referred to as NICMOS) originally developed for the IOTA interferometer by Millan-Gabet et al. (1999). We use the same dewar, readout electronics and control software approach as in the original implementation. The NICMOS3 is a CMOS Field-Effect Transistor (FET) switch array, developed by Rockwell International Science Center, and arranged in four 128x128 pixel quadrants. In 2007 however, the original NICMOS3 array failed and was replaced by another engineering grade NICMOS3 array, kindly loaned by National Optical Astronomy Observatory (NOAO). Although the replacement array has a larger number of bad pixels, which can easily be avoided, the noise characteristics remain similar. Camera control has been integrated into the CHARA environment as described in section 4.10. The main JouFLU CPU coordinates the Newport XPS and the MS-DOS machine that communicates with NICMOS. A software delay triggers reading of the camera so that data are only collected when the fringe-scanning stage (OPD Scan) is moving with constant velocity. These data are then sent to the JouFLU computer for real-time display and recording.

Table 4.4: Detector specifications for NICMOS (Millan-Gabet et al. 1999).

<table>
<thead>
<tr>
<th>Detector Specifications</th>
<th>NICMOS3 FPA HgCdTe</th>
</tr>
</thead>
<tbody>
<tr>
<td>Technology</td>
<td></td>
</tr>
<tr>
<td>Spectral range</td>
<td>0.9 - 2.5 ( \mu \text{m} )</td>
</tr>
<tr>
<td>Array format</td>
<td>4 - 128x128 quadrants</td>
</tr>
<tr>
<td>Pixel pitch</td>
<td>40 ( \mu \text{m} )</td>
</tr>
<tr>
<td>Pixel clock rate</td>
<td>400 kHz</td>
</tr>
<tr>
<td>Pixel operability</td>
<td>98% for ROI</td>
</tr>
<tr>
<td>ADC</td>
<td>16 bits</td>
</tr>
<tr>
<td>Out voltage</td>
<td>0 - 0.5 V with 3 V offset</td>
</tr>
<tr>
<td>QE</td>
<td>50% for 0.9-2.5 ( \mu \text{m} )</td>
</tr>
<tr>
<td>RON</td>
<td>16.6 e- (double read)</td>
</tr>
<tr>
<td>Dark current</td>
<td>0.2-100 e-\text{/s}</td>
</tr>
<tr>
<td>Gain</td>
<td>3.2 e-/Analog to Digital Unit (ADU)</td>
</tr>
<tr>
<td>Well capacity</td>
<td>300,000 e-</td>
</tr>
</tbody>
</table>
Figure 4.13: The results of gain and readout noise tests of the NICMOS camera for each of the four read pixels. The camera was operated in destructive mode at 500 Hz with 2 loops and 3 reads of each pixel. Measurements were determined by calculating mean counts and variance as light levels were incrementally increased. The first 200 counts were treated as linear, and a regression was performed to determine the gain. Mean readout noise for the four pixels read is 2.56 ADU with a mean gain of 5.12 $e^{-}$/ADU.

In Figure 4.13 we show the results of updated measurements of the camera gain and readout noise.

The total noise, in ADU for a camera is given by

$$N = \sqrt{N_{\text{read}}^2 + N_{\text{shot}}^2 + N_{\text{fixed}}^2}$$

where $N$ is the total noise, $N_{\text{read}}$ is the camera readout noise, $N_{\text{shot}}$ is the shot noise from the photon flux and $N_{\text{shot}} = \text{counts/gain}$, and $N_{\text{fixed}}$ is the Fixed Pattern Noise (FPN) from the non-uniform response of the chip and scales with the flux. The signal shot noise equals the square root of the signal, $S$, and the FPN is proportional to the signal, $N_{\text{fixed}} = S \times PRNU$. Photo Response Non-Uniformity (PRNU) is the difference between the measured response from a detector and a uniform response, and is related to FPN. Plugging this into
Equation 4.5 yields:

\[ N = \sqrt{N_{\text{read}}^2 + S + (S \times PRNU)^2}. \]  

(4.6)

Measurements of these noise characteristics can be obtained by recording the camera output under varying amounts of photon flux. The mean number of counts can then be plotted against the variance or standard deviation. When plotted, the y-intercept gives the read noise, and the slope of the curve gives the inverse of the gain. The slope is linear if it is dominated by shot noise. Flat-fielding removes the FPN, which is from inhomogeneities between pixels.

The readout mode used for lab testing is the same as used on-sky. As in previous implementations, we use multiple sampling for noise reduction: the array is reset for each sample, after which the pixels of interest are sampled continuously as they discharge (Fowler & Gatley 1991).
The integration time for each recorded data point is set by the time needed to sample each target pixel, possibly multiple times. For a given camera read frequency, an optimum number of \texttt{Nreads} and \texttt{Nloops} for the pixels is determined to give the correct sample time and minimize read noise. For example, to get a 500 Hz camera readout the four pixels are read out 11 times in one loop. This combination was found to give a 2 ms readout with minimal readout noise. The timing of pixel readout for various modes was checked by oscilloscope. Figure 4.14 shows noise reduction as a function of the number of multiple reads per sample, i.e., per data point in the fringe scan.

\texttt{NICMOS} has two readout modes: a movie mode runs at about 8.1 Frames Per Second (\texttt{FPS}) and the data collection mode. Figure 4.15 shows a large spike in the background noise level of the camera. This occurs regularly during the movie readout mode at roughly 12 second intervals and in all conditions. Nothing is known to have that period on the mountain. One plausible explanation for this behavior is electronic interference from distant weather, air traffic, or military radar sweeps. Radar sweep has been known to affect astronomical instrumentation and would match the periodicity of the observed noise pattern.

The data collection mode by default runs at 500 Hz but can be 125 Hz, 250 Hz, 500 Hz, 750 Hz, or 1000 Hz. The decision of what rate to collect data requires a balance of two factors: the signal and the piston. If the object pair, target and calibrator, is bright, a short integration time is preferred. The longer the integration time, the greater the signal, which depends upon the brightness of the object and the seeing conditions. However, if the piston of the atmosphere is bad, then a long integration time will result in loss of the fringe packet from the
data scan. The integration time must be long enough to collect a strong signal, but short enough to “freeze” the piston.

This mode reads four pixels representing the P1, S2, S1, and P2 channels from MONA. NICMOS can be operated in destructive or non-destructive modes. Destructive mode is a series of RESET-READ-READ sequences. The difference between the two consecutive reads yields the number of counts. As the chip is reset between every sample, this method is less prone to saturation on bright targets. The downside of this approach is that each reset adds noise. Non-destructive mode has a single reset at the beginning of the record sequence followed by many reads. The measured counts are the difference between consecutive reads. This method avoids the additional reset noise, but the initial reset is noisier. Also, the electron well is finite and will eventually saturate if the sequence is too long.

Figure 4.16 shows a map of the counts of each pixel for one frame of NICMOS in movie
mode with no light incident upon the camera. Figure 4.16 also shows the median counts and SD of each pixel in movie mode for 100 frames. The top of the frame shows higher counts because it is the location of the readout shift register; where it is warmer and more thermal electrons are generated. Figure 4.17 and Figure 4.18 show the Photon Transfer Curve (PTC) for NICMOS.

![Image of three frames from a movie mode](image)

Figure 4.16: One frame of the movie mode (left). The median of 100 frames of the movie mode (center). The median SD of 100 frames of the movie mode (right).

Figure 4.19 and Figure 4.20 show the readout rate for NICMOS in destructive and non-destructive mode, respectively, for spectral dispersion of 1 (non-dispersed), 3, and 5 channels per readout region. The camera readout is held at one Nloops while Nreads is increased to produce the desired camera read rates of 125, 250, 500, 750, or 1000 Hz.

Inside the NICMOS dewar, there is a filter wheel housing H, K’ (Figure 4.21), 2.101, and 2.36 μm filters as well as an Aluminum “cold” stop and an empty position.
Figure 4.17: Counts vs. SD, PTC for nicmos. The top left panel gives the counts vs. SD for nicmos. The black ‘+’ symbols denote each pixel, while the colored circles are the specific pixels read for data. The top right panel is a histogram of the SD. The bottom left panel is a histogram of the counts.
Figure 4.18: Counts vs. variance for the median of all 100 frames on NICMOS. Note: the variance is plotted here at integer levels.

Figure 4.19: NICMOS destructive mode read rate vs. Nreads for 1, 3, and 5 spectral channels.
Figure 4.20: NICMOS non-destructive mode read rate vs. Nreads for 1, 3, and 5 spectral channels.

Figure 4.21: The NICMOS K' filter transmission curve. (image courtesy: Rafael Millan-Gabet)
4.7 Motion Control

The most immediately apparent improvements to the FLUOR optical bench have been mechanical. FLUOR previously consisted of fold mirrors, a piezo-stack dither mirror to scan fringes, the OAPs to inject the fibers, and stepper motors to direct light to the OAPs. Alignments were done in the lab by hand and with an alignment telescope before the night's observations began. Now the addition of a Newport XPS Motion Controller/Driver has enabled the use of multiple mechanized stages to offer various configurations of the optical bench.

The Newport XPS Controller is capable of driving 8 axes of motion. We have connected 7 stages to the XPS. There is a stage for moving the FTS beamsplitter into position (subsection 4.8.4), two OPD stages, two alignment stages, a stage to adjust the focus for the H-band pupil camera (subsection 4.8.2), and the OUTPUT. The OPD Scan and alignment stages are discussed in subsubsection 4.5.2.1. The OUTPUT stage of JouFLU is a motorized rotating circular platform. It has three configurations: an open position for normal observations, a red LED for retro-injecting light for alignment procedures, and a Zinc Selenium (ZnSe) biprism to produce spectrally dispersed fringes (subsection 4.8.3).

<table>
<thead>
<tr>
<th>Stage</th>
<th>Model</th>
<th>Payload</th>
<th>Purpose</th>
<th>Range</th>
<th>Positions</th>
</tr>
</thead>
<tbody>
<tr>
<td>OPD SCAN</td>
<td>XMS50</td>
<td>dihedral</td>
<td>Fringe scanning</td>
<td>50 mm</td>
<td>—</td>
</tr>
<tr>
<td>OPD STAT</td>
<td>UTS100PP</td>
<td>dihedral</td>
<td>OPD correction</td>
<td>100 mm</td>
<td>FTS, Out, In</td>
</tr>
<tr>
<td>OUT</td>
<td>URM100PP</td>
<td>Backlight, Prism</td>
<td>alignment, spectral dispersion</td>
<td>360°</td>
<td>Out, Prism, Backlight</td>
</tr>
<tr>
<td>ALIU A</td>
<td>UTS150PP</td>
<td>Mirror, Dichroic</td>
<td>pupil observation, alignment</td>
<td>150 mm</td>
<td>Out, Dich, Mirror</td>
</tr>
<tr>
<td>ALIU B</td>
<td>UTS150PP</td>
<td>Mirror, Dichroic</td>
<td>pupil observation, alignment</td>
<td>150 mm</td>
<td>Out, Dich, Mirror</td>
</tr>
<tr>
<td>ALIU L2</td>
<td>MFA-PPD</td>
<td>L2 lens</td>
<td>focus for IRcam</td>
<td>25 mm</td>
<td>—</td>
</tr>
<tr>
<td>FTS</td>
<td>UTS100PP</td>
<td>Mirror &amp; BS cube</td>
<td>FTS mode</td>
<td>100 mm</td>
<td>Out, In</td>
</tr>
</tbody>
</table>
4.8 Observation Modes

JouFLU has five different configurations: (1) a normal observing mode for recording data on-sky, (2) an alignment mode allowing light to be retro-injected through the system and the alignment of either beam checked, (3) a pupil imaging mode that enables the CHARA system pupil to be checked while on-sky, (4) a spectrally-dispersed fringes mode, and (5) a FTS mode.

4.8.1 Alignment Mode

Figure 4.22: JouFLU in the configuration used when aligning the fiber injection. One of the alignment stages is moved to the dichroic position and the output stage is moved to the LED position. The LED projects red light back through MONA and to the ALIU dichroic. The tip/tilt stage is adjusted to center the JouFLU LED source with the CHARA white light source. The alignment stage is then moved out and the process can be repeated with the other beam. Figure B.1 shows more detail of this section of the bench.

A major focus of JouFLU is to improve the quality and ease of optical alignment procedures so as to increase the observing efficiency of FLUOR. The addition of ALIU improves the
accuracy and repeatability of alignment for the beams. These are useful if the initial alignment of the instrument is poor, or if the observer is operating remotely. These new stages also add the ability to view both the image and pupil planes. Previously, FLUOR alignment procedures could only be performed during the day and required personnel to be in the lab. The addition of ALIU provides a method for reliable alignment adjustments to be performed during the night with little interruption of data collection, or to be performed remotely. ALIU consists of two long-travel Newport stages, one for each beam, that carry a payload of a mirror and a dichroic. These stages are placed at 45° to the beam path and have three set positions: Open, Dichroic, and Mirror. During science observing ALIU is clear of the beam path. When an alignment needs to be performed one of the stages can position the dichroic into the path of the beam that is to be aligned (see Figure 4.22). Then on the JouFLU OUTPUT stage a red Light Emitting Diode (LED) is turned on and rotated into position to be retro-injected through MONA. After leaving the fibers, the light from the LED hits the ALIU dichroic and is reflected to a corner cube. The corner cube directs the light through the dichroic and to another system consisting of a mirror and focusing lenses (see Figure B.1). The LED spot is imaged by a Prosilica GE 640 visible light camera. The position of this spot is compared with that of a green laser spot produced by CHARA. The Zaber stages are then adjusted to overlay the two spots and conjugate the system. The final possibility is the insertion of a mirror that directs light to the Indium Gallium Arsenide (InGaAs) detector, that allows the pupil plane to be viewed in the H-band.
Figure 4.23: JouFLU in the configuration used when aligning with the ALIU system and the visible camera (VISCAM). The ALIU stage for the desired beam is moved to the dichroic position, and the light from the external source or star passes through a focusing lens (L1) and another dichroic (D2) to reach a fold mirror (M2), which directs it to the alignment camera (left). To check the CHARA pupil for vignetting or other possible loss of flux, an ALIU stage is moved to the mirror position and light is passed through the same lens used by the VISCAM, reflects off of dichroic D2, and reaches a focusing lens mounted on a stage (L2). The CHARA pupil is then recorded with the infrared camera (right). (image credit: LESIA - CNRS/Observatoire de Paris)

4.8.2 Pupil Imaging Mode

Viewing the CHARA pupil is accomplished with a new camera operating in H-band (0.9 - 1.7 $\mu$m) that is situated next to the visible alignment camera. This camera is a commercially available electronically cooled 320x256 InGaAs detector. The use of a 1.319 $\mu$m metrology laser by the CHARA Array necessitated the inclusion of two notch filters to prevent possible damage to the pupil-plane camera. To utilize the pupil camera, either of the ALIU stages moves to the mirror position. The beam passes through a lens and is then redirected at a right angle with a dichroic. It passes through the two notch filters and a focusing lens. Figure 4.24 shows an image of the CHARA pupil. The global response function for this
system is shown in Figure 4.25.

Figure 4.24: An image of the CHARA pupil taken by Maxime Motisi while observing Altair. The lines crossing the face of the star result from the CHARA telescope secondary mirror support spider.
Figure 4.25: The global response function for the H-band pupil imaging camera. Also plotted are the response functions for a single notch filter, the camera response function, and the dichroic (D2) reflection function. The global or “effective” function is the product of the camera’s response, the two notch filters, and the D2 reflection.

4.8.3 Spectral Dispersion Mode

The inclusion of a low-dispersion ZnSe biprism on the OUTPUT stage of JouFLU permits the measurement of dispersed fringes across up to five channels of the four outputs. The NICMOS readout rate limits the maximum number of spectral channels to five. By increasing the number of sampled visibilities, the simultaneous independent measurement of fringes across these channels increases the statistical accuracy for sources bright enough not to be dominated by detector noise.

Spectral dispersion allows simultaneous observations of the stellar photospheres in the molecular bands, such as CO, H$_2$0, and the continuum. Spectral resolution channels give access to the slope of the spectra of the target. This allows conclusions to be drawn as to the nature of the emission of the object. For instance, the spectral slope of excess emission, indicative of an exozodiacal dust disk, can suggest either thermal emission or photon
scattering; constraints may be placed on the location of the emission; and it can give support for models of the properties of dust grains.

Since the spectral bandpass of each channel is reduced, the width of each fringe packet is wider. Corresponding to this, the length of the OPD Scan must be increased to encompass the entire fringe packet. The JouFLU K-band central wavelength is 2.152 microns with a bandwidth of 0.326 microns. In spectrally dispersed mode the maximum bandwidth in a channel is approximately 0.2 microns (see Table 4.6). This changes the coherence length from 14.2 to 23.2 microns. So the dither scan length must be increased by 63% during spectral dispersion mode.

The spectral dispersion mode is of particular importance when the science star and the calibrator star are of different spectral types. The measurement of visibilities in different wavelengths enables the removal of chromatic bias or bandwidth smearing. Figure 4.26 illustrates the effect of bandwidth smearing. Because of the dependence of visibility on wavelength, if the calibrator stars are of different spectral type it will lead to an increase in the uncertainty of the science target visibility measurement. For a one mas diameter UD model star, a 0.1 micron bandwidth leads to a 2.6% difference in visibility at long baselines. To reach the desired < 1% visibility precision, the spectral type of calibrators must be taken into account. One way to do this is measure the visibilities in spectrally dispersed mode.

Another use of spectral-dispersion mode is to obtain the differential phase. Normally with two-telescope interferometry, no information of the phase is available. However, by measuring across multiple wavelengths, the relative phase difference between wavelengths can be
obtained. In a similar manner, differential visibility becomes an accessible measurable quantity. This enables the collection of a large amount of data simultaneously, leading to high statistical precision. Also, as these are differential quantities, calibration becomes less critical to the final measurement precision.

As a feasibility study of this technique, spectrally dispersed fringes were first recorded in 2004 as part of FLUOR (Mérand et al. 2006). Since 2014, spectrally dispersed fringes are an operational capability of JouFLU. Table 4.6 shows the effective wavelength and bandwidth per channel with $R_{\text{mean}} = \lambda_{\text{mean}} / \Delta \lambda_{\text{mean}} \approx 15$, as measured by fitting model fringes to the recorded fringes in each channel with knowledge of the stage velocity.
Figure 4.27: A frame from the “movie” readout during spectrally-dispersed mode. Now instead of four single pixels, each region corresponds to the dispersed output from one of the bundle fibers.

Table 4.6: Pixels 1-5 correspond to interferometric channel S1 and pixels 6-10 correspond to interferometric channel S2. Spectrally dispersed fringes are fit for wavelength and bandwidth to model fringes for each scan, this is then averaged over all scans per pixel and the results are given here. Note: that the last pixels in each dispersed signal region, i.e., numbers 3, 4, and 10, have low flux and hence larger errors.

<table>
<thead>
<tr>
<th>pixel</th>
<th>mean wavelength $\bar{\lambda}$ ($\mu$m)</th>
<th>mean filter width $\bar{\delta\lambda}$ ($\mu$m)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>2.218 ± 0.013</td>
<td>0.074 ± 0.057</td>
</tr>
<tr>
<td>2</td>
<td>2.119 ± 0.084</td>
<td>0.092 ± 0.037</td>
</tr>
<tr>
<td>3</td>
<td>2.024 ± 0.510</td>
<td>0.192 ± 0.517</td>
</tr>
<tr>
<td>4</td>
<td>2.166 ± 0.419</td>
<td>0.340 ± 1.544</td>
</tr>
<tr>
<td>5</td>
<td>2.052 ± 0.048</td>
<td>0.137 ± 0.114</td>
</tr>
<tr>
<td>6</td>
<td>2.216 ± 0.032</td>
<td>0.081 ± 0.051</td>
</tr>
<tr>
<td>7</td>
<td>2.130 ± 0.027</td>
<td>0.110 ± 0.084</td>
</tr>
<tr>
<td>8</td>
<td>2.145 ± 0.087</td>
<td>0.127 ± 0.268</td>
</tr>
<tr>
<td>9</td>
<td>2.063 ± 0.062</td>
<td>0.090 ± 0.076</td>
</tr>
<tr>
<td>10</td>
<td>1.944 ± 0.933</td>
<td>0.202 ± 0.707</td>
</tr>
</tbody>
</table>

mean 2.108 ± 0.222  0.144 ± 0.345
Figure 4.28: Data recorded during spectrally dispersed lab fringes. All but the top three fringes are from one of the pixels in the S1 and S2 regions. The first (yellow) and third (red) fringes from the top are the sum of each pixel in the S1 and S2 regions, respectively. The largest fringe (second from the top and shown in cyan) is the difference signal taken from the summed S1 and S2 fringes.
4.8.4 The Fourier Transform Spectrograph

Figure 4.29: JouFLU in the configuration used when calibrating with the Fourier Transform Spectrograph. The alignment stages are moved out, the FTS beamsplitter is moved into place, and OPD Stat changes position to compensate for the change in OPD.

Michelson first applied the principles behind Fourier transform spectroscopy; he observed the effect on intensity and made Zeeman effect estimates, by eye. He also invented a specific kind of mechanical computer or “harmonic analyzer” to perform Fourier transforms for FTS. With this he resolved several spectra, including doublets. There have been multiple reviews of the state of Fourier transform spectroscopy, which is now a common laboratory technique, and this section is by no means exhaustive; the FTS mode of JouFLU is an on-going effort. Nevertheless some qualitative discussion of the features and benefits of FTS are mentioned. The reader is referred to the work by Connes (1970); Bell (1972); Ridgway & Brault (1984) for more depth to the subject.

Recalling from the introduction, the fringe envelope of the interferogram is the Fourier
transform of the bandpass. By taking the inverse Fourier transform of this, one obtains the spectrum of the source. Fourier transform spectroscopy has several unique advantages over grating or prism-based spectroscopy.

With an FTS the entire spectra, throughout the passband, is measured simultaneously. This is termed the multiplex, or Fellgett, advantage. This arises from the fact that the FTS is rapidly scanning in Fourier space. An added benefit from this is that any variation in intensity affects all wavelengths equally which allows for uniform frequency calibration for the entire bandpass. Spectra from a FTS are self-normalized; that is changes in photon flux affect the entire spectrum simultaneously. The FTS does not suffer from the problem of scattered light contaminating the spectrum, which plagued many spectrographs. These features increase the photometric accuracy possible with a FTS.

The second major advantage in using an FTS is the Jacquinot advantage: a FTS typically has a large, circular entrance aperture as opposed to the narrow slit of a dispersive spectrograph, this gives the FTS an higher throughput, or étendue.

All of this leads to an instrument capable of very high Signal-to-Noise Ratio (SNR) with a wide spectral resolution range that can be varied continuously, and does not suffer distortion from the instrument.

The resolution of a FTS varies with the maximum OPD,

$$\delta\sigma = \frac{1}{2L} \Rightarrow R = \frac{2L}{\lambda},$$  \hspace{1cm} (4.7)
where $L$ is the path length range, $\Delta x$, in the general fringe equation (Equation 1.29). This corresponds to the first zero crossing of the sinc function in the intensity distribution.

**OPD Scan** has a travel range of 50 mm, and after taking into account the double pass of the dihedral mirror setup on the stage, this puts the maximum $R = 18000$. However, this is an overly optimistic estimate that doesn’t take into account the stability and tolerances of the scanning stage. With a more modest range, which is more likely to reach the strict tolerances required and has already been tested for velocity stability, this value becomes $R = 2000$.

However, the size of the source can be the limiting factor to the resolution. The interferogram decreases as the optical path moves from the 0 OPD position, there is an OPD where the interferogram SNR falls below unity. Recording the interferogram for large OPD only increases the noise in the spectrum and does not improve the resolution. This limit, plus the presence of false sidelobes in the sinc function, necessitates apodization. False sidelobes are introduced because optical path displacements are finite (Bell 1972). Apodization is performed by multiplying the interferogram by a function to remove these false sidelobes. Apodization broadens lines, reducing the spectral resolution. Thus, the resolution of a FTS is ultimately limited by the strength of the signal.

The retrieval of both spatial and spectral information about a source is double Fourier interferometry. The addition of a FTS mode to JouFLU provides wavelength calibration of the science camera. In FTS mode, a single beam is taken from CHARA. The single beam does not suffer from differential seeing and atmospheric piston. This allows longer scans at 100 Hz. The single beam is passed through a cubic beamsplitter and a fold mirror mounted
on a motorized stage. The beamsplitter generates the two beams necessary to feed into MONA for combination (See Figure 4.29). The beam splitter is made of CaF$_2$ and is non-polarized. Transmission for the cube is 41\% for one face and 44\% for the other at 2.2 \( \mu \)m. A mirror is also mounted on the FTS stage to direct the reflected portion of beam A into JouFLU beam B. OPD Stat adjusts the OPD to account for the additional path introduced to beam B. OPD Scan modulates the OPD. Because the source is unresolved by a single telescope, the resulting interferogram is affected only by the spectrum instead of spatial structure of the source. This spectrum can then be used to provide photometric and spectral calibrations of the instrument; to retrieve IR spectra of stars; and potentially provide information on the spectral slope of stars that display excess NIR emission. This mode has not yet been tested on-sky.

4.9 Remote Operation

Network control of the XPS along with integration of the XPS, NICMOS, and other subsystems to the CHARA software environment provides the ability to correct alignments or reconfigure the instrument while on sky without going into the lab. It also allows the possibility for completely remote operations from any of the CHARA remote observing control rooms located in Atlanta, Meudon, Nice, Sydney, Ann Arbor, and Bonn. These remote facilities are equipped with computer workstations whose hard-drives have been imaged with clones of the CHARA operating software. The stations are directly connected to the CHARA Array via a VPN. It is thus possible, from any of the facilities, to control any aspect of running the Array.
In fact, remote operations quite frequently only require an on-site operator for nightly start-up, weather monitoring, safety/security of the instruments, and shut-down.

### 4.10 Software

As part of the JouFLU integration, it was decided to make the beam combiner compliant with the CHARA software environment. CHARA Technical Report No. 70 (ten Brummelaar 1998) describes the standard for coding practices for the Array. All programs are to be written in ANSI C, C++, or Bourne shell scripting language, functions must be prototyped, and compiled with gcc using the `-Wall` flag. There are standard libraries provided to facilitate uniform code.

The software changes for JouFLU occurred in two stages:

1. The original FLUOR algorithms (Coudé du Foresto et al. 1997) were ported to a modern C compiler. At this stage, nothing about FLUOR had changed apart from the environment in which the software was running. The hardware was tested without additional change.

2. The camera readout was converted from serial (RS-232) readout to ethernet readout via Category 5 (cat 5) cable and socket communications. This greatly improved the data bandwidth, enabling higher camera read rates and more data channels for spectral dispersion mode.

Beyond this were numerous additions to the server software, including the ability to communicate with the new hardware, XPS, fiber injection tip/tilt stages (Zabers), and CHARA
subsystems. For ease of use, an efficient Graphical User Interface was added to control the server. Finally, JouFLU was incorporated into the CHARA observation sequencer, “Cosmic Debris”, that controls the numerous steps that are required, from moving to a target to finding fringes and recording data.

4.10.1 The Move to C

The previous Command/Control system for FLUOR consisted of software originally written for use while FLUOR was at IOTA. The software was written in an early version of LabView and operated parallel to the normal functions of CHARA. As such it was unable to take advantage of some of the features present at the CHARA Array due to the integrated software environment. To remedy this, the FLUOR software system was rewritten as the JouFLU server and Graphical User Interface (GUI). This required an extensive ground-up conversion, by Theo ten Brummelaar, of the proprietary LabView code to C code modeled on existing CHARA beam combiner functions. After this conversion, new features were added, by Nicholas Scott and Theo ten Brummelaar, to the code to provide for the new hardware capabilities of JouFLU. With the new software JouFLU can be integrated with other CHARA systems such as the fringe tracker, CHAMP. The major software changes occurred over a two year period, from mid-2011 until mid-2013, with improvements and development continuing indefinitely.

There are several major advantages to the new software system. The JouFLU server runs on a modern rack-mounted PC running the CentOS distribution of the Linux operating system.
Operations from any CHARA remote facility are now possible. JouFLU will be kept up to date with any system software updates to the CHARA operating environment. The JouFLU Data Reduction Software (DRS) began as a modified version of the very well understood CLASSIC CHARA DRS. Now, the JouFLU DRS is implementing the newest data reduction strategies and the two combiners can share a common DRS. Theo ten Brummelaar, with the assistance of Paul Nuñez, has spent months creating this highly sophisticated DRS.

4.10.2 New Camera Code

The NICMOS camera is controlled by single computer running MS-DOS that is on an isolated ethernet network with the JouFLU computer. Updates have been made to the software that runs NICMOS to enable the replacement of the RS-232 serial port readout to an ethernet network-based readout. This increased the available bandwidth and allows up to 20 pixels total (5 per channel) to be read-out at 500 Hz.

4.10.3 The JouFLU Server

The software at the CHARA array has been developed by Theo ten Brummelaar (ten Brummelaar 1998) and relies on socket-based communication. There is a server, typically running some aspect of hardware, and a client. The client is typically a Graphical User Interface (GUI). One and only one instance of a server is allowed to run. Servers are typically run on a specific computer and can connect to multiple instances of a client. The server has one or more sockets that are bound to specific ports. It listens for connection to a port from a
client. Once the connection is made, the client can communicate instructions to the hardware through the server. This is the interface between the user and the instrument hardware. A full description of the JouFLU server is located in section A.1.

Figure 4.30: The JouFLU server gives information on the XPS stage positions, the NICMOS camera readout mode and rate, and offers direct control over all JouFLU systems.

4.10.4 The Graphical User Interface

A GUI is presented to the observer that provides identical functionality as the server but in a more intuitive format. The JouFLU GUI has tabs for the data collection setup, the camera setup, the XPS stage positions and control, viewing the H-band pupil camera, the alignment process and Zaber control, movie or camera frame readout mode and camera alignment, photometry monitoring, data collection, and CHARA configuration. More detailed information for each tab and all functions of the GUI are given in section A.2.
Figure 4.31: The original **FLUOR GUI** from **IOTA**, operating under LabView. (image credit: Vincent Coudé du Foresto)

Figure 4.32: The **JouFLU GUI** (*bottom*) shown while recording white light lab fringes. Clockwise from top left: the server, the fringes, photometry per channel, fringe power spectrum, a waterfall plot of the fringes, and a waterfall plot of the power spectrum.
4.11 Polarization and other issues

Early tests both in the lab and on sky revealed a lower than expected instrument visibility from JouFLU. These low visibilities raised concern that the combiner may be exhibiting either differential polarization rotation, differential polarization delay, or a combination of both.

Figure 4.33: The effects of adjustment of the fiber polarization for each of the two beams on white light fringe visibility. Each fiber has a knob which affects the bend of the fiber. Each knob was incrementally adjusted in steps of 60 (arbitrary units) to give 100 data points over the full range of travel. Based on these results MONA was set to the maximum V (480 and 180 for the top and bottom knobs, respectively).

As part of the JouFLU project, the fiber combiner box MONA was sent to LVF for re-calibration and adjustment. The fiber heads were cleaned and two knobs with numbered scales were added to provide precise adjustment of the polarization rotation (“Mickey ears”) of the fibers. Once installed back at the CHARA Array, the interferometric signal throughput was maximized. This was achieved by scanning through the range of the polarization
adjustment for each beam while measuring the visibility of lab fringes generated by a controlled WL (see Figure 4.33). However, even after this optimization process, the observed visibilities were lower than expected.

To determine the cause of the low visibility, a battery of tests was conducted using two linear K-band polarizers on rotating mounts. Table 4.7 describes the various test conditions.

Universally for all tests the following guidelines were followed: one data file was recorded for each move of rotated polarizer; one polarizer was always held stationary at either 0° or 90°; and the other polarizer was rotated in steps of 10° from 0° to 180°.

Table 4.7: *should have fringes, others are tests of photometry. Test 0 is to check the polarizers with respect to each other to determine their polarization axis. Polarizers were either placed at the CHARA WL source, the beam inputs at the edge of the JouFLU optical table, or at the OUTPUT stage immediately before the camera aperture.

<table>
<thead>
<tr>
<th>Name</th>
<th>Polarizer 1 position</th>
<th>Polarizer 1 rotation (°)</th>
<th>Polarizer 2 position</th>
<th>Polarizer 2 rotation (°)</th>
<th>Comments</th>
</tr>
</thead>
<tbody>
<tr>
<td>baseline</td>
<td>—</td>
<td>—</td>
<td>—</td>
<td>—</td>
<td>no polarizers</td>
</tr>
<tr>
<td>Test 0-b</td>
<td>b input</td>
<td>0</td>
<td>b input</td>
<td>rotated</td>
<td>beam a blocked</td>
</tr>
<tr>
<td>Test 0-a</td>
<td>a input</td>
<td>0</td>
<td>a input</td>
<td>rotated</td>
<td>beam b blocked</td>
</tr>
<tr>
<td>Test 1b-0</td>
<td>output</td>
<td>0</td>
<td>b input</td>
<td>rotated</td>
<td>beam a blocked</td>
</tr>
<tr>
<td>Test 1a-0</td>
<td>output</td>
<td>0</td>
<td>a input</td>
<td>rotated</td>
<td>beam b blocked</td>
</tr>
<tr>
<td>Test 1b-90</td>
<td>output</td>
<td>90</td>
<td>b input</td>
<td>rotated</td>
<td>beam a blocked</td>
</tr>
<tr>
<td>Test 1a-90</td>
<td>output</td>
<td>90</td>
<td>a input</td>
<td>rotated</td>
<td>beam b blocked</td>
</tr>
<tr>
<td>*Test 2b-0</td>
<td>b input</td>
<td>rotated</td>
<td>a input</td>
<td>0</td>
<td></td>
</tr>
<tr>
<td>*Test 2a-0</td>
<td>a input</td>
<td>rotated</td>
<td>b input</td>
<td>0</td>
<td></td>
</tr>
<tr>
<td>*Test 2b-90</td>
<td>b input</td>
<td>rotated</td>
<td>a input</td>
<td>90</td>
<td></td>
</tr>
<tr>
<td>*Test 2a-90</td>
<td>a input</td>
<td>rotated</td>
<td>b input</td>
<td>90</td>
<td></td>
</tr>
<tr>
<td>*Test 3-0</td>
<td>WL</td>
<td>rotated</td>
<td>output</td>
<td>0</td>
<td></td>
</tr>
<tr>
<td>*Test 3-90</td>
<td>WL</td>
<td>rotated</td>
<td>output</td>
<td>90</td>
<td></td>
</tr>
</tbody>
</table>

For Test 1, a single polarizer was placed in either beam A or beam B, and the other beam was blocked; another polarizer was placed at the combiner output, before the detector. The photometry of each beam was recorded as the input polarizer was rotated (See Figure 4.34).

As expected with two linear polarizers in-line with each other, the photon flux at the output
varies sinusoidally with the angle of rotation of the polarizer. However, this sinusoidal pattern does not match between the two beams and a difference in phase is apparent. This is indicative of the combiner displaying uncorrected differential polarization rotation. By tracking the shift in phase, the amount of differential polarization rotation can be determined (See Table 4.8).

The amount of differential polarization rotation can be found by looking at Figure 4.34 and tracing the polarization angle where the signal peak lies,

For $0^\circ$ output polarization:

$$A_1^0 - B_1^0 = 100^\circ - 40^\circ = 50^\circ$$

$$A_2^0 - B_2^0 = 180^\circ - 100^\circ = 80^\circ$$

For $90^\circ$ output polarization:

$$A_{190} - B_{190} = 160^\circ - 125^\circ = 35^\circ$$

$$A_{290} - B_{290} = 100^\circ - 170^\circ = -70^\circ = 110^\circ$$

So for the difference between the two beams in each polarization:

$$\left( A_2^0 - B_2^0 \right) - \left( A_1^0 - B_1^0 \right) = 80^\circ - 50^\circ = 30^\circ$$

$$\left( A_{290} - B_{290} \right) - \left( A_{190} - B_{190} \right) = 110^\circ - 35^\circ = 75^\circ$$

<table>
<thead>
<tr>
<th>Name</th>
<th>Signal 1 ($^\circ$)</th>
<th>Signal 2 ($^\circ$)</th>
<th>Difference ($^\circ$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Test1-0</td>
<td>50</td>
<td>80</td>
<td>30</td>
</tr>
<tr>
<td>Test1-90</td>
<td>35</td>
<td>110</td>
<td>75</td>
</tr>
</tbody>
</table>
Figure 4.34: Photometry of the JouFLU interferometric signal channels for tests 1b-0, 1a-0, 1b-90, 1a-90. As expected when two polarizers are placed in front of each other and one is rotated, there is a sinusoidal variation in the amount of flux transmitted. The evidence for differential polarization rotation comes from the difference in phase of these variations. The left two plots are for the 0° output polarization, while the right two plots are for the 90° output polarization.

Test 2 consisted of placing one polarizer at the input of each beam and rotating one of the polarizers, while holding the other stationary. This was performed on both the CLASSIC CHARA beam combiner and the JouFLU combiner. The results are shown in Figure 4.35 and Figure 4.36, respectively. Here it is apparent that there are complex polarization effects affecting JouFLU and resulting in a disturbance to the expected fringe contrast.

The final experiment, Test 3, placed one polarizer at WL source and one at the combiner output in front of the detector. This polarizer was then rotated. Figure 4.39 and Figure 4.40 compare the results of this test on both JouFLU and CLASSIC.

The results of these tests, in particular Figure 4.40, match the model condition given by Equation 4.19 below where $d\phi = 109^\circ$ with a 90° periodicity (Figure 4.41). This periodicity
Figure 4.35: Visibility measurements with CLASSIC for tests 2b-0, 2a-0, 2b-90, and 2a-90. These visibility plots are as expected for the case of superposition of two linearly polarized beams.

Figure 4.36: Visibility measurements with JouFLU for tests 2b-0, 2a-0, 2b-90, and 2a-90. In this case it is apparent that there are more complex polarization effects present than the expected superposition of two linearly polarized beams.

suggests differential phase delay is occurring inside the beam combiner. Differential phase delay, along with polarization rotation, finite fringe and temporal sampling, bandwidth smearing, and beam imbalance, results in a maximum visibility of $V_{\text{max}} \approx 0.73$. 
Figure 4.37: CLASSIC test 3 photometry shows no significant variation in flux.

Figure 4.38: JouFLU test 3 photometry shows no significant variation in flux compared to CLASSIC. This rules out variation in flux as a cause of the modulation of visibility seen during test 3.

This is our case, test 3, of having one polarizer at the WL source and one at the camera. So we have a two linearly polarized beams entering MONA, there they are superimposed and gain some polarization state, then pass through the output polarizer and are again linearly polarized. We essentially have the case of circularly polarized light, superimposed and then passing through a linear polarizer.
If we assume the source is circularly polarized and we measure two orthogonal polarizations in the $x$ and $y$ planes and start with the two input beams:

$$
E_1 = \cos \theta \hat{x} + \sin \theta \hat{y}
$$

$$
E_2 = \cos \phi e^{i\phi} \hat{x} + \sin \theta e^{i(\phi + \Delta \phi)} \hat{y}.
$$

(4.11)

where $\phi$ is the differential phase of polarization $x$ between the two beams, $\phi + \Delta \phi$ is the differential phase for polarization $y$, $\Delta \phi$ is the maximum phase delay between the two polarizations, and $\theta$ is the angle of a linear polarizer inserted into the beam.
The superposition of the polarized waves gives the total intensity, \( I \), as

\[
I = |E_1 + E_2|^2. \tag{4.12}
\]

Substituting Equation 4.11 into Equation 4.12 gives:

\[
I = \left| \cos \theta \left( 1 + e^{i\phi} \hat{x} + (\sin \theta \left( 1 + e^{i(\phi+\Delta\phi)} \hat{y} \right) \right)^2
\]

\[
= \cos^2 \theta \left| 1 + e^{i\phi} \right|^2 + \sin^2 \theta \left| e^{i(\phi+\Delta\phi)} \right|^2
\]

\[
= \cos^2 \theta \left( 1 + e^{i\phi} \right) \left( 1 + e^{-i\phi} \right) + \sin^2 \theta \left( 1 + e^{i(\phi+\Delta\phi)} \right) \left( 1 + e^{-i(\phi+\Delta\phi)} \right). \tag{4.13}
\]

From the relationship between sine, cosine, and exponential forms: \( e^{\pm\phi} = \cos \phi \pm i \sin \phi \), this becomes,

\[
I = \cos^2 \theta \left( 2 + \cos \phi - i \sin \phi + \cos \phi + i \sin \phi \right) +
\]

\[
\sin^2 \theta \left( 2 + \cos (\phi + \Delta\phi) - i \sin (\phi + \Delta\phi) + \cos (\phi + \Delta\phi) + i \sin (\phi + \Delta\phi) \right). \tag{4.14}
\]

The imaginary terms cancel, leaving,

\[
I = \cos^2 \theta (2 + 2 \cos \phi) + \sin^2 \theta (2 + 2 \cos (\phi + \Delta\phi)). \tag{4.15}
\]

This can be further simplified and making use of the identity: \( \cos^2 \phi + \sin^2 \phi = 1 \), takes the form

\[
I = 2 \left[ 1 + \cos^2 \theta \cos \phi + \sin^2 \theta \cos (\phi + \Delta\phi) \right]. \tag{4.16}
\]

Using the identity, \( \cos (\phi + \Delta\phi) = \cos \phi \cos (\Delta\phi) - \sin \phi \sin (\Delta\phi) \), this can be written as,

\[
I = 2 \left[ 1 + \cos^2 \theta \cos \phi + \sin^2 \theta (\cos \phi \cos (\Delta\phi) - \sin \phi \sin (\Delta\phi)) \right]
\]

\[
= 2 \left[ 1 + a \cos \phi - b \sin \phi \right]. \tag{4.17}
\]
where \( a = \cos^2 \theta + \cos (\Delta \phi) \sin^2 \theta \) and \( b = \sin (\Delta \phi) \sin^2 \theta \). Now we have \( I \) in terms of the phase, \( \phi \), and it can be expressed as,

\[
I = 2 \left( 1 + \sqrt{a^2 + b^2} \left[ \frac{a}{\sqrt{a^2 + b^2}} \cos \phi - \frac{b}{\sqrt{a^2 + b^2}} \sin \phi \right] \right).
\]

(4.18)

Knowing we want \( I \) of the form \( I = 2(1 + V \cos (\phi + \phi_0)) \) means that we can state \( V = \sqrt{a^2 + b^2} \) and differential delay with respect to polarization gives

\[
I = 2 \left[ 1 + \cos \phi \cos^2 \theta + \cos (\phi + \Delta \phi) \sin^2 \theta \right].
\]

(4.19)

This is modeled in Figure 4.41 and matches the empirical results shown in Figure 4.40. We find that there is a differential delay in phase between orthogonal polarizations equal to 109°.

![Figure 4.41: Results of model to determine the amount of polarization phase delay.](image)

As a result of these series of tests, we have found evidence of differential polarization rotation and differential phase delay. Differential polarization rotation is analogous to the visibility loss
resulting from beam intensity mismatch with the following effects:

\[ V_{obs} = V \frac{2 \cos(\alpha)}{1 + \cos^2(\alpha)} \]  \hspace{1cm} (4.20)

where for

\[ 50^\circ \rightarrow V_{max} \approx 90\% \]

and

\[ 70^\circ \rightarrow V_{max} \approx 60\%. \]

Additionally, we noticed that CHARA laboratory white-light source (WL) is probably not circularly polarized. We observe an average \( V \approx 0.42 \) instead of the expected 0.59. So, we may conclude that the WL is elliptically polarized (close to 50° or 150°).

External polarization rotation controllers that were used on a previous beam combiner have been sent to CHARA and, if necessary, can be added to the JouFLU optical bench. This would provide an extended range of polarization rotation control when used with the full range of the Mickey ears. It remains to be seen if the gain from matching polarization rotation would compensate for the sensitivity that would be lost by this method.

To correct the polarization phase delay, lithium niobate plates are planned for incorporation in the beams prior to reaching the JouFLU optical table. These plates would be AR coated for K band and mounted on rotation stages with the faces perpendicular to the beam axis. Differential rotation of the plates would induce a delay in the polarization phase of that beam by increasing the path for that beam polarization. A similar method to control polarization
phase delay was introduced to PIONIER, a 4-beam H band beam combiner at the VLTI (Lazareff et al. 2012). This technique enabled them to reach instrumental contrasts of \( \approx 98.5\% \) with a throughput loss of only \( \approx 10\% \).

### 4.12 Current Status

JouFLU first obtained fringes on the sky in May 2012 and was on-line for regular operation during 2013. Observing runs were carried out during the 2013, 2014, and 2015 observing seasons. A complete realignment of the JouFLU bench has been completed as well as the development of new alignment procedures. Fringes have been obtained on objects as faint as FU Ori (\( K_{\text{mag}} = 5.16 \)). However these data were of questionable quality. The best results have come from calibrated data points restricted to K magnitudes less than 4 to 4.5. The mean error in V of a single data point for JouFLU \( \approx 1\% \), (compared to \( \approx 5\% \) for CLASSIC). For multiple bracketed points over several nights, this error can be significantly reduced.

#### 4.12.1 Magnitude Limit Calculation

A functional approximation of the limiting magnitude for the instrument can be determined on sky that omits any assumptions about internal reflections.

The flux from Vega scaled for the apparent magnitude of the target star is given by

\[
F_\nu = f_\nu \times 10^{\frac{-m_K}{2.5}},
\]

(4.21)
where \( f_\nu \) is the Vega flux value, 640 Jy, for the given filter (Bessell et al. 1998), \( m_K \) is the apparent magnitude of the target star in K, and \( F_\nu \) is the Vega corrected flux.

The energy of a photon is, of course, given by

\[
E = h\nu = \frac{hc}{\lambda},
\]

where \( h = 6.63 \times 10^{-34} \text{Js} \).

The spectral bandpass is defined by

\[
\Delta \nu = \nu_1 - \nu_2 = \frac{c}{\lambda_1} - \frac{c}{\lambda_2} \approx \frac{c\Delta \lambda}{\lambda_0^2},
\]

which involves the relation

\[
\frac{\Delta \nu}{E} = \frac{\Delta \lambda}{h\lambda_0} = \frac{1}{hR}.
\]

The bandpass is narrow so we can assume for these purposes that all photons share the same energy.

R in Equation 4.24 is the spectral resolving power,

\[
R = \frac{\lambda_0}{\Delta \lambda}.
\]

This gives the flux in number of photons per second per square meter for a given filter and integration time, \( t_{\text{int}} \), as

\[
n_{\text{incident}} = \frac{F_\nu \cdot 10^{-26} \cdot t_{\text{int}}}{Rh}.
\]
The overall efficiency in terms of number of photons entering the telescope of area, $A_{\text{tel}}$, compared to the number of photons detected is

$$\text{eff} = \frac{n_{\text{detector}} \cdot A_{\text{tel}}}{n_{\text{incident}}}, \tag{4.27}$$

and the photon flux at detector is

$$n_{\text{counts}} = \frac{n_{\text{detector}} \cdot \text{gain}}{\text{gain}}. \tag{4.28}$$

To get the flux in number of photons per second per square meter for a filter in Jansky units, one would divide the scaled flux of the object by the energy of a single photon and multiply this by the spectral bandpass, $\Delta \nu$.

Then taking the number of counts detected at the camera, one can calculate the overall efficiency for the system and the minimum number of counts necessary to get fringes. Knowing this, it is possible to solve for limiting magnitude.

Rearranging Equation 4.21 and Equation 4.26 gives:

$$m_{\text{k,lim}} = -2.5 \log \frac{F_{\nu}}{f_{\nu}} \tag{4.29}$$

and

$$F_{\nu} = \frac{n_{\text{incident}} \cdot R \cdot h}{10^{-26} \cdot t_{\text{int}}}, \tag{4.30}$$

respectively.
Finally, rearranging Equation 4.27 gives

$$n_{\text{incident}} = \frac{n_{\text{counts}} \cdot \text{gain} \cdot A_{\text{tel}}}{\text{eff}}.$$  \hspace{1cm} (4.31)

Putting this all together in Equation 4.29 and setting $n_{\text{counts}}$ to the minimum number of counts on the detector where decent fringes are detected gives

$$m_{\text{k,lim}} = -2.5 \log \frac{n_{\text{counts}} \cdot \text{gain} \cdot A_{\text{tel}} \cdot R \cdot h}{\text{eff} \cdot 10^{-26} \cdot t_{\text{int}} \cdot f_{\nu}}.$$  \hspace{1cm} (4.32)

Here, the efficiency and $n_{\text{counts}}$ are found empirically and used along with the known gain to predict the limiting magnitude. This assumes that the throughput is a constant.

In practice, 30 counts at 500 Hz is a threshold for good data quality with FLUOR, and the typically observed efficiency based on the number of counts at the detector for a given magnitude star is 0.22\% in average seeing. Using this gives a total system throughput and a limiting magnitude of $K_{\text{mag}} = 4.54$.

4.12.2 Comparison of JouFLU with FLUOR and CLASSIC

HD 34411 is a G1V star with a $K_{\text{mag}} = 3.24$. It has no detected infrared excess. In order to characterize part of the performance of JouFLU on-sky, this star was selected to compare with well-established results from the CLASSIC beam combiner. The diameter of HD 34411 from Boyajian et al. (2012) was compared with that found with JouFLU. Five data points from a single baseline, E2/W2, were taken with JouFLU in just 2.5 hours on January 16, 2014. From these data, we derive a uniform diameter of $0.991 \pm 0.028$ mas for HD 34411 with a
$\chi^2 = 1.09$ (Figure 4.42 and Figure 4.43).

Figure 4.42: The visibility of HD 34411 from JouFLU.

Figure 4.43: The visibility of HD 34411 from JouFLU with the V plot range constrained.
Boyajian et al. (2012) used 18 points from 4 nights and 2 baselines (S1/E1 and W1/E1) to fit their diameter for HD 34411 (Figure 4.44). From this they obtained a uniform disk diameter of $0.958 \pm 0.015$ mas and a reduced $\chi^2 = 1.07$.

Our results agree within less than 1.5 $\sigma$, despite JouFLU having many fewer data points and using a much shorter baseline. From this it is apparent that JouFLU is operating well and can deliver high-precision visibility measures.

4.13 Future

Looking ahead, there are plans to implement further upgrades to add capability to JouFLU. More data will be collected utilizing the spectral dispersion mode, and the FTS will be used on
sky to calibrate the camera. Refinements to the camera readout software will add more flexibility to reading the camera as well as on-the-fly changes of the ROI.

Future improvements will include modification to work with the fringe tracker, CHAMP (Berger et al. 2008). It should also be possible to modify JouFLU to use the CHARA CLASSIC beam combiner as a fringe tracker. Fringe-tracking will enable much higher precision data to be obtained with JouFLU by reducing or removing the remaining atmospheric piston error. This will allow longer tracking on fringes and a higher quantity of shorter scans, increasing the data throughput and statistical precision. This will lead to higher accuracy measurements of raw visibility. Fringe tracking will also improve observation during periods of high atmospheric turbulence. This will increase the observing efficiency of the instrument.

Work has begun at the CHARA Array to bring adaptive optics to the six telescopes of the interferometer (Che et al. 2013). JouFLU will be among the five other beam combiners at the CHARA Array to benefit from this.

Currently, JouFLU performs poorly if \( r_0 < 5 \text{ cm} \) in V band due to the severe drop in fiber injection efficiency. Figure 4.45 shows the expected effect of close-loop AO on the Strehl at 1.85\( \mu \text{m} \). The addition of AO will allow JouFLU to work at its current magnitude limit during all but the worst seeing conditions.

Strehl ratio is functionally defined as the ratio of the maximum measured peak intensity Point Spread Function (PSF) to the maximum intensity for a diffraction-limited PSF,

\[
S = \frac{I}{I_0},
\]  
(4.33)
where $I_0$ is the maximum of the diffraction-limited PSF. Strehl can be more formally defined as,

$$S = \frac{1}{\pi^2} \left| \int_0^1 \int_0^{2\pi} \rho \, d\rho \, d\theta \, e^{i\psi} \right|^2,$$

for a circular aperture where $\psi$ is a function which represents the wavefront distortion and $S$ ranges from 0 to 1. The relation of Strehl ratio to the wavefront errors (Maréchal approximation) is (Tatarski 1961; Fried 1965; ten Brummelaar et al. 1995; Lawson 2000)

$$S = e^{-\sigma_\phi^2}$$

where $\sigma_\phi^2$ is defined in Equation 3.13 and
\[ r_0 \propto \lambda^{\frac{6}{5}}. \] (4.36)

If \( r_0 = 5 \) cm in V band, then \( r_0 \approx 26 \) cm in K band. For a 1-m aperture telescope, this and Equation 4.35 gives a Strehl ratio of \( 6 \times 10^{-5} \). An AO system with 50% Strehl would hence be a greater than 8000-fold improvement.

The median \( r_0 \) for CHARA is 11 cm in V band during the peak observing season. This is equivalent to 58 cm in K band. This gives a Strehl of 0.08. The AO system would then yield a 6-fold improvement in flux. This would push the limiting magnitude of JouFLU from \( \approx 4.5 \) to \( \approx 6.5 \). For the exozodi survey this would dramatically increase the number of potential targets from \( \approx 300 \) to \( \approx 3,000 \).

From a purely technical point of view, AO will bring the following improvements.

- Visible \( r_0 = 5 \) to 10 cm providing a K-magnitude limit gain of 1 to 2.5 depending on actual seeing conditions and a gain of a factor of 2 or more in observation efficiency and/or accuracy.
- Visible \( r_0 = 2 \) to 5 cm achieving a gain of 2.5 to 4 K-magnitudes, often making the difference between staying closed or being able to observe at all.

FLUOR was a productive, high-precision instrument that fulfilled a niche for high-precision visibility measurements at the CHARA Array. The upgrades described herein continue that role while adding much needed improvements. JouFLU adds higher efficiency, increased
data throughput, new science capabilities, greater ease of use, and more user accessibility to this established instrument. The addition of an FTS mode and the option of collecting spectrally-dispersed fringes offer new possibilities for science with JouFLU. Better integration with the CHARA Array allows remote operation as well as rolling updates to the software control systems and data reduction software. Better alignment and AO leading to higher throughput and faster collection of high quality data scans should allow even higher precision visibilities for JouFLU.
“A man said to the universe:

“Sir, I exist!”

“However,” replied the universe,

“The fact has not created in me

A sense of obligation.”

— Stephen Crane
Part III

Science
Dust is ubiquitous. It permeates our own Solar system and exists throughout other stellar systems. Rings of cold dust dwell far from the Sun in the Kuiper belt. Closer to our star, dust surrounds the Earth and extends to within a few radii of the Sun. Sunlight scattering from this dust disk creates the zodiacal light that is visible across the ecliptic after sunset. The disk formed by zodiacal dust has an immense surface area that, when viewed from a distance, makes it the brightest object in the system after the Sun.

The Solar System’s zodiacal disk is composed of small, \(< 100 \ \mu m\), particles deposited by comet infall and asteroidal collisions. This dust is short lived and continually replenished by spontaneous disruptions of short-period comets. Particles ejected from comets undergo a collisional cascade, eventually reaching sub-\(\mu\)m sizes, where they are strongly influenced by Poynting-Robertson (PR) drag, radiation pressure, and sublimation while drifting in to \(< 1 \text{ AU}\) of the Sun. These factors are discussed in section 5.5. During the Late Heavy Bombardment (LHB), \(\approx 3.8\) Gyr ago, this inner region of the zodiacal cloud could have been
more than $10^4$ times brighter than today (Nesvorný et al. 2010).

Stellar light is absorbed by dust grains and reemitted at IR wavelengths. Exozodiacal disks are measured in terms of “zodi” or fractional dust luminosity, defined as

$$L_{\text{dust}} / L_\odot \sim 2 \times 10^{-7},$$  \hspace{1cm} (5.1)

where is the Sun’s luminosity, $L_\odot = 3.839 \times 10^{26} W$ (Nesvorný et al. 2010). In the optically thin case, it is proportional to the dust mass, but is strongly affected by dust grain size and composition. A 1% NIR excess in a Sun-like zodiacal distribution is $> 1000$ zodis.

Spectroscopic dust detections are frequently referred to as “excesses”. The term refers to higher than expected emission in the Spectral Energy Distribution (SED) of the object in question. The implication is that this extra emission, over that expected from blackbody or stellar models, is the result of the presence of additional blackbody radiation in the form of a reservoir of small grains at lower than stellar temperature.

In 1984, the The Infrared Astronomical Satellite (IRAS) mission detected the first large infrared excess from a main sequence star that wasn’t undergoing significant mass loss (Aumann et al. 1984). Subsequent observations with observatories such as IRAS, The Infrared Space Observatory (ISO), Spitzer Space Telescope (SST), Akari, Herschel, and Wide-field Infrared Survey Explorer (WISE) revealed that distant cold debris disks, $< 100$ K, were common. About 20% of nearby solar-type stars are surrounded by an optically thin, gas-poor disk of cold debris (Trilling et al. 2008; Carpenter et al. 2009; Eiroa et al. 2013). This proportion is even greater in binary systems (Beichman et al. 2006). These cold debris disks
are distinguished from proto-planetary disks by their significantly greater age, ranging from hundreds of Myr to Gyrs, lack of gas or accretion onto the star, and by being optically thin. The cold reservoir of dust is likely embedded with asteroids and comets, analogous to our own Kuiper belt. The dust is not primordial; the stars are old, the dust is small, < 100 µm, and has a short lifespan due to collisions. Some of these disks have been resolved by visible or sub-millimeter imaging and generally show a lack of dust in their inner regions. This inner void may be the result of dynamical interactions with unseen planets.

The Far Infrared (FIR) excess from cold debris disks is not the only source of excess IR emission. Just as our solar system has a zodiacal disk of warm and hot dust, other stars have exozodiacal disks. There are hot (> 1000 K) components to some dust disks, which show up as excess in NIR emission. Details on exozodiacal disks are covered in Roberge et al. (2012); Kennedy & Wyatt (2013); Kennedy et al. (2015).

NIR interferometric surveys with FLUOR (42 stars, K-band) and PIONEER (88 stars, H-band) instruments indicate that 15-20% of solar-type stars have hot, exozodiacal dust (Absil et al. 2013; Ertel et al. 2014), and this is probably an underestimate due to instrumental limitations. Thus, we have an architecture of stellar systems that may be composed of a distant cold debris disk (< 100 K), a warm exozodiacal disk (≈ 300 K), and a hot inner dust disk (1000 K or more) Mennesson et al. (2014).

This NIR emission is easier to detect than the MIR emission originating from a warm exozodiacal dust population. There is likely a connection between these hot interferometrically detected dust disks and the harder to detect warm exozodiacal dust in the
habitable zone. In this way, NIR interferometric studies can observe the tip-of-the-iceberg of stellar systems exozodiacal dust, providing details such as composition and grain size of dust, as well as statistics on the correlation of dust populations and stellar properties.

Inner dust regions may exhibit a high degree of variability and may hint at the dust origin and replenishment mechanisms. Many of these disks reach the sublimation radius of their host star even though PR drag or radiation pressure should clear the region. In the reference frame of a dust grain orbiting a star, the light from the star appears to be coming from slightly forward along the grain’s orbit. This effect is known as stellar aberration. As a result of this, when the grain absorbs light from the star there is a small component of the force which acts in opposition to the direction of the grain’s orbit. This leads to a drop in the grain’s angular momentum, ultimately causing it to spiral in toward the star.

Models suggest that dust located in the region where NIR disks are found has a removal timescale on the order of a few years (Wyatt 2008b). These dust grains should either be dragged into the star by PR drag or blown out by radiation pressure, yet interferometric surveys have resolved warm and hot dust around a large fraction of stars observed (Absil et al. 2006). The dust must be continually replaced or dust production may be punctuated and chaotic due to collisional grinding and/or comet infall. This situation would lead to variability in the density of the disk and thus a variation in the incoherent flux. A third possibility is that small grains are trapped in the host star’s magnetic field (Su et al. 2013). Continual replenishment is unlikely (Lebreton et al. 2013), so it remains to be determined if the system in a steady state or varies over time.
Study of warm dust disks close to stars is crucial for a deeper understanding of planetary systems and their formation. These dust disks are common in stellar systems including those known to host exoplanets (Meyer et al. 2007). Exoplanets may shepherd the dust, forming gaps in the disk that reveal undetected exoplanets or resonances, and clumping effects within the disk may mimic exoplanets. A comprehensive picture of the nature of this exozodiacal phenomenon is important to future exoplanet detections. If abundant, dust in this zodiacal analog region confounds exoplanet detections by scattering light, hiding or mimicking planetary emission. The origin and composition of dust on this region and its relation to other reservoirs of dust is of particular interest for planetary formation models.

Interferometry with the CHARA Array provides the angular resolution necessary to detect NIR excesses originating within 1 AU of the star. The recently upgraded JouFLU is capable of measuring the NIR visibility to a very high precision ($< 0.3\%$), hence providing high dynamic ranges at very small angular separations. The capability of CHARA/JouFLU to study this inner region to such high precision is unique in the Northern hemisphere.

### 5.1 Timeline of Discoveries

Since the first hints of exozodiacal dust were detected, the field has progressed rapidly from studies of individual stars to large-scale survey efforts. Recent significant discoveries include:

- 2001: Hints of an excess around Vega with PTI. (Ciardi et al. 2001)
• 2006: CHARA/FLUOR detection around Vega, 1.29 ± 0.19% (Absil et al. 2006)

• 2007: FLUOR, ε Eri (no detection) & τ Ceti (detection) (di Folco et al. 2007)

• 2008: 5 non-detections & ζ Aql (Absil et al. 2008b)

• 2009: β Leo & ζ Lep detections (Akeson et al. 2009)

• 2009: VLTI/VINCI, Fomalhaut (Absil et al. 2009)

• 2011: IOTA/IONIC detection around Vega (Defrère et al. 2011)

• 2011: Palomar Fiber Nuller (PFN) non detection of Vega: new constraints on location (Mennesson et al. 2011a)

• 2011: Coronagraphs see predicted companions (Mawet et al. 2011)

• 2011-12: First spectroscopic detections of very hot excesses ((Lisse et al. 2012; Weinberger et al. 2011)

• 2012: VLTI/PIONIER detection around β Pic (Defrère et al. 2012a)

• 2013: CHARA initial survey of 40+ single MS stars says it is fairly common (11/40). (Absil et al. 2013)

• 2014: VLTI/PIONIER survey, larger VLTI dispersed H-band survey sees it too, but less often (9/85) (Ertel et al. 2014)

• 2014: VLTI/PIONIER survey - binary companions (Marion et al. 2014)
5.2 The Observables

Exozodiacal dust produces deficits in interferometric visibility and excesses in infrared emission. For interferometric observations the signature of a disk lies between that of a uniform disk and an annular ring. The resulting visibility function is:

\[ V^2 = [f V_{CSE} + (1 - f) V_*]^2, \]  

which can be written as,

\[ V^2 \approx (1 - 2f)^2 V_*^2, \]  

or

\[ V^2 \approx (1 - 2f) \left( \frac{2J_1(\pi b \theta / \lambda)}{\pi b \theta / \lambda} \right)^2, \]  

where \( f \) is the disk/star flux ratio, \( b \) is the baseline, \( \lambda \) is the wavelength, and \( \theta \) is the angular size. Equation 5.3 holds true as long as the circumstellar emission is extended enough to be over resolved at the shortest baseline, that is \( V_{CSE} \approx 0 \) for \( b > b_{\text{min}} \). Equation 5.4 is true if \( f \ll 1 \) is also true. di Folco et al. (2007) provides a derivation of the V deficit formulation.

Conceptually, one could imagine simple fringes, for example from Young’s double-slit experiment, assuming a visibility of 1, or perfect contrast with zero flux in the dark fringes and 100% flux in the bright fringes. If a source of incoherent light is added, the fringes will remain,
but the dark fringes will have some additional flux. Perfect contrast is lost, and the visibility is reduced. This is analogous to the addition of a disk around a star, it adds an incoherent source of flux that reduces the contrast of the fringes and lowers the system visibility.

Figure 5.1 shows the visibility curve for an A0V star and the visibility curve for the same star with a 1% disk. The figure also shows the resulting visibility error for a given stellar diameter error and baseline. Figure 5.2 gives an example SED fit, illustrating the precision of FLUOR interferometric results compared to photometric measurements.

The CHARA/FLUOR FoV of 0.9′′ Full Width at Half Maximum (FWHM) and VLTI/PIONIER FoV of 0.4′′ FWHM coupled with the separation sensitivity of the PFN (See Figure 5.3) suggest that location of the excess must be very close in to the star’s surface. The PFN results put an upper limit on the excess flux for Vega at different separations that is well below the detection level of CHARA/FLUOR unless the excess flux is from a source that is closer than 0.2 AU or greater than 2 AU from Vega (Mennesson et al. 2011a).
Figure 5.1: (top) Expected visibility for an A0V star (solid line) and for the same star surrounded by a (fully resolved) exozodiacal disk producing 1% of the photospheric flux at K band (dashed line). Note that the visibility deficit is larger at short baselines. The solid line in the V deficit plot is the resulting change in visibility with increasing baseline for a 5% error in the stellar diameter. At short baselines, the visibility deficit dominates over error in the stellar diameter measurement. (bottom) The resulting visibility error for a 5% error in a star’s diameter for each baseline. For this program we use the shorter baselines to measure the visibility deficit. The program goal is \(< 0.1\%\) error in dust disk visibility.
Figure 5.2: Example fit to the SED of $\tau$ Ceti showing the FLUOR data point along with published photometric measurements with 1-$\sigma$ error bars (grey boxes), a NextGen spectrum (dot-dash line), a 60 K blackbody (dotted line), the residual excess with 1-$\sigma$ error bars (diamonds), and the total emission of the hot disk obtained by subtracting the best-fit stellar spectrum (solid line), including scattering. Also shown is just the thermal emission from the grains (dashed line). (image credit: di Folco et al. 2007)

Figure 5.3: This plot illustrates the relative flux constraints on a geometrically thin annular dust ring located at various distances from Vega. The CHARA curve (dash-dot line with 1 $\sigma$ boundaries) is derived from the excess reported by Absil et al. (2006), while all other curves are upper limits (3 $\sigma$). In the inner 0.1 to 1 AU region, the PFN data are the most constraining. For the CHARA and PFN measurements to be compatible, any significant near infrared excess can only reside within 0.2 AU and therefore is of thermal or scattering origin, otherwise it must lie outside of 2 AU of Vega and would come from scattering, a less likely solution. (image credit: Mennesson et al. 2011b)
Figure 5.4: A toy disk model with Gaussian edges and a uniform disk star. The solid lines represent the intensity of the model star+disk system, the dotted lines are various Gaussian curves which are used to define the toy model's shape (i.e., the Gaussian damped edges). The height of the solid line represents the relative flux, with the central star having a flux of 1.0. The dust model has a central hole where there is no flux from the star or disk, representing an inner sublimation region. The x-axis is the angular distance from the stellar center. The scale of the axes is arbitrary.

Figure 5.5: The Fourier transform of the toy disk model. Here the y-axis is the visibility in the K-band and the x-axis is the baseline. Note the drop in visibility at short baselines. The top, black solid line is the visibility curve of a naked star, the cyan curve below it is the visibility of the star+disk. The bottom, black solid line is the visibility curve of the disk alone. The “ringing” of the function is due to the properties of taking the Fourier transform of a sharp edge.

Figure 5.6: A Toy Model and its Transform
5.2.1 Models

Extensive work has gone into attempting to model the composition of the dust. As with any modeling, numerous assumptions must be made initially, which may then be refined as additional data are acquired. To model exozodiacal dust, it is assumed that the flux is coming either from thermal emission or from scattering from the dust; which, based on the distance which has been constrained by observations, is likely very close to the star and therefore approaches its sublimation temperature. Of course, the sublimation temperature depends upon the composition of the dust; and the emission properties of the dust depend heavily upon the dust grain size; it is therefore evident that models can rapidly escalate into multiple dimension parameter-spaces.

Current modeling of exozodiacal dust requires high precision stellar photosphere models. These typically take input values of log $g$, effective temperature, and stellar luminosity and generate the stellar SED. This energy output from the star excites the various dust grains. How this energy interacts with the dust first depends upon the geometry and location of the dust disk. For simplification, an optically-thin, two-dimensional model that follows a power law is often used. Finally the dust grain properties must be accounted for. Dust emissivity is highly characterized by grain size and composition. Dust is currently thought to be sub-micron in size and composed of either carbonaceous material like graphite or silicates such as olivine. Lebreton et al. (2013) include a thorough discussion of the dust modeling procedure.
5.2.2 Suitability of JouFLU

Short, 34 to 65-m, baselines give access to the inner region of the exozodiacal disk, but the measurements require extremely high precision (see Figure 5.1). The JouFLU beam combiner at the CHARA is uniquely suited to performing these observations, providing <0.3% precision on 1% stellar excess hot/warm dust disks. The single mode fibers provide spatial filtering, improving precision (Coudé du Foresto et al. 1997), while simultaneous recording of the photometric signal further improves precision (Perrin 2003). JouFLU is capable of spectrally dispersed visibility measurements. The baselines available at the CHARA Array range from 34 to 331 m. The stellar photosphere is constrained by models and/or direct long baseline diameter measurements, while the shorter baseline observations are used to detect visibility deficit due to the disk.

Incoherent flux produces a visibility deficit on all baselines. To measure this requires < 1% precision on measurements of V. Figure 5.7 shows the precision required to discern between disk models, and Figure 5.1 shows the impact of stellar diameter error on the visibility curve. From this, it is apparent that in order to detect small visibility drops, shorter baselines are best suited. For example, for an unresolved star, < 1 mas or the sun at 10 parsec (pc), the photosphere can be ignored at these baselines.

JouFLU is currently involved in a large survey of spectral types A through K stars that potentially harbor exozodiacal dust. The goal is to provide statistics about dust disk occurrence in relation to their host stars and the presence of cold dust reservoirs.
Figure 5.7: This plot shows the expected visibility for various disk models for Vega as a function of baseline and H band data from IOTA/IONIC. The models include: the limb-darkened stellar photosphere with a 5% uncertainty on its diameter, the star with a uniform disk, the star with a zodiacal disk, and the star with a narrow ring. The data point colors indicate which night the data were taken and with what polarization state. (image credit: Defrère et al. 2011)

Complementing this survey is a project to re-observe the earliest excess detections in order to determine their variability. In addition, NASA’s InfraRed Telescope Facility (IRTF) provides a method for spectrophotometric detections of excess stellar flux corresponding to the presence of hot/warm exozodiacal dust. Combining multiple NIR interferometric instruments as well as medium-resolution spectroscopy gives a sensitive and efficient method of discovering inner disks and characterizing nearby habitable zone environments.
5.3 System Architecture

In the schematic (Figure 5.8), the cold debris disk is the outer broad ring in the left hand plot. Debris disks are thought to be left over from planetary formation or the Late Heavy Bombardment (LHB) period. These disks have been detected by the measurement of FIR excesses, through sub-millimeter (sub-mm) imaging, visible imaging, and radio interferometry. These debris disks may exhibit asymmetries and resonances which could both indicate or mimic the existence of a planet (Kuchner & Holman 2003). Disks may not be smooth and general asymmetries, bands, or clumps can be formed by a planet. These clumps may have a different orbital period than a planet and show a spectrum similar to the planet. Clumpy structures in the dust distribution may point toward dynamical interaction with planets (Stark & Kuchner 2008). A clump may also be misinterpreted as an actual planet, since bright dust clumps may outshine the planetary flux if there are more than ten zodis of flux (Defrère et al. 2012b). It is clear that in order to reach exo-Earth imaging, we need to measure the dust and know its location and effect on imaging.

Much closer in to the star we find exozodiacal analogs. This circumstellar dust is closer than 10 AU and ranges from $< 1 - 100 \, \mu m$ in size in the inner solar system. Table 5.1 shows how close to the host star, in stellar radii, that JouFLU is capable of resolving.

From the current observations performed by many investigators using multiple instruments and facilities, we can provide some constraints on the location of the exozodiacal dust in a system. The FoV for FLUOR is roughly 10 AU at 10 pc. This corresponds to the right hand
The innermost bright ring of this diagram represents the location of the hot exozodiacal dust. It is of great interest to determine the source of this dust and if there exists any correlations between the presence of dust in this region with the stellar properties, the existence of more distant warm dust which may be located in the region of habitable zone planets, and the distant, cold, outer dust reservoir.

Exozodiacal and debris disks are tenuous but huge in surface area. The inner warm (300 K) region of dust disks remains largely unprobed. Such observations require high dynamic range and high angular resolution. For an example of just how difficult it is to probe this region see the Kalas et al. (2008) Hubble Space Telescope (HST) coronagraphic images of
exosolar planet Fomalhaut b and its dust belt.

Table 5.1: Potential excesses would be resolvable as close as 5 mas with JouFLU.

<table>
<thead>
<tr>
<th>Object</th>
<th>Spectral Type</th>
<th>$R_{\text{star}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\tau$ Cet</td>
<td>G8V</td>
<td>4.7</td>
</tr>
<tr>
<td>10 Tau</td>
<td>F9V</td>
<td>8.7</td>
</tr>
<tr>
<td>$\eta$ Lep</td>
<td>F1V</td>
<td>9.2</td>
</tr>
<tr>
<td>$\lambda$ Gem</td>
<td>A3V</td>
<td>11.3</td>
</tr>
<tr>
<td>$\beta$ Leo</td>
<td>A3V</td>
<td>7.0</td>
</tr>
<tr>
<td>$\xi$ Boo</td>
<td>G8V</td>
<td>7.9</td>
</tr>
<tr>
<td>Altair</td>
<td>A7V</td>
<td>2.8</td>
</tr>
<tr>
<td>Vega</td>
<td>A0V</td>
<td>2.8</td>
</tr>
<tr>
<td>110 Her</td>
<td>F6V</td>
<td>9.6</td>
</tr>
<tr>
<td>$\zeta$ Aql</td>
<td>A0V</td>
<td>11.0</td>
</tr>
<tr>
<td>$\alpha$ Cep</td>
<td>A7IV</td>
<td>6.0</td>
</tr>
</tbody>
</table>

In this inner region, which corresponds to our interferometric FoV, either an optically thin/low emissivity disk or a single bright, narrow ring could generate the visibility deficits that we observe. We have no way to break this ambiguity without ancillary data.

Kate Su’s multiple publications (Su et al. 2013, 2005, 2006) which address this topic combine FIR and MIR spectra that suggest the existence of two disk populations with different temperatures. However, the data cannot provide accurate temperature estimates.

Jeremy Lebreton finds that blackbody fits get the location of the disk wrong. Also, based on his modeling work the dust is smaller than 1 micron, the same size as typical smoke particles; so perhaps smoke would be a better term than dust for general public discussion (Lebreton 2015). He also finds indications that two dust populations within a few AU may be common.

The best models suggest a hot ring with a warm belt surrounding it. The warm belt transports dust to the hot ring by PR drag, then the dust is broken down. Dust tends to accumulate next to the sublimation region, typically at a radius of 0.1 to 0.5 AU for an A-type star like Vega,
and 0.01 to 0.02 AU, or a few stellar radii, for a solar-type star. The inner region of the dust may not have a distinct edge either. Kimura & Mann (1998) and Hahn et al. (2002) show that, in our solar system, the distribution extends continuously and with increasing surface density from a few AU from the Sun down to a fraction of an AU where it forms the Solar F-corona.

5.4 Composition

There have been many substances hypothesized as the primary dust material. The most common of these are: olivines, particularly fosterite, astrosilicates, SiC, SiO, fayalite, enstatite, carbonaceous compounds such as graphite, iron, quartz, corundum, and water ice. Independent work by Kate Su and Carey Lisse is highly suggestive of magnesium-rich olivine as the main constituent of the dust (Lisse 2015; Su 2015). Forsterite is magnesium-rich olivine, \( \text{Mg}_2\text{SiO}_4 \); the common gemstone peridot is the familiar form of this.

There is a general consensus that the dust grains range from \( \leq 1 \) – 100 \( \mu \text{m} \) in size, with sub-micrometer dust being likely. Exozodiacal dust detected with NIR interferometry is composed of sub-micrometer sized grains that are much smaller than the radiation blow-out size, wherein radiation pressure removes the grain. The dust likely has a high albedo and undergoes weathering processes. It may be generated as debris from comets, asteroids, collisions, and outgassing.

Carey Lisse has an on-going IRTF Spectroscopic Survey (Lisse 2015). From this, he has found that the Kuiper belt dust is dark and is only detected in FIR, NIR dust is low-iron olivine,
not pyroxene or water ice or carbon, and MIR dust looks like cometary dust. Also, there are two populations of dust, of different composition, with a 5% inclination in the height of dust, which is similar to the solar system’s asteroid belt. Possibly this started out as cometary dust, but when brought in close to star the organics were removed, leaving olivine.

Jeremy Lebreton also finds astrosilicates and forsterite as plausible components of the dust. Forsterite is more likely, but it must have a high albedo. As soon as forsterite is contaminated even slightly, its albedo drops (Lebreton 2015).

5.5 Limiting Factors

Inner dust is extremely short-lived, lasting only a few years (Wyatt 2008b). This region may be highly chaotic and variable over short timescales. Many factors may remove dust from this region: sublimation, radiation pressure blow-out, collisions, and Poynting-Robertson (P-R) drag in toward the star. Which factor is dominant depends upon the properties of the dust and of the host star. Collisions between larger bodies may produce dust. Then grain-grain collisions rapidly reduce the size of the dust, in a collisional cascade, ultimately reaching a small enough size that PR drag or blow-out from radiation pressure becomes important.

Dust Destruction factors —

Poynting-Robertson — Due to stellar aberration, or the apparent shift of a light source when viewed from a moving reference frame, from the point of view of the dust grain there is an asymmetry in how light is absorbed relative to the grain’s
orbit. This is the cause of PR effect, which is essentially the tangential component of radiation pressure relative to the grain’s motion. This force opposes the dust grain's orbital motion causing it to lose angular momentum and slowly spiral into the star. PR drag brings $10^{-13}$ Earth masses per year into the exozodiacal zone, and we need $10^{-9}$ Earth masses per year or more to match what is observed Wyatt (2005); Defrère et al. (2011). This, of course, depends upon the lifespan of the dust, trapping, and collisions. There is substantial work that indicates the insignificance of PR drag (Wyatt 2005; Wyatt et al. 2007a; Wyatt 2015). PR drag supplies a small amount of dust to the sublimation zone, but a fully consistent treatment yields a maximum amount of dust that is about 7 times lower than that given by analytical estimates (van Lieshout et al. 2014a). The process is thought to proceed something like this:

1. Grains spiral in toward the star due to PR drag.
2. Sublimation reduces the grain, resulting in stalling drift as PR and radiation pressure balance. Grains pile-up on the inner edge of the sublimation zone.
3. The remains are blown out by radiation pressure, and an inner cavity forms (Kobayashi et al. 2009).

The PR timescale to drag dust onto the surface of the star is

$$\tau_{PR} = 2 \times 10^6 \text{yrs} \left( \frac{r}{50\text{AU}} \right)^2 \left( \frac{a}{\mu\text{m}} \right) \left( \frac{\rho_g}{g \cdot \text{cm}^{-3}} \right) \left( \frac{\pi a^2}{C_{ph}} \right) \left( \frac{L_*}{L_\odot} \right)^{-1}, \quad (5.5)$$
where $C_{\text{ph}} = \pi a^2 Q_{\text{rad}}$ is the effective cross-section, $r$ is the radius of the disk, $a$ is the area of the grains, $\rho$ is the density of the grains, and $L_*$ is the luminosity of the star (Minato et al. 2006).

**Sublimation** — Sublimation, the temperature at which material turns from solid to gas, defines an inner edge to the exozodiacal disk. Dust sublimation has a strong dependence on temperature as well as the size and optical properties of the dust grains. In addition, it is affected by the presence of gas. According to Bryden et al. (2006) the effective temperature of the dust is found by

$$\frac{L_{\text{dust}}}{L_*} = \frac{F_{\text{dust}}}{F_*} \frac{kT_{\text{dust}}^4 \left( e^{\frac{\hbar \nu}{kT_{\text{dust}}}} - 1 \right)}{\hbar \nu T_*^3},$$

where $L$ is the luminosity, $F$ is the flux, $T$ is the temperature in Kelvin, $k$ is the Boltzmann constant, $\hbar$ is the Planck constant, and $\nu$ is the frequency. Akeson et al. (2009) also discuss this topic and the properties of the dust. Lamy (1974) includes sublimation temperatures for various grain compositions and finds that the time it takes for dust to sublimate, $\tau_{\text{sub}}$, for 1 $\mu$m grains may be $< 10^4$ seconds, but smaller grains may persist for $> 10^2$ yr. For amorphous grains, Akeson et al. (2009) uses a sublimation temperature of 1600 K. Sublimation temperatures typically used for “hot” exozodiacal disk analysis range from 1000 K to 1500 K, matching those of silicates (Kobayashi et al. 2011). The sublimation temperature for graphite is 2800 K (Lamy 1974). The sublimation temperatures for other grains are: Olivine, 1300K, and Pyroxene, 1700K Kobayashi et al. (2011).
**Collisions** — (Stark & Kuchner 2009) The collisional timescale between grains in a disk is

$$
\tau_{\text{coll}} = 2 \times 10^3 \text{yrs} \left( \frac{r}{50 \text{AU}} \right)^{\frac{3}{2}} \left( \frac{a}{\mu m} \right) \left( \frac{\rho_g}{\text{g} \cdot \text{cm}^{-3}} \right) \left( \frac{\pi a^2}{S_z} \right) \left( \frac{M_*}{M_{\odot}} \right)^{-\frac{1}{2}} \left( \frac{M_{\text{dust}}}{10^{-3} M_{\odot}} \right)^{-1},
$$

(5.7)

where $a$ is the grain radius, $r$ is the orbital radius of the grain, $\rho_g$ is the dust grain density, and $S_z = \pi a^2 Q_{\text{coll}}$ is the effective grain cross-section with $Q_{\text{coll}}$ being an empirically derived factor (Minato et al. 2006; Najita & Williams 2005). However, dust does not lie in a true disk, but instead is confined to a narrow ring. Plavchan et al. (2009) reformulated $\tau_{\text{coll}}$ to reflect this,

$$
\tau_{\text{coll}} = 2 \times 10^3 \text{yrs} \times \gamma \left( \frac{r}{50 \text{AU}} \right)^{\frac{3}{2}-\alpha} \left( \frac{a}{\mu m} \right) \left( \frac{\rho_g}{\text{g} \cdot \text{cm}^{-3}} \right) \left( \frac{\pi a^2}{S_z} \right) \left( \frac{M_*}{M_{\odot}} \right)^{-\frac{1}{2}} \left( \frac{M_{\text{dust}}}{10^{-3} M_{\odot}} \right)^{-1}.
$$

(5.8)

Here $\gamma$ is a corrective factor and $\alpha$ is an exponential factor proportional to the surface density profile.

**Radiative Blowout** — Radiation pressure is simply defined as,

$$
P_{\text{rad}} = \frac{1}{3} A T^4,
$$

(5.9)

where $A$ is the radiation constant, $A = 4 \sigma / c = 7.565767 \times 10^{-16} J/m^3 K^4$. The balance between the forces of radiation pressure and gravitation is typically depicted by the beta factor,

$$
\beta = \frac{F_{\text{rad}}}{F_{\text{grav}}} = \frac{3 L Q_{\text{PR}}}{16 \pi G M c \rho_g a},
$$

(5.10)

for a spherical particle, where $Q_{\text{PR}}$ is the Mie scattering coefficient, $L$ is the luminosity, $G$ is the gravitational constant, $M$ is the star mass, $c$ is the speed of light, $\rho_g$ is the dust grain density, and $a$ is the area of a grain (Burns et al. 1979).
The associated tangential component of this force is

\[ \beta_{PR} = 5000 \tau_{eff} \sqrt{\frac{r}{M_*}} \]  

(5.11)

where \( \tau_{eff} \) is the effective optical depth, \( r \) is the orbital radius of the dust grain, and \( M_* \) is the mass of the star \text{Wyatt (2005)}.

Particles of different sizes fall into different regimes, often classified by \( \beta \) and \( \beta_{PR} \):

- \( \beta \ll \beta_{PR} \) are confined to a belt
- \( \beta \approx \beta_{PR} \) are affected by PR drag
- \( 0.1 < \beta < 0.5 \) are bound and extended in distribution
- \( \beta > 0.5 \) are blown out by radiation pressure on hyperbolic orbits

This determines the number density of various particle sizes within a dust population \text{Wyatt 2008a}.

**Stellar Wind Drag** — Stellar winds are flows of neutral or charged gas or particles ejected from the upper atmosphere of a star. Like PR drag, stellar wind drag makes the grains lose angular momentum and spiral toward the star. Stellar wind will drag dust into the star on a timescale of

\[ \tau_{SW} = 6 \times 10^6 \text{yrs} \left( \frac{r}{50 \text{AU}} \right)^2 \left( \frac{a}{\mu \text{m}} \right) \left( \frac{\rho_g}{\text{g} \cdot \text{cm}^{-3}} \right) \left( \frac{\pi a^2}{C_{sw}} \right) \left( \frac{M_*}{M_\odot} \right)^{-1} \]  

(5.12)

where \( r \) is the orbital radius of the dust grain, \( a \) is the area of a dust grain, \( \rho_g \) is the dust grain density, \( M_* \) is the mass of the star, and \( C_{sw} = \pi a^2 Q_{sw} \) is the effective
cross-section with $Q_{sw}$ being an empirically derived factor (Minato et al. 2006).

5.6 Dust Sources

The dust production mechanism is poorly understood, but it is most likely from small body dynamics. Given the various factors at work that rapidly destroy or clear away the hot dust, producing the amount of dust detected requires a very active source of replenishment. For Vega, this rate is $10^{-9} M_{\oplus}/\text{yr}$ (Defrère et al. 2011). Competing models exist to explain the persistence of dust, the most prominent are: continuous replenishment, magnetically trapped nano-grains, or comet in-fall/out-gassing (Wyatt 2008b; Su et al. 2013; Lebreton et al. 2013).

Sources of dust — in situ:

**Trapping** — If the dust grains are sufficiently small, i.e., nanograins, then magnetic trapping could take place. Kate Su and others have suggested that the magnetic trapping of nanograins, 10 to 20-nm, of any composition is sufficient to explain the observed properties of dust disks (Su et al. 2013). The problem is how to keep the grains close in to the star. Nanograins charge rapidly by the photoelectric effect and every A star should have a remnant, or fossil field that is not very active and would not reverse direction. In such a case, nanograins, < 10 nm, could be locked in the weak magnetic field. Grains fall into the sublimation radius, then decompose into nanograins. Once captured by the magnetic field, a grain gyrates around the field lines instead of following a Keplerian orbit (Ragot & Kahler 2003; Czechowski
Stellar origin — Winds, circumstellar material, and flares/mass ejection are possible factors that could affect or produce dust. If the source of the excess emission is closer to the host star, these factors must be considered more carefully.

Late Heavy Bombardment — Evidence from our own Solar System suggests that the LHB period of $\approx 3.8$ to $4.1$ Gyr ago was a time of intense collisional activity, and therefore an extremely high dust production rate. However, Bonsor et al. (2013a) shows that the occurrence of catastrophic events due to dynamical instabilities in a planetary system is unlikely, with a less than 1% chance of being observed.

Evaporating planets — Inside an in situ mass reservoir capable of producing dust could be a hot Mercury-like planet (Owen & Wu 2013). The slow evaporation of planets the size of Mercury produce the right amount of dust to explain exozodiacal observations (Rappaport et al. 2012). This equates to $\approx 10^{-9} M_{\oplus}/yr$ (van Lieshout et al. 2014b).

Mean Motion Resonances — Mean Motion Resonances (MMR), or orbital resonances where the ratio of orbital periods can be expressed as the ratio of two small integers, trap dust as it is migrating inwards due to PR drag. More optical depth can accumulate by this method but not by more than a factor 10 over PR drag alone (Shannon et al. 2015).

Steady state — Asteroid belt/ongoing planet formation is a possible source of dust,
but only if the system is very young or if the system's asteroids or planetesimals have extreme eccentricities. Nevertheless, these large objects are a significant potential source of dust mass. The evolution of the steady state size distribution is dependent upon the mass of objects at the large end of the scale. However, the hot dust cannot originate in a recent collision in an asteroid belt. This is because there is a limit to which collisions sufficient to reproduce the dust luminosity can occur. The planetesimal belt in these systems must be located farther from the star than the dust. The total mass of an exozodiacal cloud is on the order of $10^{-10}$ to $10^{-9}M_\oplus$, equivalent to that of an asteroid a few km in radius (Augereau 2009). The disk is also dense enough for collisions to be frequent. Unlike cold debris disks, replenishment in situ by collisions between planetesimals does not work for exozodiacal dust. Any local population of planetesimals will not survive long enough to sustain a sufficiently high level of dust. Mark Wyatt concludes that the most likely origin for transient dust is dynamical instability, during an LHB event wherein planetesimals are scattered inward from a more distant reservoir and releasing dust via collisions and sublimation (Wyatt et al. 2007a).

**Stochastic** — This would be a random large event, such as a recent impact (Wyatt et al. 2007b). This process requires a 20 km body every 3 years or a 100 km body every 300 years in order to produce the observed amount of dust. So collisions are not a likely source of dust (Kral et al. 2015).

**Sources of dust — external:**
**Poynting-Robertson drag** Although explained as a source of dust destruction, PR drag is also a source for dust, as it serves to transport dust inward toward the star, where it warms and becomes a stronger IR emitter, before being eventually destroyed. Reidemeister et al. (2011) presents PR drag as the primary means of transporting dust from the outer Kuiper belt analog to the inner region of the $\epsilon$ Eridani system.

**Comets** — In the Solar System, the evaporation of Jupiter family comets is thought to be the most effective dust production mechanism. 90% of zodiacal dust comes from comets (Nesvorný et al. 2010). For exozodiacal disks, it is possible that this is also true. This could be revealed by observing spectral features of primitive material & impact produced material (Lisse et al. 2012). This would require a very high rate of star-grazing comets. However, this would be compatible with systems undergoing a period of planet/planetesimal formation and LHB. There are multiple possible ways for comets to be brought into the inner regions of a system.

- The outward migration of a planet toward a planetesimal belt similar to the Kuiper belt could be responsible for a cometary bombardment (Augereau 2009).

- Dynamical action — Processes such as MMR, Kozai resonance, or that of the Nice model could deliver comets. The Kozai resonance, or resonance with Jupiter, is the source of most comets in the Solar System, with the ejected body becoming a Sun-grazer (Bailey 1992). The Nice model (Gomes et al. 1996).
2005) illustrates a period of late and intense cometary activity, in an LHB-like event; however, if all exozodiacal disks were formed this way, the detection rate would be only \( \approx 0.1\% \) (Bonsor et al. 2013a). It has been suggested that the observed detections could occur after a single large collision, such as in the case of the 12 Myr old HD 172555 (Lisse et al. 2009); or after a LHB-like event for Vega, \( \sim 500 \) Myr old (Absil et al. 2006), and for \( \eta \) Crv, at 1 Gyr old (Lisse et al. 2012; Wyatt et al. 2007a).

- **Steady State** — The comets are scattered semi-regularly, or dust is dragged in. But this requires an outer belt, and a correlation between cold outer belts, or FIR excesses, has not been confirmed. Cold debris disks are usually explained by steady-state collisional grinding of planetesimals (Backman & Paresce 1993). But there is no correlation with exozodi dust levels and age of system.

- **Stochastic** — Again a recent dynamical instability could deliver comets which then produce the dust.

**Scattering** — Gravitational interactions involving planetesimal dynamics or chains of planets could also work to deliver large amounts of dust to near the sublimation zone (Bonsor et al. 2012). This process works best with tightly packed low-mass planets, so somewhat contrived architectures are necessary. This does function to transport material from an outer belt to the inner parts of the system (Bonsor & Wyatt 2012; Bonsor et al. 2012) and outward planetary migration can sustain
cometary activity on Gyr timescales (Bonsor et al. 2014). This activity fades with time and is often insufficient to explain the brightest exozodiacal disks. It is possible that the phenomenon could have a longer duration if the gravitational feedback of the belt on the planets is taken into account, resulting in a planetesimal-driven migration of the outermost planet(s) (Bonsor et al. 2014; Raymond & Bonsor 2014).

At this point, it may be apparent that there are contradictions and conflicting views on this subject. It is clear that the theory has not yet developed to the point of fully explaining the results the field, as a whole, has observed. Further work on both ends, theory and observation, is needed to explain this topic. It would not be surprising to find multiple methods involved in producing, maintaining, and destroying dust in a complex balance of different physics and stellar environments.

5.7 Key Questions

We generally observe centro-symmetric emission at 1% of the stellar level at H/K bands within ≈ 1” FoV. For CHARA the observations are made with one baseline, while with the Precision Integrated-Optics Near-infrared Imaging ExpeRiment (PIONIER) 6 baselines are used but all the baselines are needed just for detection.

The following is a list of questions we are left with about the nature of this observed phenomenon. These questions, ranging from the general to the specific, indicate that this is a
very active and intriguing topic of research, with much to discover about the nature of this phenomenon, and it is driving the development of more sophisticated theories and modeling. Some questions are my own, some are from the literature, and many are from discussions at the Hot Dust around Main Sequence Stars (HDMSS) meeting at Caltech in May 2015. Also included are questions from email feedback on the meeting:

**Dust** — Questions related to the nature of the dust.

- Is it real? Is it dust? Where does it come from and what are the dust grain properties?
- What are the relationships between the hot inner dust, warm exozodiacal dust, the outer cold debris disk, the host star, and planet formation?
- Is the dust variable? If so, at what rate?
- What is the typical dust grain lifetime and what is the dominant dust removal mechanism?

**Theory** — Questions related to the underlying theory as we understand it now.

- How much gas should be included in the models?
• Is the dust from a rocky or icy source? Is the dust similar to that from a Kuiper belt source, an asteroid belt source, or a cometary source? Something different? Can we tell?

• Models seem to require small dust with high sublimation temperatures, such as carbon, what is the implication for the dust origin?

• Is it safe to assume exozodiacal dust distributions have uniform disk-like shapes? Different geometries could increase the available surface area. Could it favorably impact the collisional rate? Then could it be reflected light? (see Defrère et al. 2012a)

• Can we assume the dust material properties are the same as for cold disks?

• Are we probing the continuum in H and K or could this be broad emission features? How about some sort of PAHs?

• The phenomenological models suggest very small dust, for which blowout may be faster than sublimation for A-type stars. The same tiny dust grains around a G- or K-type star would remain bound, because of the lower stellar luminosity and the low optical efficiencies $Q_{pr}$ at those grain sizes. Does this imply different removal mechanisms for A-type and Sun-like stars? Would that imply qualitatively different dust belt morphologies / dust levels for different stellar types? Is this consistent with the differences seen in the recent surveys?
• If NIR excess is not coming from the Habitable Zone (HZ), can it still tell us about dust in the HZ?

• If the dust is due to comets should it correlate with metallicity and/or planet-formation?

**Observations —** Issues with how observations are made, or suggestions for ones that should be made.

• What is the wavelength dependence?

• Is it thermal emission or scattering?

• Can the differences in detection rates be explained by differences in the reduction codes used by PIONIER and FLUOR?

• Could it be systematic error that comes and goes? Or is it real variability? Need long term monitoring of high excess stars or to find a strong astrophysical correlation.

• What is the repeatability and period of the variability?

### 5.8 The Observing Programs

These are some of the observing programs that have led to our current understanding of the dust phenomenon.
5.8.1 Keck Interferometry Nuller

The Keck Interferometry Nuller (KIN) links the two iconic Keck 10-m telescopes on Mauna Kea into a nulling interferometer with an 85-m baseline (Colavita et al. 2009). A nulling interferometer takes advantage of the destructive interference property of light superposition. The central dark fringe, or null, is positioned over a star, canceling its light. This allows faint circumstellar emission to be detected and enables measurement of high contrast objects. In this way a nulling interferometer can perform the role of a coronagraph, which uses a physical mask to block light from the central star. KIN operates in the N-band (8.0 to 13.0 µm) and has ten spectral channels. Its FoV is \(\sim 300 \text{ mas} \) FWHM at 8.5 µm. It is sensitive to inner dust to 5 mas (0.05 AU at 10 pc).

The KIN was utilized in a survey program of 25 non-dusty main-sequence stars from 2008 until 2011 (Millan-Gabet et al. 2011). This was a subset of a survey of 47 nearby main-sequence stars, including 22 stars with known NIR or FIR excesses (Mennesson et al. 2014). From this survey, one star (\(\eta\) Corvi) showed a large excess, while four more showed significant (> 3\(\sigma\)) excesses (\(\beta\) Leo, \(\beta\) UMa, \(\zeta\) Lep, and \(\gamma\) Oph). The KIN found that excesses were overall more frequent around A-type stars. Statistical results found that stars with known FIR excesses tend toward higher NIR excesses. This hints at a dynamical link between inner, hot- and outer, cold- belt dust populations. No such link was found between known NIR excesses and MIR excesses measured. This suggests that if the visibility deficits are due to dust, the grains must be very hot, very small, and pile-up very close to the sublimation zone.
5.8.2 The Hot Exozodiacal Disks Survey

The long-baseline interferometric survey of hot exozodiacal disks is broken up into four major publications. This represents a united effort of many collaborators, utilizing FLUOR at the CHARA Array and PIONIER at the VLTI.

5.8.2.1 Part 1

The initial stage of the overall survey project first probed two stars, $\epsilon$ Eri and $\tau$ Ceti, with FLUOR (di Folco et al. 2007). A fractional excess of $0.98 \pm 0.21 \times 10^{-2}$ was detected around $\tau$ Ceti. Possible causes of this excess that are not dust, such as a faint binary companion or background star, were excluded. An upper limit was placed on any excess around $\epsilon$ Eridani of $0.6 \times 10^{-2}$.

5.8.2.2 Part 2

The second phase of the survey included six bright A- and F-type stars. A 5 $\sigma$ excess was detected around $\zeta$ Aquila (Absil et al. 2008b). AO and Radial Velocity (RV) observations show a low-mass companion is a likely explanation for this excess. However, a $1.29 \pm 0.31\%$ NIR excess could also match the data. At this point in the survey, Vega remains the only definitive hot dust detection (Absil et al. 2006).
5.8.2.3 Part 3

In 2013, the first statistics on 42 stars observed with FLUOR were presented (Absil et al. 2013). Here, Olivier Absil observed nearby stars ranging in spectral type from A to K. This was a magnitude-limited sample, with $K \approx 4$, that aimed to determine the frequency of dust around these stars and to correlate NIR dust detections with stellar spectral type, age, metallicity, and the presence of cold dust. Roughly 1% NIR excesses were detected around 13 of the 42 stars. Follow-up study revealed that one of these detections ($\epsilon$ Cep) was likely due to a companion. 11 of the detections were around main-sequence stars, resulting in 28% of the sample displaying hot dust. A-type stars were found to be more likely to exhibit dust (See Table 5.2 and Figure 5.9). It is this work that provides the baseline measurement for the exozodiacal dust disk variability study.

<table>
<thead>
<tr>
<th>Spectral Type</th>
<th>A</th>
<th>F</th>
<th>G-K</th>
<th>Total</th>
</tr>
</thead>
<tbody>
<tr>
<td>Cold Disk</td>
<td>8</td>
<td>6</td>
<td>6</td>
<td>20</td>
</tr>
<tr>
<td>No outer disk</td>
<td>4</td>
<td>7</td>
<td>9</td>
<td>19</td>
</tr>
<tr>
<td>Unknown</td>
<td>0</td>
<td>2</td>
<td>0</td>
<td>2</td>
</tr>
<tr>
<td><strong>Total</strong></td>
<td>12</td>
<td>15</td>
<td>15</td>
<td>42</td>
</tr>
</tbody>
</table>
(a) A histogram of the significance of the excesses in the sample. Also shown is a Gaussian distribution (*dotted line*). Note the bimodal distribution.

(b) The significant excess occurrence rate, separated based on spectral type and presence of cold, outer dust reservoirs.

Figure 5.9: Key results from Absil et al. (2013), showing the frequency of excess occurrence among 42 main sequence A-K stars.

### 5.8.2.4 Part 4

The most recent chapter in the near-infrared interferometric survey of debris-disk stars was led by Steve Ertel (Ertel et al. 2014). This project sought to constrain the properties of the hot exozodiacal dust. Here the sample consisted of 92 southern stars, magnitude-limited to $H \leq 5$, observed with PIONIER at VLTI. Nine out of 85 targets displayed a bright H-band excess, making an 11% detection rate. An additional 3 detections were less conclusive. Again, earlier-type stars showed an enhanced detection rate, as did older stars. This study did not confirm a correlation between outer and inner dust reservoirs. Spectrally dispersed data from this study suggest that the dust must either be very hot or that scattered light is the dominant source of NIR dust emission.
Figure 5.10: Results from Ertel et al. (2014), showing the exozodiacal detection rate by spectral type and cold excess.

Figure 5.11: Separate and combined results from the Absil et al. (2013) and Ertel et al. (2014) surveys. (Image from Ertel et al. 2014)
Figure 5.12: This plot shows how temperature affects the normalized spectra. Spectral slope obtained from just a few spectral channels can therefore reveal the dust temperature. (From Ertel et al. 2014)
5.9 Criticisms

Understandably, there is not yet universal agreement as to the nature of exozodiacal dust. Why are some people not convinced? In most cases, the detection level is only a few sigmas and there are possible alternative explanations to match the observations. Troubling is the fact that interferometry typically gives a different visibility when something goes wrong; either in calibration or instrumentally. There is also a spectral type mismatch in using K giants to calibrate A stars. The visibility deficits could be due to a binary that cannot be completely ruled out using imaging constraints, $RV$, or closure phase. Or it could be explained by a background object, but this has a very low probability. There are stellar phenomenon that could influence detections: models of NIR emission from stellar winds around early type stars indicate it should be very close to the star, 1 to 2 $R_\star$, and drop off quickly. Even in the best known stellar model, for the Sun, there is a 2% uncertainty on spectral features. More particular examples have also been called into question, for example $\eta$ Crv might not be representative.

5.10 Response to Criticisms

While many of the criticisms are valid questions that need to be answered, the quantity and variety of the data obtained so far does place some weight to the idea that something is going on here that is not easily dismissed. Is this a real astrophysical effect and not an instrumental systematic error? Alternating between long and short baselines shows the accuracy of the
instrument. Long-baseline measurements fit the photosphere visibility curve very well. Using K giants as calibrators means that if a giant has an excess it would give us a false negative, but never a false positive exozodiacal disk detection. We prefer calibrators that are as bright or brighter than the target for the same reason. RV measures are checked to rule out binarity; but this does not rule out a background star. This is an extremely rare possibility, as a simple back-of-the-envelope calculation shows,

\[ N_{\text{field}} = \left( \frac{\theta}{2} \right)^2 \pi \frac{N_{\text{stars}}}{41253 \text{ deg}^2}, \]  

(5.13)

where \( N_{\text{field}} \) is the number of stars in the instrumental FoV, \( \theta \) is the field of view, \( N_{\text{stars}} \) is the number of stars brighter than a certain magnitude, and \( N_{\text{stars}}/41253 \text{ deg}^2 \) is then the number of stars in the entire sky. This assumes that the stars are distributed homogeneously and are isotropic. The magnitude limit of JouFLU is approximately 5, and we want to be sensitive to \( \sim 1\% \) flux levels, this gives a magnitude limit for possible background contamination of 10 in the K-band. The number of stars from the Tycho catalog with \( K < 10 \) is 620,000. Using the JouFLU FoV, this gives a probability that a background star is in the field of less than one in 13.5 million. Even using the entire Set of Identifications, Measurements and Bibliography for Astronomical Data (SIMBAD) database with \( K < 10 \) still only puts this probability at about one in a half million.

Following-up FLUOR detections with MIRC could use closure phase to rule out possible binary systems, and this work has begun. As to the question of stellar phenomenon and NIR
emission from stellar wind very close to star, 1 to 2 $R_\star$: the spectral slope of free-free emission does not match what is observed, for f-f excess should go up from NIR to MIR. In the particular case, the spectra of $\eta$ Crv shows 4% olivine dip at 0.9 $\mu$m to 1.4 $\mu$m (Lisse et al. 2012). Presently, some form of exozodiacal dust is the most promising answer to the question of what is causing these visibility deficits and spectrophotometric excesses.

5.11 Concluding Remarks

The study of the planetary and habitable zone environment is an emerging scientific field. As such, there are many unanswered questions and observational challenges. However, if we seek to understand this aspect of stars and exo-planets, and ultimately to study stellar systems similar to our own, the fundamental physical processes and interactions that lead to the production and destruction of exozodiacal dust must be further studied. Work at the CHARA Array with JouFLU, VLTI with PIONIER, and at LBTI with the Hunt for Observable Signatures of Terrestrial Systems (HOSTS) survey (Danchi et al. 2014) will continue to improve our understanding of this phenomenon. This region of stellar systems may be dynamic and constantly evolving, and the production of dust may be tied to the formation of planetesimals, exo-planets, exo-comets and their dynamics. Understanding the role of this dust is critical, as the dust may serve as a marker for exo-Earths as well as a hindrance to their detection.
“Studying the behavior of large whales has been likened to astronomy. The observer glimpses his subjects, often at long range; he cannot do experiments, and he must continually try to infer from data that are usually inadequate.”

— Hal Whitehead
6

The Hot Exozodiacal Disks Survey Extension

6.1 Introduction

In an effort to answer some of the questions that face the field of exozodiacal disk research, the Hot Exozodiacal Disks Survey Extension is an on-going, three year exozodi survey of ≈100 nearby MS stars of spectral type A through K. Hot, 1000 to 1500 K, dust is expected in 25 to 30% of MS systems. The goal is to detect 0.5% excesses at the 5 σ level for m_K < 5 and to determine grain properties, disk morphology, and correlations between stellar properties. The presence of hot or warm dust in these systems has been an important question for several years. Neither aperture photometry, nor single-pupil infrared imaging have been able to answer that question. Infrared interferometry has provided the first unambiguous resolved detections of hot dust around main sequence stars (Absil et al. 2006; di Folco et al. 2007; Absil et al. 2008b; Akeson et al. 2009; Absil et al. 2009, 2013). The JouFLU beam combiner at the CHARA Array is currently one of the only instruments in the Northern hemisphere capable of detecting small visibility drops at short baselines, indicating
the presence of circumstellar emission at the 1% level or below.

Extending the hot exozodiacal survey has been funded as part of a NASA Origins of the Solar System project, with Bertrand Mennesson as the principal investigator. Additional financial support has been provided by Georgia State University (GSU) and the National Science Foundation (NSF). We are in the process of extending the existing CHARA/FLUOR hot exozodiacal survey with JouFLU to \( \approx 100 \) stars over 3 years. Our primary goal is to improve on the detection statistics of the initial sample of 40 nearby MS single stars previously surveyed (Absil et al. 2013). So far, we have obtained data on 46 targets. Table 6.1 gives the complete list of possible targets for the survey.

The identification of correlations between hot, warm, and cold debris disks and stellar characteristics is a primary goal for understanding the formation and evolution of planetary systems. Our extension program started successfully in October 2013 with the observation of 16 new science targets and reached a new magnitude limit with JouFLU.

This project increases the statistics of hot debris disks in an attempt to find correlations with either stellar properties, such as spectral type, age, metallicity, rotation velocity, or cold dust population properties. Together with temporal monitoring, this can constrain the dust evolution models that we are developing within two different working groups. In the target list the occurrence of hot dust around stars is compared with the presence of cold disks known from previous far-infrared space missions and stars that lack any disks. By increasing our detection statistics, from both the JouFLU upgrades as well as an increased sample size, the possible scenarios for the high dust replenishment rate required in the inner disk can be
constrained. Once completed, we will combine CHARA K-band measurements with other relevant high contrast observations (VLTI/PIONIER, PFN, LBTI N-band nuller). Finally, we are developing new models and numerical simulations of the dynamical evolution of these dusty systems, including the effect of the gas that is produced by sublimation.
Table 6.1: The FLUOR source list id list gives the identification and various physical details on the FLUOR survey target list. Kcorr is the bandwidth magnitude, corrected for estimated visibility: $K_{\text{corr}} = K - 2.5 \times \log(V_{\text{est}})$, where $V_{\text{est}}$ is the visibility estimated from the LD diameter. Note in the table below, V is the V band magnitude.

<table>
<thead>
<tr>
<th>HD</th>
<th>H4P</th>
<th>Name</th>
<th>Sp Type</th>
<th>RA</th>
<th>DEC</th>
<th>Dist</th>
<th>LD</th>
<th>diam</th>
<th>K</th>
<th>V</th>
<th>Kcorr</th>
</tr>
</thead>
<tbody>
<tr>
<td>166</td>
<td>544</td>
<td>* V49 And</td>
<td>K0 V</td>
<td>00 06 36.7482</td>
<td>+29 01 17.4083</td>
<td>13.70</td>
<td>0.65</td>
<td>4.21</td>
<td>6.13</td>
<td>4.32</td>
<td></td>
</tr>
<tr>
<td>432</td>
<td>746</td>
<td>* bet Cas</td>
<td>F2 V</td>
<td>00 09 10.8518</td>
<td>+59 08 59.2100</td>
<td>17.68</td>
<td>2.10</td>
<td>1.44</td>
<td>2.27</td>
<td>1.47</td>
<td></td>
</tr>
<tr>
<td>656</td>
<td>917</td>
<td>LTT 75</td>
<td>F6 Vp</td>
<td>00 11 15.5817</td>
<td>-15 29 04.7205</td>
<td>18.75</td>
<td>0.34</td>
<td>3.64</td>
<td>8.93</td>
<td>3.64</td>
<td></td>
</tr>
</tbody>
</table>

(...continued...)
6.2 Observing Strategy

To insure the most robust measurements and improve our measurement precision, great care has gone into developing our observing strategy and calibration of interferometric visibilities.

When possible we use three calibrator stars in a bracketing sequence of Cal1 —Object —Cal2 —Object —Cal3 —Object —Cal1. A minimum of five bracketed data points, where each point is a sequence of 150 to 200 interferograms on the science object, are required before data analysis is considered valid.
depending upon target brightness and data quality. Figure 6.1 shows the effect of target brightness on visibility measurement error. If at all possible, the object is observed before transit, close to transit, and after transit; as well as at elevations $> 30^\circ$ in order to reduce the effects of differential refraction from the atmosphere.

![Percent Visibility Error](image)

Figure 6.1: This plot shows the average error for a single calibrated data point as a function of target K band magnitude. For N bracketed points, this error can be reduced by $\sqrt{N}$. (image credit: Paul Nuñez)

### 6.2.1 Calibrator Selection

A process of experimentation and review of the reduced data has revealed that the best quality calibrators are those nearby the object. It is particularly important that the calibrators and object match as closely as possible in elevation and, as a result, have similar airmass. As such, initially-chosen, well-characterized calibrators from the Borde (Bordé et al. 2002) and Mérand (Mérand et al. 2005) catalogs have been rejected in favor of our own selection
criteria. It was found that the uncertainties in the transfer function at different zenith angles is
greater than the uncertainties in calibrator diameter.

To select our calibrators we use a Set of Identifications, Measurements and Bibliography for
Astronomical Data (SIMBAD) astronomical database query (Wenger et al. 2000), here is an
example code used to search by criteria for calibrators for Vega:

rah > 17 & rah < 19 & dec > 23 & dec < 53 & Vmag < 9.5 & Vmag > 0 & Kmag > 0
& Kmag < 3.5 & splum = III & maintypes!=Pu* & maintypes!=sr* & maintypes!=SB*

This search string provides upper and lower bounds to the right ascension hour and the
declination angle. Bounds are placed on both the V and K magnitudes. This search string
also includes several important exclusions for calibrator selection. Setting splum = III limits
our search to class III giants. Pulsating variable and semi-regular variable stars, and known
spectroscopic binary stars are excluded from the potential calibrator list.

Luminosity class III giants are preferred as calibrators instead of A - K class V dwarfs, as it is
possible that the dwarfs could have exozodiacal disks that would give us excesses on the
calibrator and contaminate our measurements. If a calibrator had an IR excess, then it would
lead to a false negative on the target object disk detection, but, importantly, not a false positive.

Spectral type G8 to M1 are limits for our calibrator spectral type, with a preference for late Gs
and Ks. Calibrators with roughly the same K-magnitude, < 0.2 Kmag difference, as the target
object and with similar airmass are always preferred.
After candidate calibrators are selected from the SIMBAD results, they are checked in the Washington Double Star (WDS) catalog (Mason et al. 2001) for multiplicity within a 5″ radius.

Potential variability is checked using American Association of Variable Star Observers (AAVSO) [http://www.aavso.org]; if $\Delta_{\text{mag}} < 0.1$ in V, then the calibrator is kept.

Finally, the most appropriate surface brightness relationship is used to determine the calibrator’s estimated LD diameter (Kervella et al. 2004b; Groenewegen 2004; di Benedetto 1998; Bonneau et al. 2006, 2011). The Kervella et al. (2004a) surface brightness relations are used for dwarfs, and the Groenewegen (2004) relations for all Giants but M-type. Otherwise, the di Benedetto (1998) surface brightness relationship is used with a 5% error bar if $V - K$ is within $-0.1$ to $3.7$; outside of this range and within ($V - K = -1.1$ to $7.0$) The Bonneau et al. (2006, 2011) surface brightness relationships were used. Extinction corrected values for V and K magnitudes should be used in these cases. Calibrator surface brightness estimated diameters should be $<2$ mas to reach our goal $V^2$ error $<0.2\%$.

Finally, from the remaining calibrator candidates, the three that will be used for the object are chosen based on their right ascension: Calibrator 1 should be 10 arcminutes earlier than the target object in RA; Calibrator 2 should be 10 arcminutes after the target object in RA; and calibrator 3 should be closest to the target object in RA.

After any data are obtained and before they are reduced, the data from calibrators that were used for that bracket are reduced and are cross-checked against one another. Any calibrator showing evidence of being resolved or a binary, is rejected.
6.3 Results

Interferometrically observed exozodiacal disks are typically presented as an estimated K-band disk/star flux ratio, or $f_{cse}$. This is discussed in di Folco et al. (2007) and Absil et al. (2008b). Circumstellar emission with uniform surface brightness that fills the interferometric FoV gives a drop in measured visibility squared compared to the expected visibility squared, or:

$$V^2 \simeq (1 - 2f_{cse}) V_{\star}^2,$$

where $V^2$ is the measured visibility squared, $V_{\star}^2$ is an estimated squared visibility of the stellar photosphere, and $f_{cse}$ is the disk/star flux ratio.

Table 6.2 gives some details about the observing parameters and conditions.
Table 6.2: Notes for exozodiacal survey extension targets.

<table>
<thead>
<tr>
<th>HD</th>
<th>Object</th>
<th>Obs date</th>
<th>Baseline</th>
<th>Number of data points</th>
</tr>
</thead>
<tbody>
<tr>
<td>5448</td>
<td>37 And</td>
<td>8/13/2015, 8/14/2015</td>
<td>S1-S2</td>
<td>5</td>
</tr>
<tr>
<td>14055</td>
<td>γ Tri</td>
<td>10/11/2013, 10/14/2013</td>
<td>E1-E2</td>
<td>12</td>
</tr>
<tr>
<td>15335</td>
<td>13 Tri</td>
<td>10/16/2013, 10/19/2013</td>
<td>E1-E2</td>
<td>7</td>
</tr>
<tr>
<td>19373</td>
<td>υ Per</td>
<td>10/13/2014</td>
<td>S1-S2</td>
<td>8</td>
</tr>
<tr>
<td>20630</td>
<td>κ01 Cet</td>
<td>10/17/2013, 10/18/2013</td>
<td>E1-E2</td>
<td>7</td>
</tr>
<tr>
<td>23249</td>
<td>δ Eri</td>
<td>10/11/2013, 10/14/2013</td>
<td>E1-E2</td>
<td>11</td>
</tr>
<tr>
<td>26965A</td>
<td>40 Eri</td>
<td>10/13/2014</td>
<td>S1-S2</td>
<td>8</td>
</tr>
<tr>
<td>28355</td>
<td>b Tau</td>
<td>10/12/2013, 10/18/2013</td>
<td>E1-E2</td>
<td>11</td>
</tr>
<tr>
<td>34411</td>
<td>λ Aur</td>
<td>10/16/2013, 10/17/2013</td>
<td>E1-E2</td>
<td>12</td>
</tr>
<tr>
<td>87901</td>
<td>α Leo</td>
<td>11/10/2014</td>
<td>S1-S2</td>
<td>6</td>
</tr>
<tr>
<td>162003A</td>
<td>ψ01 Dra A</td>
<td>7/23/2014</td>
<td>E1-E2</td>
<td>6</td>
</tr>
<tr>
<td>164259</td>
<td>ζ Ser</td>
<td>6/15/2015</td>
<td>E1-E2</td>
<td>10</td>
</tr>
<tr>
<td>165777</td>
<td>72 Oph</td>
<td>6/19/2015, 6/20/2015</td>
<td>E1-E2</td>
<td>12</td>
</tr>
<tr>
<td>168151</td>
<td>36 Dra</td>
<td>7/23/2014</td>
<td>E1-E2</td>
<td>6</td>
</tr>
<tr>
<td>182572</td>
<td>b Aql</td>
<td>7/23/2014, 6/15/2015, 6/20/2015</td>
<td>E1-E2</td>
<td>6</td>
</tr>
<tr>
<td>182640</td>
<td>δ Aql</td>
<td>10/13/2014</td>
<td>S1-S2</td>
<td>8</td>
</tr>
<tr>
<td>184006</td>
<td>τ Cyg</td>
<td>6/20/2015</td>
<td>E1-E2</td>
<td>6</td>
</tr>
<tr>
<td>187691A</td>
<td>o Aql</td>
<td>6/20/2015</td>
<td>E1-E2</td>
<td>8</td>
</tr>
<tr>
<td>190360</td>
<td>LHS 3510</td>
<td>10/16/2013, 10/19/2013</td>
<td>E1-E2</td>
<td>8</td>
</tr>
<tr>
<td>202444</td>
<td>τ Cyg</td>
<td>11/12/2014</td>
<td>S1-S2</td>
<td>5</td>
</tr>
<tr>
<td>210418</td>
<td>θ Peg</td>
<td>10/12/2013, 10/14/2013</td>
<td>E1-E2</td>
<td>8</td>
</tr>
<tr>
<td>213558</td>
<td>α Lac</td>
<td>10/11/2013, 10/12/2013</td>
<td>E1-E2</td>
<td>11</td>
</tr>
<tr>
<td>215648</td>
<td>LHS 3851</td>
<td>8/13/2015</td>
<td>S1-S2</td>
<td>10</td>
</tr>
<tr>
<td>217014</td>
<td>51 Peg</td>
<td>10/16/2013, 10/19/2013</td>
<td>E1-E2</td>
<td>9</td>
</tr>
<tr>
<td>219134</td>
<td>HR 8832</td>
<td>10/13/2014</td>
<td>S1-S2</td>
<td>8</td>
</tr>
<tr>
<td>222368</td>
<td>ι Sc</td>
<td>10/17/2013, 10/18/2013</td>
<td>E1-E2</td>
<td>8</td>
</tr>
</tbody>
</table>

6.4 Discussion

Table 6.3 lists the exozodiacaal flux ratio for each target and its corresponding significance of excess $\sigma = \frac{f_{\text{CSE}}}{\sigma_{f_{\text{CSE}}}}$. This information is shown in Figure 6.2. Figure 6.3 shows a
Below are the results of data reduction for 26 of the survey targets, for which we have histogram of these results.

Table 6.3: Exozodi Survey Extension Results, $f_{\text{cse}}$, along with reduced $\chi^2$ for each, and the significance of excess $\sigma = f_{\text{CSE}}/\sigma_{f_{\text{CSE}}}$.

<table>
<thead>
<tr>
<th>HD</th>
<th>Object</th>
<th>$f_{\text{cse}}$ (%)</th>
<th>$\chi^2$</th>
<th>$\sigma$</th>
</tr>
</thead>
<tbody>
<tr>
<td>5448</td>
<td>37 And</td>
<td>2.943 ± 0.505</td>
<td>0.716</td>
<td>5.8</td>
</tr>
<tr>
<td>14055</td>
<td>$\gamma$ Tri</td>
<td>-1.396 ± 1.080</td>
<td>0.891</td>
<td>-1.3</td>
</tr>
<tr>
<td>15335</td>
<td>13 Tri</td>
<td>0.605 ± 1.415</td>
<td>0.651</td>
<td>0.4</td>
</tr>
<tr>
<td>19373</td>
<td>$\iota$ Per</td>
<td>-0.356 ± 0.280</td>
<td>2.267</td>
<td>-1.3</td>
</tr>
<tr>
<td>20630</td>
<td>$\kappa$ 01 Cet</td>
<td>1.393 ± 1.030</td>
<td>0.830</td>
<td>1.4</td>
</tr>
<tr>
<td>23249</td>
<td>$\delta$ Eri</td>
<td>1.404 ± 0.770</td>
<td>0.092</td>
<td>1.8</td>
</tr>
<tr>
<td>26965A</td>
<td>40 Eri</td>
<td>0.011 ± 0.765</td>
<td>1.945</td>
<td>0.0</td>
</tr>
<tr>
<td>28355</td>
<td>b Tau</td>
<td>-0.860 ± 1.235</td>
<td>4.157</td>
<td>-0.7</td>
</tr>
<tr>
<td>34411</td>
<td>$\lambda$ Aur</td>
<td>0.626 ± 0.450</td>
<td>0.797</td>
<td>1.4</td>
</tr>
<tr>
<td>87901</td>
<td>$\alpha$ Leo</td>
<td>-0.118 ± 0.790</td>
<td>0.539</td>
<td>-0.1</td>
</tr>
<tr>
<td>162003A</td>
<td>$\psi$ 01 Dra A</td>
<td>7.005 ± 0.525</td>
<td>1.431</td>
<td>13.3</td>
</tr>
<tr>
<td>164259</td>
<td>$\zeta$ Ser</td>
<td>0.761 ± 0.880</td>
<td>2.748</td>
<td>0.9</td>
</tr>
<tr>
<td>165777</td>
<td>72 Oph</td>
<td>3.269 ± 1.465</td>
<td>1.464</td>
<td>2.2</td>
</tr>
<tr>
<td>168151</td>
<td>36 Dra</td>
<td>1.842 ± 0.760</td>
<td>3.301</td>
<td>2.4</td>
</tr>
<tr>
<td>182572</td>
<td>b Aql</td>
<td>0.030 ± 0.600</td>
<td>2.336</td>
<td>0.1</td>
</tr>
<tr>
<td>182640</td>
<td>$\delta$ Aql</td>
<td>5.138 ± 0.460</td>
<td>1.837</td>
<td>11.2</td>
</tr>
<tr>
<td>184006</td>
<td>$\iota$ Cyg</td>
<td>-0.539 ± 0.875</td>
<td>1.581</td>
<td>-0.6</td>
</tr>
<tr>
<td>187691A</td>
<td>$\phi$ Aql</td>
<td>1.171 ± 1.870</td>
<td>0.478</td>
<td>0.6</td>
</tr>
<tr>
<td>190360</td>
<td>LHS 3510</td>
<td>-0.012 ± 0.540</td>
<td>1.457</td>
<td>0.0</td>
</tr>
<tr>
<td>202444</td>
<td>$\tau$ Cyg</td>
<td>2.833 ± 1.350</td>
<td>3.439</td>
<td>2.1</td>
</tr>
<tr>
<td>210418</td>
<td>$\theta$ Peg</td>
<td>1.727 ± 0.520</td>
<td>2.879</td>
<td>3.3</td>
</tr>
<tr>
<td>213558</td>
<td>$\alpha$ Lac</td>
<td>-1.172 ± 0.860</td>
<td>0.428</td>
<td>-1.4</td>
</tr>
<tr>
<td>215648</td>
<td>LHS 3851</td>
<td>0.269 ± 0.420</td>
<td>3.701</td>
<td>0.6</td>
</tr>
<tr>
<td>217014</td>
<td>51 Peg</td>
<td>-0.032 ± 0.760</td>
<td>0.624</td>
<td>0.0</td>
</tr>
<tr>
<td>219134</td>
<td>HR 8832</td>
<td>0.582 ± 0.540</td>
<td>1.614</td>
<td>1.1</td>
</tr>
<tr>
<td>222368</td>
<td>$\iota$ Psc</td>
<td>1.360 ± 0.275</td>
<td>1.549</td>
<td>4.9</td>
</tr>
</tbody>
</table>
collected five or more data points. For each object, the plots are given for the $V^2$ vs. baseline, hour angle, and phase angle. Here, phase angle is the angle of the baseline projected on the sky, relative to the $U$ coordinate, or $\text{phase angle} = \arctan \left( \frac{V}{U} \right)$. In these plots, the star model is simply the Fourier Transform of a limb-darkened disk. If available, previously measured diameters and limb-darkening coefficients from Boyajian et al. (2012) were used. In cases where measured diameters were not available, the surface-brightness relationships described in subsection 6.2.1 were used. For the star+dust model, Equation 6.1 was used.
Figure 6.3: A histogram of the JouFLU exozodiacal survey extension results with a bin size of 0.4%.
6.4.1 37 And

Figure 6.4: HD5448
6.4.2 γ Tri

Figure 6.5: HD14055
6.4.3 13 Tri

Figure 6.6: HD15335
6.4.4 \( \iota \) Per

Figure 6.7: HD19373
Figure 6.8: HD20630
6.4.6 δ Eri

Figure 6.9: HD23249
6.4.7 40 Eri

Figure 6.10: HD26965
6.4.8 b Tau

![Graphs showing baseline, hour angle, and phase angle measurements for HD28355.](image)

**Figure 6.11: HD28355**
6.4.9 λ Aur

![Graphs showing data for HD34411 with baseline, hour angle, and phase angle plots.](image)

(a) HD34411 baseline

(b) HD34411 hour angle

(c) HD34411 phase angle

Figure 6.12: HD34411
6.4.10 α Leo

Figure 6.13: HD87901
6.4.11 $\psi_{01}$ Dra A

Figure 6.14: HD162003
6.4.12 $\zeta$ Ser

Figure 6.15: HD164259
6.4.13 72 Oph

Figure 6.16: HD165777
6.4.14 36 Dra

(a) HD168151 baseline

(b) HD168151 hour angle

(c) HD168151 phase angle

Figure 6.17: HD168151


Figure 6.18: HD182572

(a) HD182572 baseline

(b) HD182572 hour angle

(c) HD182572 phase angle

6.4.15 b Aql
6.4.16 δ Aql

Figure 6.19: HD182640
6.4.17 \( \epsilon \) Cyg

(a) HD184006 baseline

(b) HD184006 hour angle

(c) HD184006 phase angle

Figure 6.20: HD184006
Figure 6.21: HD187691
6.4.19 HD 190360

(a) HD190360 baseline

(b) HD190360 hour angle

(c) HD190360 phase angle

Figure 6.22: HD190360
6.4.20 $\tau$ Cyg

Figure 6.23: HD202444
6.4.21 θ Peg

Figure 6.24: HD210418
Figure 6.25: HD213558
Figure 6.26: HD215648
6.4.24 51 Peg

Figure 6.27: HD217014
Figure 6.28: HD219134
Figure 6.29: HD222368
6.5 Discussion of individual targets

As part of the exozodiacal survey extension, this is the first time results on exozodiacal disk/star flux ratio have been reported for these objects. We have found five stars with excesses at greater than 1% of the stellar flux and at greater than the 3σ level, four of these are near or above the 5σ level. Most of the strong detections should be followed-up with additional observations to check for binarity and to confirm the excess.

6.5.1 37 And (HD 5448).

This star is an example of a strong excess detection, with $f_{\text{cse}} = 2.94 \pm 0.51\%$. This is a 5.8σ result with a good model fit, $\chi^2 = 0.72$.

6.5.2 γ Tri (HD 14055).

No excess is detected for this star. The negative excess ratio is possibly due to poor calibration and observing difficulties. Note that calibration issues typically result in negative $f_{\text{cse}}$ values, so any positive detections are difficult to attribute to calibration error.

6.5.3 13 Tri (HD 15335).

No excess is detected for 13 Tri, but only three data points passed DRS quality checks. Nevertheless, the model fit is reasonable mainly due to a single high-quality data point.
6.5.4 \( \nu \) Per (HD 19373).

These data are of very good quality, and the resulting non-excess detection has small error bars.

6.5.5 \( \kappa \) 01 Cet (HD 20630).

This star shows the possible presence of an excess, \( f_{\text{cse}} = 1.39 \pm 1.03\% \). However, the error is large and the resulting \( \sigma \) is just 1.4 so no conclusions should be drawn.

6.5.6 \( \delta \) Eri (HD 23249).

\( \delta \) Eri also shows the possible presence of an excess, \( f_{\text{cse}} = 1.4 \pm 0.77\% \). However, in this case the error is smaller, the resulting \( \sigma \) is 1.8, and the model fit is excellent with a \( \chi^2 = 0.09 \).

6.5.7 40 Eri (HD 26965A).

No excess is detected for 40 Eri.

6.5.8 b Tau (HD 28355).

b Tau shows no excess, and the data are of poor quality.
6.5.9  λ Aur (HD 34411).

λ Aur may show a slight excess of $0.63 \pm 0.45\%$, but the confidence for this detection is below the detection criteria at only $1.4\sigma$.

6.5.10  α Leo (HD 87901).

α Leo shows no excess.

6.5.11  ψ01 Dra A (HD 162003).

This is the strongest excess detections, with an excess ratio of $7.0 \pm 0.53\%$. The data pass quality checks and have small error bars. This is a $13\sigma$ excess from six data points.

6.5.12  ζ Ser (HD 164259).

This star could be labeled a tentative, if small, excess. However, conclusions should not be drawn on $< 3\sigma$ detections.

6.5.13  72 Oph (HD 165777).

This is one of the strongest excess detections at $f_{\text{sec}} = 3.27 \pm 1.47\%$. However, the data quality is poor and the errors are large, making this an only $2.2\sigma$ detection. It is recommended this star be re-observed.
6.5.14 36 Dra (HD 168151).

This star could be labeled a tentative excess with a ratio of $1.84 \pm 0.76\%$. However, with such low significance the author refrains from conclusions.

6.5.15 b Aql (HD 182572).

No excess is detected for this target.

6.5.16 δ Aql (HD 182640).

This star shows a disk/star excess of $5.14 \pm 0.46\%$ and is one of the strongest excess detections. This is an $11\sigma$ detection.

6.5.17 ι Cyg (HD 184006).

No excess is detected for this target.

6.5.18 o Aql (HD 187691A).

This is a possible excess detection, but the errors are very large.

6.5.19 LHS 3510 (HD 190360).

This is a non-detection of an excess.
6.5.20  τ Cyg (HD 202444).

This is one of the strongest excess detections in terms of flux ratio percentage. The ratio, \( f_{\text{cse}} \) is \( 2.83 \pm 1.35\% \). However, the overall data quality is poor and the errors are large, making the resulting detection only at the \( 2\sigma \) level.

6.5.21  θ Peg (HD 210418).

θ Peg shows an excess of \( 1.73 \pm 0.52\% \). The data quality is good.

6.5.22  α Lac (HD 213558).

No excess is detected for this star.

6.5.23  LHS 3851 (HD 215648).

No excess is found.

6.5.24  51 Peg (HD 217014).

51 Peg shows no excess.

6.5.25  HR 8832 (HD 219623).

No strong excess is found.
6.5.26 ι Psc (HD 222368).

This star shows a strong excess of $1.36 \pm 0.28\%$ and a good model fit of $\chi^2 = 0.62$. This makes ι Psc an almost five $\sigma$ detection.
“Astronomy is useful because it raises us above ourselves; it is useful because it is grand; . . . . It shows us how small is man’s body, how great his mind, since his intelligence can embrace the whole of this dazzling immensity, where his body is only an obscure point, and enjoy its silent harmony.”

—– Henri Poincaré
The primary goal of this study is to determine if the exozodiacal dust phenomenon is a constant feature of stellar systems or a transient event, possibly the recent of dynamical interactions. Table 7.1 shows the amount of exozodiacal dust excess in relation to the significance of that detection. The strongest indicators of exozodiacal excess variability are found in the stars $\upsilon$ And, $\kappa$ CrB, and $\gamma$ Ser. Two stars $\upsilon$ And and $\kappa$ CrB show an increase in previously detected excesses, while $\gamma$ Ser is a new excess detection where previously there was no excess found. Other stars, such as $\zeta$ Aql, show a change in the significance of the excess detection but not in the overall dust excess flux ratio, or $f_{\text{ce}}$. 
Table 7.1: Variability study results, in terms of significance of excess $\sigma = \frac{f_{\text{CSE}}}{\sigma_{\text{CSE}}}$ for both Absil et al. (2013) and this work. Note that $\sigma$ here denotes the ratio of flux to its Gaussian error in the same convention used in Absil et al. (2013). No conclusion is made on the presence of an excess for 10 Tau or $\alpha$ Lyr for reasons given in the discussion section. Results shown for $\alpha$ Cep are from 2015 and use the rapid rotator model from van Belle et al. (2006) and results from $\alpha$ Aql shown are for the rapid rotator model from Monnier et al. (2007). Note: this table is in terms of excess relative to its significance, not $f_{\text{cse}}$.

<table>
<thead>
<tr>
<th>HD Object</th>
<th>$\sigma^*$ excess</th>
<th>$\sigma^{**}$ excess</th>
</tr>
</thead>
<tbody>
<tr>
<td>9826 $\upsilon$ And</td>
<td>3.1 n 6.3 y</td>
<td></td>
</tr>
<tr>
<td>22484 10 Tau</td>
<td>11.0 y -3.8 —</td>
<td></td>
</tr>
<tr>
<td>40136 $\eta$ Lep</td>
<td>4.2 y 0.9 n</td>
<td></td>
</tr>
<tr>
<td>102647 $\beta$ Leo</td>
<td>3.6 y 2.1 n</td>
<td></td>
</tr>
<tr>
<td>131156 $\zeta$ Boo</td>
<td>3.7 y -0.8 n</td>
<td></td>
</tr>
<tr>
<td>142091 $\kappa$ CrB</td>
<td>5.9 y 6.5 y</td>
<td></td>
</tr>
<tr>
<td>142860 $\gamma$ Ser</td>
<td>-0.2 n 3.1 y</td>
<td></td>
</tr>
<tr>
<td>161868 $\gamma$ Oph</td>
<td>0.5 n -0.8 n</td>
<td></td>
</tr>
<tr>
<td>172167 $\alpha$ Lyr</td>
<td>4.7 y 2.4 —</td>
<td></td>
</tr>
<tr>
<td>173667 110 Her</td>
<td>3.8 y -1.6 n</td>
<td></td>
</tr>
<tr>
<td>177724 $\zeta$ Aql</td>
<td>6.3 y 3.2 y</td>
<td></td>
</tr>
<tr>
<td>185144 $\sigma$ Dra</td>
<td>0.9 n -1.1 n</td>
<td></td>
</tr>
<tr>
<td>187642 $\alpha$ Aql</td>
<td>12.8 y 8.3 y</td>
<td></td>
</tr>
<tr>
<td>203280 $\alpha$ Cep</td>
<td>4.8 y -0.2 n</td>
<td></td>
</tr>
</tbody>
</table>

* Absil et al. (2013)

** this work

7.1 Introduction

Infrared interferometry has provided the first unambiguous resolved detections of hot dust around main sequence stars (Absil et al. 2006), showing an unexpectedly dense population of sub-micrometer dust grains close to their sublimation temperature. Current models of circumstellar debris disks suggest that for the inner region within one AU of the disk, the
timescale for complete removal of this dust is on the order of a few years (Wyatt 2008b).

Interferometric surveys have resolved warm/hot dust around a large fraction of stars observed. The presence of dust close to the star is surprising because most cold debris belts detected are collisionally dominated. Mutual collisions grind the dust down to the size where radiation pressure pushes the dust out before Poynting-Robertson drag has a chance to pull the dust inward.

As discussed in Chapter 5, competing models exist to explain the persistence of this dust. In a steady state model, the dust would need to be continually replaced, but this is unlikely as the timescale for radiation pressure to remove the warm dust is on the order of weeks (Lebreton et al. 2013). A bombardment model, in which dust production is a punctuated and chaotic process fueled by asteroid collisions and comet infall, would show variability on timescales of a few years (Wyatt 2008b). This model suggests that these systems are undergoing a several million year period of instability where asteroid grinding and comet outgassing are producing the dust observed. In this case, some variations should be present in high-precision data over time scales of ten years or less. The rate of material production required for a system like Vega is $10^{-9} \, \text{M}_\odot/\text{yr}$ (Defrère et al. 2011). Recently a third model has emerged wherein the star's magnetic field is responsible for trapping nano-dust grains (Su et al. 2013). Probing variations in this region could help determine whether it is a steady state system such as with magnetic trapping, or a more chaotic one as favored by bombardment models.

These discoveries raise questions about the origin and physical properties of such dust
grains. Furthermore, because future missions for spectroscopic characterization of exo-Earths could be severely affected by exozodiacal scattering, the systems with the strongest dust emission need to be identified and characterized. In this context, our group is currently involved with the CHARA Array, VLTI, and LBTI to study inner dust disks with high precision.

Concurrent with the exozodiacal survey extension, we are working on a project to resolve the question of the variability of the exozodiacal disks. Knowledge of the variability of the dust provides evidence to support a formation model and places constraints on the various models of inner debris disks. To produce the survey, we take advantage of the long temporal baseline we have in the exozodiacal data. By utilizing the survey results from Absil et al. (2013), we have a record of the fractional amount of exozodiacal dust present for these stars ranging from 2005 to 2011. For the NIR dust variability study, we initially selected 12 bright, spectral type A-K stars with previously detected excesses out of the 42 star exozodiacal survey list from Absil et al. (2013). Since then, our list has been expanded to include the objects presented here. The circumstellar disk flux ratio, $f_{\text{CSE}}$, is compared between the archival data taken with FLUOR by Absil et al. and the $f_{\text{CSE}}$ taken more recently, within the past 3 years, by JouFLU. The upgrade of FLUOR and in particular the development of a new DRS poses a challenge to the validity of comparing the results from the two instruments. However, we have addressed this by re-reducing the original data with the new pipeline to show that the two instruments are consistent and compatible.
7.2 Results

Our data are separated by date, so that plots shown are only for data collected within a few nights. Data from different runs are reduced separately. Table 7.2 gives some details about the observing parameters and conditions.
<table>
<thead>
<tr>
<th>HD</th>
<th>Object</th>
<th>Obs date</th>
<th>Baseline</th>
<th>Number of scans</th>
<th>Notes</th>
</tr>
</thead>
<tbody>
<tr>
<td>9826</td>
<td>υ And</td>
<td>10/15/2013, 10/16/2013</td>
<td>E1-E2</td>
<td>9</td>
<td>~3% (non-physical) excess with a &gt; 3 sigma significance. Bad calibration, only 4 points.</td>
</tr>
<tr>
<td>22484</td>
<td>10 Tau</td>
<td>7/22/2014, 11/10/2014</td>
<td>S1-S2</td>
<td>7</td>
<td></td>
</tr>
<tr>
<td>40136</td>
<td>γ Lep</td>
<td>10/13/2014, 11/10/2014</td>
<td>S1-S2</td>
<td>9</td>
<td></td>
</tr>
<tr>
<td>102647</td>
<td>β Leo</td>
<td>5/1/2015, 5/27/2015, 5/28/2015</td>
<td>S1-S2</td>
<td>20</td>
<td></td>
</tr>
<tr>
<td>131156</td>
<td>ξ Boo</td>
<td>5/1/2015</td>
<td>S1-S2</td>
<td>8</td>
<td></td>
</tr>
<tr>
<td>142091</td>
<td>κ CrB</td>
<td>5/1/2015, 6/18/2015, 6/19/2015, 6/20/2015</td>
<td>S1-S2,E1-E2</td>
<td>17</td>
<td></td>
</tr>
<tr>
<td>142860</td>
<td>γ Ser</td>
<td>5/15/2013, 5/25/2014, 6/16/2015</td>
<td>E1-E2</td>
<td>22</td>
<td>No excess found by Absil et al. (2013), but we find one.</td>
</tr>
<tr>
<td>161868</td>
<td>γ Oph</td>
<td>6/17/2015</td>
<td>S1-S2,E1-E2</td>
<td>8</td>
<td>No excess found by Absil, and none found with E1-E2. Poor calibration.</td>
</tr>
<tr>
<td>172167</td>
<td>α Lyr</td>
<td>7/22/2014, 5/27/2015, 8/14/2015</td>
<td>S1-S2</td>
<td>21</td>
<td></td>
</tr>
<tr>
<td>173667</td>
<td>110 Her</td>
<td>6/16/2015</td>
<td>S1-S2</td>
<td>16</td>
<td>periodic appearance of data indicates companion 110 Her is in WDS Catalog, but min separation is &gt; 20&quot;</td>
</tr>
<tr>
<td>177724</td>
<td>ζ Aql</td>
<td>6/18/2015</td>
<td>E1-E2</td>
<td>9</td>
<td>new object to var list.</td>
</tr>
<tr>
<td>185144</td>
<td>σ Dra</td>
<td>8/15/2015</td>
<td>S1-S2</td>
<td>4</td>
<td>Abris et. al. (2013) detected no excess, we confirm that.</td>
</tr>
<tr>
<td>187642</td>
<td>α Aql</td>
<td>7/23/2014, 6/17/2015</td>
<td>E1-E2,S1-S2</td>
<td>15</td>
<td>minimum and maximum UD models and fast rotator model from Monnier et al. (2007).</td>
</tr>
</tbody>
</table>
7.3 Discussion

Table 7.3: Variability study results, $f_{\text{cse}}$ (%) of both Absil et al. (2013) and this work, along with reduced $\chi^2$ for each, and the difference: $\Delta f_{\text{cse}}(\%) = f_{\text{cse}}^\dagger - f_{\text{cse}}^{\dagger \dagger}$. Results shown for $\alpha$ Cep are from 2015 and use the rapid rotator model from van Belle et al. (2006), results from $\alpha$ Aql use the rapid rotator model from Monnier et al. (2007). $\Delta f_{\text{cse}}$ error is found by adding $f_{\text{cse}}$ errors in quadrature.

<table>
<thead>
<tr>
<th>HD</th>
<th>Object</th>
<th>$f_{\text{cse}}$ (%) $^\dagger$</th>
<th>$\chi^2$ $^\dagger$</th>
<th>$f_{\text{cse}}$ (%) $^{\dagger \dagger}$</th>
<th>$\chi^2$ $^{\dagger \dagger}$</th>
<th>$\Delta f_{\text{cse}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>9826</td>
<td>$\upsilon$ And</td>
<td>0.53 ± 0.17</td>
<td>3.12</td>
<td>3.01 ± 0.48</td>
<td>0.63</td>
<td>2.48 ± 0.50</td>
</tr>
<tr>
<td>22484</td>
<td>10 Tau</td>
<td>1.21 ± 0.11</td>
<td>11.00</td>
<td>-2.75 ± 0.73</td>
<td>0.08</td>
<td>-3.96 ± 0.74</td>
</tr>
<tr>
<td>40136</td>
<td>$\eta$ Lep</td>
<td>0.89 ± 0.21</td>
<td>4.24</td>
<td>0.45 ± 0.48</td>
<td>2.03</td>
<td>-0.44 ± 0.52</td>
</tr>
<tr>
<td>102647</td>
<td>$\beta$ Leo</td>
<td>0.94 ± 0.26</td>
<td>3.62</td>
<td>1.09 ± 0.53</td>
<td>0.90</td>
<td>0.15 ± 0.59</td>
</tr>
<tr>
<td>131156</td>
<td>$\zeta$ Boo</td>
<td>0.74 ± 0.20</td>
<td>3.70</td>
<td>-0.32 ± 0.43</td>
<td>1.40</td>
<td>-1.06 ± 0.47</td>
</tr>
<tr>
<td>142091</td>
<td>$\kappa$ CrB</td>
<td>1.18 ± 0.20</td>
<td>5.90</td>
<td>3.40 ± 0.53</td>
<td>1.85</td>
<td>2.22 ± 0.56</td>
</tr>
<tr>
<td>142860</td>
<td>$\gamma$ Ser</td>
<td>-0.06 ± 0.27</td>
<td>-0.22</td>
<td>1.63 ± 0.54</td>
<td>0.70</td>
<td>1.69 ± 0.60</td>
</tr>
<tr>
<td>161868</td>
<td>$\gamma$ Oph</td>
<td>0.25 ± 0.48</td>
<td>0.52</td>
<td>-0.61 ± 0.81</td>
<td>16.48</td>
<td>-0.86 ± 0.94</td>
</tr>
<tr>
<td>172167</td>
<td>$\alpha$ Lyr</td>
<td>1.26 ± 0.27</td>
<td>4.67</td>
<td>1.82 ± 0.78</td>
<td>16.99</td>
<td>0.56 ± 0.82</td>
</tr>
<tr>
<td>173667</td>
<td>110 Her</td>
<td>0.94 ± 0.25</td>
<td>3.76</td>
<td>-0.47 ± 0.30</td>
<td>14.20</td>
<td>-1.41 ± 0.39</td>
</tr>
<tr>
<td>177724</td>
<td>$\zeta$ Aql</td>
<td>1.69 ± 0.27</td>
<td>6.26</td>
<td>1.23 ± 0.38</td>
<td>1.32</td>
<td>-0.46 ± 0.47</td>
</tr>
<tr>
<td>185144</td>
<td>$\sigma$ Dra</td>
<td>0.15 ± 0.17</td>
<td>0.88</td>
<td>-1.11 ± 1.00</td>
<td>1.88</td>
<td>-1.26 ± 1.01</td>
</tr>
<tr>
<td>187642</td>
<td>$\alpha$ Aql</td>
<td>3.07 ± 0.24</td>
<td>12.90</td>
<td>6.12 ± 0.74</td>
<td>2.76</td>
<td>3.05 ± 0.78</td>
</tr>
<tr>
<td>203280</td>
<td>$\alpha$ Cep</td>
<td>0.87 ± 0.18</td>
<td>4.70</td>
<td>-0.14 ± 0.78</td>
<td>0.74</td>
<td>-1.01 ± 0.80</td>
</tr>
</tbody>
</table>

$^\dagger$ Absil et al. (2013)

$^{\dagger \dagger}$ this work

Figure 7.1 shows the results from our exozodical dust variability study along with the original results from Absil et al. (2013) in terms of $f_{\text{cse}}$, while Figure 7.2 shows the change in $f_{\text{cse}}$.

Table 7.3 lists the overall results of this study in comparison to the results from Absil et al. (2013). In order to investigate any claim of variability, we must first verify that any changes cannot be explained as result of the concurrent changes to the instrument and software. To
verify that the hardware changes are not a cause, the stellar diameters measured with JouFLU were compared with those found with the CLASSIC beam combiner. This was shown explicitly in subsection 4.12.2, and a fit to a stellar model is also an integral part of the DRS for all data collected. To verify that the DRS can satisfactorily replicate the Absil et al. (2013) results, we have re-analyzed some of the original data collected by Olivier Absil for Absil et al. (2013). Our DRS yields excess measurements consistent, within errors, of his findings.

There are some differences between the data quality criteria used for the original work and our follow-up study. The original FLUOR DRS reduced the I1 and I2 channels separately and only did the difference at the end. The new JouFLU DRS uses the difference for every scan and treats it as an extra channel.

Data quality starts at the individual fringe-scan level. We verify that the fringe packet is not too close to the edge of the scan window, i.e., there should be at least $\sim 3$ fringe envelopes, central fringe packet+side lobes, within the scan window. Then we check that the power of the fringe power spectrum is sufficiently greater than the noise, which can be estimated from the power spectrum away from the fringe peak location. The quantity estimating this, $\frac{f_{\text{on}} - f_{\text{off}}}{f_{\text{on}}}$, should be greater than 0.5, where $f_{\text{on}}$ and $f_{\text{off}}$ refer to the on-fringe and off-fringe peak respectively. If there are less than 50 scans remaining, we do not include the point in the analysis due to lack of statistics.

After the raw visibilities are computed, using the median estimator, and subsequently calibrated, we perform the following checks to assess the quality of calibrated points. First, we check that the variability of the transfer function is not too large. The final error bar takes
into account the transfer function variability as a systematic error, but we have empirically shown that it is good practice, but not critical, to reject calibrated points whose systematic error is $\sim 30$ times greater than the statistical error determined by using the data set itself to provide the statistical distribution of the data, also known as data bootstrapping. Most importantly, we then check for the agreement of the calibrated visibility between the two interferometric channels via a type of $\chi^2$ analysis described in Perrin (2003), and reject calibrated points whose $\chi^2$ is greater than 3.0.
Figure 7.1: JouFLU findings (red, filled circles) compared to Absil et al. (2013) published exozodi measurements (cyan, filled circles). Also shown are original data from Absil et al. (2013) that have been re-reduced with the JouFLU data reduction pipeline (green, filled squares, offset by +0.25 in the x-axis). Results from the new and old reduction are consistent, at least for the four stars for which we have done the check. Stars are listed by their HD number. For Altair, HD 187642, results are shown for a rapid rotator stellar model fit, from Monnier et al. (2007), for both the new and original data (filled red circle and green square, respectively). In addition, fits for the minimum and maximum UD diameter fit for Altair are shown in downward and upward triangles, respectively. For α Cep, HD 203280, the red, filled circle represents the 2014 data, while the blue, filled diamond is the result from the 2015 data, results from both years rely upon the rapid rotator model from van Belle et al. (2006). There is evidence for time variability for some stars, with one caveat, the baseline lengths and orientations are different between epochs. Due to technical issues, S1-S2 was not always available for the new data, so the E1-E2 baseline was used instead.
Figure 7.2: JouFLU findings (red, filled circles) for the difference in circumstellar flux percentage: $\Delta f_{\text{cse}}(\%) = f^+_{\text{cse}} - f^+_{\text{cse}}$. Results shown for $\alpha$ Cep are from 2015 and use the rapid rotator model from van Belle et al. (2006), results from $\alpha$ Aql shown use the rapid rotator model from Monnier et al. (2007). The $\Delta f_{\text{cse}}$ errors are found by adding $f_{\text{cse}}$ errors in quadrature. HD22484 is included as an example of the problem with analysis of undersampled, low-quality data, not necessarily changes in the star. As a result, no change should be concluded for HD22484.

† Absil et al. (2013)
‡ this work
7.3.1 2013 JouFLU Observations

S2 was out of use for some of the 2013 observing runs, due to the AO upgrades at the telescope. Much of the data that was collected with JouFLU during this period was used as for engineering and not usable for science.
7.3.1.1 υ And

Figure 7.3: HD9826
7.3.2 2014 JouFLU Observations

Most of 2014 data used only one interferometric channel, due to JouFLU fiber problems. Some data had both channels but one was significantly weaker, in this case a weighted mean was used instead of the difference signal. Also in 2014, S1 had a telescope drive motor failure, preventing use of the S1-S2 baseline for the remaining observing runs from May to August.
7.3.2.1 10 Tau

(a) HD22484 baseline

(b) HD22484 hour angle

(c) HD22484 phase angle

Figure 7.4: HD22484
Figure 7.5: HD40136
7.3.2.3 2014 α Cep

Figure 7.6: 2014 HD203280
7.3.2.4 2014 rapid rotator $\alpha$ Cep

![Graphs showing baseline, hour angle, and phase angle for HD203280.](image)

(a) HD203280 baseline

(b) HD203280 hour angle

(c) HD203280 phase angle

Figure 7.7: 2014 HD203280 rapid rotator model
7.3.3 2015 JouFLU Observations

For 2015, both JouFLU interferometric channels worked properly, unbalanced fiber flux issues with JouFLU were resolved, and both short baselines, S1-S2 & E1-E2, were available. As a result, the 2015 data is the best data collected with JouFLU to date.
7.3.3.1 β Leo

Figure 7.8: HD102647
7.3.3.2 ζ Boo

Figure 7.9: HD131156
Figure 7.10: HD142091
7.3.3.4 γ Ser

Figure 7.11: HD142860
7.3.3.5 γ Oph

Figure 7.12: HD161868
7.3.3.6 α Lyr

Figure 7.13: HD172167
Figure 7.14: HD173667
7.3.3.8 ζ Aql

Figure 7.15: HD177724
7.3.3.9 α Dra

Figure 7.16: HD185144
7.3.3.10 minimum UD ∆ Aql

Figure 7.17: HD187642 - minimum UD model
Figure 7.18: HD187642 - maximum UD model
7.3.3.12 rapid rotator model $\alpha$ Aql

![Graphs showing data for HD187642 with baseline, hour angle, and phase angle.]

Figure 7.19: HD187642 - rapid rotator model
Figure 7.20: 2015 HD203280
7.3.3.14 2015 rapid rotator model $\alpha$ Cep

Figure 7.21: 2015 HD203280 rapid rotator model
7.4 Absil Data Re-Reduction

The following plots are from original FLUOR data from Absil et al. (2013), but reduced using the new JouFLU DRS. These plots use the shutter sequence as the background estimation method, and the median, actually $\sqrt{\text{median}(V^2)}$, as the visibility estimator. Also applied were our own data quality tests, devised by Paul Nuñez, and mentioned above.

Note: In general the Absil et al. (2013) results are reproducible. However, the choice of the background estimation method and the visibility estimator, may significantly affect the final result for some objects, resulting in an excess or a non-excess. In these cases, we have tried to be as conservative as possible.
7.4.1 2008 \( \nu \) And

Figure 7.22: HD9826 - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
Figure 7.23: HD142091 - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
7.4.3 2011 γ Ser

Figure 7.24: HD142860 - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
7.4.4 2005 α Lyr

Figure 7.25: HD172167 - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
7.4.5 2011 110 Her

Figure 7.26: HD173667 - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
Figure 7.27: HD187642 - minimum UD model - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
7.4.7 2011 maximum UD α Aql

Figure 7.28: HD187642 - maximum UD model - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
7.4.8 2011 rapid rotator model $\alpha$ Aql

![Graph](image)

Figure 7.29: HD187642 - rapid rotator model - original data from Absil et al. (2013) reduced with modern JouFLU DRS.
7.5 Discussion of individual targets

7.5.1 **υ And (HD 9826).**

This is one of the sources for which we see the most variability. This star now shows an excess detection of $6.3\sigma$ with a $f_{\text{cse}}$ of $3.01 \pm 0.48\%$, up from the $0.53 \pm 0.17\%$ in Absil et al. (2013) where this was considered a non-detection. They did report that the small excess they detected was robust and showed good data quality. We propose that this is an excess star that shows variability in the disk/star flux ratio. Re-analysis of the original FLUOR data, processed with the new JouFLU DRS, shows an excess of $0.23 \pm 0.28\%$. This is consistent with the published Absil et al. (2013) results and lends validity to the JouFLU DRS. However, none of the points analyzed were published in Absil et al. (2013). The points are mostly between hour angle 0 and 0.6. Some of the original data seems to be missing and unaccounted for, so this reduction is based on an incomplete data set. Also found were some old long-baseline data that matches the stellar model quite well.

7.5.2 **10 Tau (HD 22484).**

HD 22484 was one of the Absil group’s most significant excess detections, at $11\sigma$. Our result of a large negative excess can only be attributed to very poor calibration and collecting only four data points. At the time of observation only two calibrators were available due to a worsened magnitude limit for JouFLU during 2014. It is possible that one of the remaining two calibrators failed to meet the calibrator criteria. From this, it is suggested that this result not
be given much weight until it can be re-observed.

7.5.3 $\eta$ Lep (HD 40136).

We find a disk/star flux ratio of $0.45 \pm 0.48\%$ for this star. This has a confidence level of $0.9\sigma$. From this, we cannot confidently say this star has an excess. However, our results are consistent, within $1\sigma$, of those found by Absil et al. who reported this as an excess. This demonstrates the differences that can arise from the number of data points, how the individual scans are filtered, and how the measurement errors are handled.

7.5.4 $\beta$ Leo (HD 102647).

This star was shown to have an excess by Akeson et al. (2009) using FLUOR data from 2006. Absil et al. (2013) expanded this data set using observations from 2009 and found a smaller excess of $0.94 \pm 0.26\%$. The group was also very thorough in constraining the possibility that this excess results from a faint companion, by using AO-assisted aperture masking at the Keck-II telescope to confirm that this visibility deficit was from an extended source. Our results confirm this excess, with a ratio of $f_{\text{cse}} = 1.09 \pm 0.53\%$ making this object a very strong argument for the excess being caused by an extended rather than point source. It is worthy of note that the Absil group reported a poor model fit, $\chi^2_{\nu} = 5.5$, whereas we find a relatively good fit of $\chi^2_{\nu} = 0.90$. 
7.5.5 ξ Boo (HD 131156).

The Absil 2013 paper reported their detection of an excess on this star as “pessimistic”. We find no excess around this star, \( f_{\text{cse}} = -0.32 \pm 0.43\% \). This suggests that either the initial detection was a false positive, or the system is displaying some variability.

7.5.6 κ CrB (HD 142091).

For this star we find an excess ratio of 3.40 ± 0.53\%, significantly higher than the 1.18 ± 0.20\% found by Absil et al. (2013). Re-analysis of the original data is consistent with the Absil et al. (2013) results, with an excess of 1.55 ± 0.27\%. This star is a sub-giant, with luminosity class III-IV. It is a so-called ‘retired’ A-star; it began as an A-star and is now nearing the red giant phase. Bonsor et al. (2013b) reports on the spatially resolved debris disk in the system and the minimum 2.1 Jupiter mass planet and likely secondary planetary companion. This system is highly interesting for the study of dust, its variability, its production/destruction mechanisms, and the relationship to exoplanets.

7.5.7 γ Ser (HD 142680).

This star was included in these results despite having only three data points. However, these points pass all of the DRS quality checks. This star previously showed no excess but now shows a strong excess. The flux ratio was measured by Absil et al. to be \(-0.06 \pm 0.27\%\), but in our results it shows an excess of 1.63 ± 0.54\%. Such a difference could be the result of true variability or it could be due to disk structure and the use of a different baseline and
orientation, e.g. E1-E2 instead of S1-S2. Only additional observations may resolve this ambiguity.

The original data from 2011 on HD 142860 data were re-analyzed and found consistent with a non-excess, $f_{\text{cse}} = -0.54 \pm 0.42\%$. Note: Depending on the background subtraction method used, a $3\sigma$ excess of $1.7 \pm 0.52\%$ is possible on the 2011 data, but with a large $\chi^2$. The recent 2015 data seem to be more robust against DRS parameters, always giving a $>3\sigma$ excess. The new data always have a poor fit, but it may be that the dust distribution is asymmetric.

7.5.8 $\gamma$ Oph (HD 161868).

We find no excess for this star, nor was one found by Absil et al. (2013), $f_{\text{cse}} = 0.25 \pm 0.48\%$ vs. $-0.61 \pm 0.81\%$ respectively. The poor model fit and high $\chi^2$ is likely due to suspected poor calibration.

7.5.9 $\alpha$ Lyr (HD 172167).

Vega is an extremely well-studied star, famous for many reasons. The presence of a circumstellar NIR excess was presented by Absil et al. (2006). This has since been confirmed and attributed to extended emission. It remains to be seen if the emission is from a uniform disk, zodiacal disc, or a narrow ring. Our results find an excess of $1.82 \pm 0.78\%$, consistent with the previous findings. However, due to the large errors on this measurement and poor model fit we refrain from drawing conclusions about the presence of dust. These large errors are attributed to difficulties arising from the brightness of Vega and calibration of the data. For
most of our observations, including those of Vega, we have found a benefit in favoring proximity over similarity in magnitude for calibrator selection. We chose calibrators that were significantly dimmer than Vega, compared to the calibrators chosen by Absil et al. The DRS reports a very large variability of the transfer function, compared to the statistical error of the raw visibilities. However, the $\chi^2$ between the two channels seems be acceptable. Variations in the photometric channels cause data scans to be rejected and are correlated with larger error bars, this effect may be very pronounced with extremely bright targets, creating a selection bias for poor observing conditions. Another possibility is that Vega is an object often observed when the weather or seeing prevents the observation of fainter targets. For our results, less than 80 scans were used in the analysis of each point. We believe our excess amount to be reliable, however we have little confidence in the reported error bars. This effect is currently being investigated.

7.5.10 110 Her (HD 173667).

HD173667 is a FLUOR 2013 paper excess star, and E1-E2 measurements show that this is likely a binary. We find no excess and an apparent periodicity with respect to the hour angle. 110 Her is in the WDS with four components but the minimum separation given is greater than $20''$. It is possible observations occurred at point in the orbit where the separation was within JouFLU’s FoV or perhaps there is another component. This eliminates it as a candidate for future exozodiacal dust observations.

Data from 2011, re-analyzed with the JouFLU DRS, do not show an excess and the result is
somewhat compatible with Absil et al. (2013). However, the data analyzed are not published and correspond to $0 < \text{HA} < 2$, a region which is missing in Olivier’s plot. The data look fine according to the observing logs and the DRS data quality checks.

7.5.11 $\zeta$ Aql (HD 177724).

We find an excess of $1.23 \pm 0.38\%$ for this star, consistent with Absil et al.’s finding of $1.69 \pm 0.31\%$. Absil et al. (2013) rules out the most likely companions for this object based on MIRC, Canada France Hawaii Telescope (CFHT), and KIN observations.

7.5.12 $\sigma$ Dra (HD 185144).

Only four data points were obtained on this object. However, it is a new object to the exozodiacal variability target list, and provides an important null result. Absil et al. (2013) detected no excess and we have confirmed that.

7.5.13 $\alpha$ Aql (HD 187642).

Altair was observed in 2015. The data were reduced using three models: a minimum UD model, a maximum UD model, and a rapid rotator model. The rapid rotator model used was from Monnier et al. (2007) and makes a difference in the model fit as Altair is somewhat resolved with the S1-S2 baseline. The excesses found for each model were: $6.63 \pm 0.57\%$, $4.639 \pm 0.60\%$, and $6.12 \pm 0.74\%$. The $\chi^2$ for each model were found to be: 4.5, 4.4, and 2.8. Compare this with the rapid rotator fit from Absil et al. (2013), which also used Monnier et al.
(2007), of 3.07 ± 0.24% and a $\chi^2 = 12.9$. This suggests an increase in extended emission for \( \alpha \) Aql.

The 2011 and 2015 results are compatible with each other and the $\chi^2$s are lower. The quality of the 2011 data is better, and almost all the points passed data quality checks. The available 2011 data set for this object was found to be incomplete and as a result only half of the original data was used for this reduction. The results are still found to be consistent. However, this may account for any minor discrepancies.

7.5.14 \( \alpha \) Cep (HD 203280).

This star was observed with E1-E2 in 2014 and S1-S2 in 2015. The data were reduced with and without modeling the star as a rapid rotator. The rapid rotator model that was applied came from van Belle et al. (2006). From our data, we find no significant excess. The Absil et al. (2013) results for \( \alpha \) Cep used a rapid rotator model from Zhao et al. (2009) and found an excess of 0.87 ± 0.18%. 
“I can live with doubt and uncertainty and not knowing. I think it is much more interesting to live not knowing than to have answers that might be wrong.”

— Richard Feynman
As powerful as optical interferometric techniques using the CHARA Array and JouFLU or the VLTI are, it is unlikely that they can fully characterize the exozodiacal phenomenon. Additional work using other instruments and observational techniques is highly desirable in order to confirm the existing results and constrain theoretical models. As such, the author is leading a pilot project that utilizes the SpeX spectrograph at NASA’s IRTF. In addition to this, my collaborator, Bertrand Mennesson, has been leading an observing campaign at the Palomar Fiber Nuller (PFN), working to constrain the physical domain of the exozodiacal dust reservoir. Finally, plans have been made to incorporate other beam combiners at the CHARA Array in order to strengthen the exozodiacal data and to rule out possible binary contamination. Suggestions and requirements for future observations are also discussed.
8.1 Exozodiacal Disk Spectrophotometric Survey

This project is an exploratory program lead by the author to utilize the newly upgraded 0.7-5.3 micron medium-resolution spectrograph and imager (SpeX) at NASA's IRTF for spectrophotometric warm/hot exozodiacal disk detections. Recent success with ground-based cold disk detections indicate warm disk detections should be possible (Lisse et al. 2013). This project has been awarded four observing runs to date, the data from which are being used to develop an independent DRS, specialized for detecting thermal excesses.

8.1.1 Introduction

High SNR relative spectrophotometric calibration of SpeX from 1-5 μm is used to search for a photometric excess corresponding to warm/hot dust around stars that show an interferometric excess in the NIR. These results place new constraints on the excesses detected by CHARA and lead to robust criteria for spectroscopic identification of stars that harbor inner debris disks. Sets of stars with known hot exozodiacal disks and those known to lack such disks, based upon recent interferometric observation, are compared.

Infrared excesses indicate exozodiacal dust disks and trace planetesimal belts and planetary system architecture. FIR observations show that roughly 15% of G-K stars show the presence of cold dust (Carpenter et al. 2009). Among A stars, this rate goes up to 30% or more (Su et al. 2006). Eiroa et al. (2013) found an incidence rate of 20% around solar-type FGK stars. The presence of a warm or hot inner disk material is not always concurrent with
the relatively common cold disk. Absil et al. (2013) found that 80% of A stars that show excess emission from $<3$ AU lack evidence of a cold disk. However, the warm dust is rarely detected in photometric surveys due to the precision and dynamic range required.

Interferometric surveys have shown that hot dust is common in main sequence star systems. However, these surveys are observing time intensive and only possible from the CHARA Array in the Northern hemisphere (Absil et al. 2013) and the VLTI in the Southern hemisphere (Ertel et al. 2014). The task of extending the survey to larger samples would best be suited to single-aperture spectrophotometry. High SNR observations that show an excess at longer wavelengths relative to the stellar model are indicative of the presence of hot dust. Thus, high SNR relative spectrophotometric calibration data from SpeX from 1-5 microns are used to search for a photometric excess corresponding to 800-1500 K dust around stars that show an interferometric excess in the NIR. In addition to confirming interferometrically derived excesses, our goal is to develop a technique by intercomparison of the two data sets for identifying excesses based on spectroscopic IR data.

This work complements the CHARA/JouFLU disks survey and improves ground-based spectrographic/photometric infrared disk detection capability. Ultimately, we seek to place new constraints on the excesses detected by CHARA and possibly lead to robust criteria for spectrophotometric identification of stars that harbor inner debris disks. Our interferometric survey may make possible the identification of statistically significant observational features that can be used as spectroscopic debris disk markers. Exploring connections between the presence of debris, the host star properties, e.g. age, rotation, metallicity, and the
replenishment of dust, may provide clues to planetary formation.

Together with the temporal monitoring of the detected NIR excesses, a complementary program led by the author, this is currently our only way to get a better understanding of this phenomenon and to constrain the dust evolution models that we are developing within two different working groups.

Figure 8.1 shows the blackbody curve for a typical A0V star, also shown is the corresponding Kurucz atmosphere model for the A0V and for the star, ξ Boo. On the same figure is a blackbody curve for a face-on disk and our first data taken with IRTF on ξ Boo. Figure 8.2 shows the region of the ξ Boo data in detail, while Figure 8.3 shows the corresponding expected and measured SNR. It is this initial proof-of-concept that led to the development of the more refined method. This initial ξ Boo data was processed solely with the normal IRTF DRS, getting accurate measures of exozodiacal excesses with this technique requires modifications to the data reduction process, described below.
Figure 8.1: Model luminosity from an A0V star+a face-on 3AU radius 1500K dust disk, at our CHARA detection limit of \( \approx 1\% \) stellar flux at 2.2 \( \mu m \).

Figure 8.2: Model luminosity from an A0V star+a face-on 3AU radius 1500K dust disk, at our CHARA detection limit of \( \approx 1\% \) stellar flux at 2.2 \( \mu m \). Relative difference in the slope of this region identifies and constrains the presence of dust.
Figure 8.3: Model SpeX SNR for an A0V star with 1500K disk, SXD+LXD mode, 0.8 arcsecond seeing, 0.4s minimum frame integration, 300s total integration irtfweb.ifas.hawaii.edu/cgi-bin/spex/spex_calc2.cgi scaled to the expected maximum SNR of 300. Typically the SNR for the disk is a factor of 100 less than that of the star at 2-5 µm. The cumulative SNR of the dust normalized to that of the star shows that over the full spectrum range our sensitivity to dust reaches SNR 100.
8.1.1.1 SpeX

**SpeX** is a medium resolution, cross-dispersed spectrograph at NASA’s **IRTF** (Rayner et al. 2003). It has a resolving power of $R \approx 2000$. The detector is a 2048x2048 HgCdTe HAWAII-2RG array with a pixel size of 18 $\mu$m. For this work, two modes were utilized, Short Wavelength Cross-dispersed Mode (SXD) and Long Wavelength Cross-dispersed Mode (LXD). These provide a wavelength coverage of 0.69 - 2.56 $\mu$m and 1.94 - 5.36 $\mu$m, respectively. The slit dimension chosen for this project was 0.3x15″. SpeX instrument Principal Investigator (PI) and co-investigator on this project, John Rayner, assists with methods to avoid saturation of the array as well as observing strategies.

8.1.2 Method

We collect relative spectrophotometry from 0.69 - 5.36 $\mu$m for a selection of our interferometrically observed excess stars, and compare it to control stars of the similar spectral type for differences in the SED slope across the NIR. Table 8.1 lists the excess stars, while Table 8.2 lists the controls. Emission < 1.4$\mu$m is dominated by the star, whereas longer wavelengths display additional flux contribution from the warm dust disk. Focusing on the relative slope of the spectrum rather than on its absolute value improves our sensitivity to warm dust.

Due to the brightness of our interferometric targets, $m_K < 4.5$, observations with IRTF have a very high SNR. Integration is optimized to obtain SNR=100 on the dust disk, when the flux of
dust is ≈1% stellar flux. Star SNR is limited by systematic array pattern noise for bright targets, and we can expect a SNR of > 300 per resolution element for the star+disk. After removal of the star SED, SNR of ≈3 should remain for the disk. Integrating over the full spectral range yields a total $\text{SNR}_{\text{dust}} * \sqrt{N_{\text{elements}}} \approx 150$ for the detection of the disk. A bracket consists of an observation of the science target in each mode, SXD and LXD, one of the control star in each mode, and one of a telluric standard star in each mode.

The short-wavelength spectra can be normalized, and fit to the stellar SED using stellar models. Excess at long wavelengths that does not fit is the dust disk spectrum. Ancillary photometry from the literature can be used to normalize the stellar part of the SED. This technique has been demonstrated by Lisse et al. (2013). The photosphere-removed spectra can then be compared to separate MIR spectra or MIR spatially resolved excess spectra when available (LBTI). A MIR excess provides constraints on the temperature of the disk, size of the dust grains in the <1 - 100 $\mu$m range, composition of the dust, mass of the disk, and albedo of the disk. These components can then be used as constraints in modeling the dust disk.

Observing runs consisting of three half nights and one full night have been fulfilled as a pilot to this program. This project should enable us to confirm the level of the 2 $\mu$m excess seen in the CHARA data, determine the temperature of the hottest dust observed, determine if there is any evidence for multiple temperature components, and constrain the solid angle of the emitting area. Once the known excess stars are observed and a clear method has been devised to identify excesses based on IRTF spectra reliably, a follow-up effort could look for
previously undetected debris disk stars.
Table 8.1: IRTF project target list. 13 Stars with confirmed exozodi disks (Absil et al. 2013). The complete target list consists of a sample of 24 main sequence stars with $K \leq 4.5$. Among these stars, 13 have a known NIR excess.

<table>
<thead>
<tr>
<th>HD</th>
<th>Object name</th>
<th>Coordinates</th>
<th>Kmag</th>
<th>Integration (SXD,LXD)</th>
<th>Co-adds/Cycles</th>
<th>Warm excess $f_{CSE}$ (%)$^1$</th>
<th>Cold Excess</th>
</tr>
</thead>
<tbody>
<tr>
<td>10700</td>
<td>$\tau$ Cet</td>
<td>$01^h 44^m, -15^\circ 56$</td>
<td>1.68</td>
<td>1,2</td>
<td>10/10,10/10</td>
<td>0.98</td>
<td>5.4 far</td>
</tr>
<tr>
<td>22484</td>
<td>10 Tau</td>
<td>$03^h 36^m, +00^\circ 4$</td>
<td>2.90</td>
<td>-</td>
<td>-</td>
<td>1.21</td>
<td>11.0 far</td>
</tr>
<tr>
<td>40136</td>
<td>$\eta$ Lep</td>
<td>$05^h 56^m, -14^\circ 10$</td>
<td>2.91</td>
<td>1.5,2</td>
<td>1/5,10/10</td>
<td>0.89</td>
<td>4.3 mid, far</td>
</tr>
<tr>
<td>56537</td>
<td>$\lambda$ Gem</td>
<td>$07^h 18^m, +16^\circ 32$</td>
<td>3.54</td>
<td>-</td>
<td>-</td>
<td>0.74</td>
<td>4.3 –</td>
</tr>
<tr>
<td>102647</td>
<td>$\beta$ Leo</td>
<td>$11^h 49^m, +14^\circ 3$</td>
<td>1.93</td>
<td>0.5,1</td>
<td>1/5,4/10</td>
<td>0.94</td>
<td>3.6 mid, far</td>
</tr>
<tr>
<td>131156</td>
<td>$\xi$ Boo</td>
<td>$14^h 51^m, +19^\circ 06$</td>
<td>2.96</td>
<td>3,5</td>
<td>4/3,10/3</td>
<td>0.74</td>
<td>3.7 –</td>
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<td>142091</td>
<td>$\kappa$ CrB</td>
<td>$15^h 51^m, +35^\circ 39$</td>
<td>2.49</td>
<td>-</td>
<td>-</td>
<td>1.18</td>
<td>5.9 far</td>
</tr>
<tr>
<td>172167</td>
<td>$\alpha$ Lyr</td>
<td>$18^h 36^m, +38^\circ 47$</td>
<td>0.00</td>
<td>-</td>
<td>-</td>
<td>1.26</td>
<td>4.7 mid, far</td>
</tr>
<tr>
<td>173667</td>
<td>110 her</td>
<td>$18^h 45^m, +20^\circ 32$</td>
<td>3.06</td>
<td>5,3</td>
<td>4/3,10/3</td>
<td>0.94</td>
<td>3.8 far</td>
</tr>
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<td>177724</td>
<td>$\zeta$ Aql</td>
<td>$19^h 05^m, +13^\circ 51$</td>
<td>2.91</td>
<td>4,3</td>
<td>4/3,10/5</td>
<td>1.69</td>
<td>6.3 –</td>
</tr>
<tr>
<td>187642</td>
<td>$\alpha$ Aql</td>
<td>$19^h 50^m, +08^\circ 52$</td>
<td>0.24</td>
<td>-</td>
<td>-</td>
<td>3.07</td>
<td>12.9 –</td>
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<tr>
<td>203280</td>
<td>$\alpha$ Cep</td>
<td>$21^h 18^m, +62^\circ 35$</td>
<td>1.85</td>
<td>-</td>
<td>-</td>
<td>0.87</td>
<td>4.7 –</td>
</tr>
<tr>
<td>211336</td>
<td>$\epsilon$ Cep</td>
<td>$22^h 15^m, +17^\circ 02$</td>
<td>3.54</td>
<td>-</td>
<td>-</td>
<td>3.25</td>
<td>4.7 –</td>
</tr>
</tbody>
</table>

$^1$ Estimated K-band disc/star flux ratio ($f_{CSE}$)

$^2$ Significance of measured disc/star flux ratio ($\chi_f = f_{CSE}/\sigma_f$)
Table 8.2: IRTF project control list. 6 Observed stars out of 27 with no interferometric exozodi disk detection (Absil et al. 2013). These serve as a control sample to compare with the hot exozodi sample.

<table>
<thead>
<tr>
<th>HD</th>
<th>Object name</th>
<th>Coordinates</th>
<th>Kmag</th>
<th>Time (SXD,LXD)</th>
<th>Co-adds/Cycles</th>
<th>Comments</th>
</tr>
</thead>
<tbody>
<tr>
<td>3651</td>
<td>54 Psc</td>
<td>00\textsuperscript{h} 39\textsuperscript{m}, +21\textsuperscript{°} 15</td>
<td>3.999</td>
<td>25,2</td>
<td>1/5,10/10</td>
<td></td>
</tr>
<tr>
<td>16895</td>
<td>θ Per</td>
<td>02\textsuperscript{h} 44\textsuperscript{m}, +49\textsuperscript{°} 13</td>
<td>2.697</td>
<td>25,2;13,2</td>
<td>1/5,10/10;1/5,10/5</td>
<td>repeated due to clouds</td>
</tr>
<tr>
<td>71155</td>
<td>30 Mon</td>
<td>08\textsuperscript{h} 25\textsuperscript{m}, −03\textsuperscript{°} 54</td>
<td>4.079</td>
<td>4,5,3</td>
<td>1/10,4/10</td>
<td>mid &amp; far but no nir excess</td>
</tr>
<tr>
<td>97603</td>
<td>δ Leo</td>
<td>11\textsuperscript{h} 14\textsuperscript{m}, +20\textsuperscript{°} 31</td>
<td>2.144</td>
<td>0.5,0.5</td>
<td>1/5,5/10</td>
<td></td>
</tr>
<tr>
<td>109085</td>
<td>η Crv</td>
<td>12\textsuperscript{h} 32\textsuperscript{m}, −16\textsuperscript{°} 11</td>
<td>3.372</td>
<td>3.8,2.1</td>
<td>1/5,5/10</td>
<td></td>
</tr>
</tbody>
</table>
The conceptual technique we developed is:

- Record spectra using the upgraded SpeX (Rayner et al. 2003) instrument from 0.7 – 5.3 µm on both known excess stars and control stars that exhibit no signs of excess.

- Using the short wavelength end of the IRTF spectrum, λ_{min} ≈ 1µm to λ_{cutoff}, a model (Kurucz, Nextgen, or MARCS) is fit to the data.

- Discrete models are interpolated in T_{eff} and log g space.

- Given the observed spectra S_{obs}(λ), or a rebinned lower resolution version, its measurement uncertainty n(λ), and the model spectrum S_{mod}(λ), one can compute, outside of strong telluric absorption bands, a mean offset significance between the model and the data: mean from λ_{min} to λ_{cutoff} of [(S_{obs}(λ) − S_{mod}(λ))/n(λ)]

- Using this derived best fit photospheric model, one can then compute a similar mean offset over the long wavelengths between λ_{cutoff} and λ_{max} (≈ 5µm).

- There are at least three parameters: λ_{cutoff}, a scaling factor for the photosphere, and its slope. The latter are dependent on log g, T_{eff}.

Relative spectrophotometry from the SXD mode range, 0.69 - 2.56 µm, is used to look for an excess in relative SED slope compared to the stellar model. A goal is to confirm the results found by high-precision interferometry with results found by spectrophotometry. Objects with known mid- and/or far-IR excesses are additionally analyzed using data from the LXD mode that continues up to 5.3 µm wavelength.
While the proposal, conceptual development, data collection, and initial data reduction and
evaluation was performed by the author, subsequent development of a DRS pipeline to deal
with the SpeX data for the purposes of this project was the subject of the Master’s thesis of
Raphaela Wagner, with the assistance of Michael Meyer, Bertrand Mennesson, and the
author. Here is presented a brief example of the technique she advanced; for a full discourse
please see Wagner (2015). This work is likely to be continued by the author and Paul Nuñez.

Analyzing data from a cross-dispersed spectrograph in this manner faces many challenges.
The waveband range is broken up into multiple bands across the detector array. These bands
are also contaminated by telluric features. These two effects make constructing a continuous
spectrum and deriving its slope difficult, but not impossible. The method we have developed
from the previously mentioned technique is as follows:

1. Each observation generates 6 to 20 individual spectra. The telluric regions are removed
   and each order is checked by its slope for inconsistencies resulting from any flux
   imbalance between bands. The regions of the cross-dispersed spectrum near the
   edges of the detector typically exhibit poor SNR and are cropped. The result is a
   spectrum with gaps. To reduce instrumental errors the entire spectral range is divided
   into 8 different wavelength bands, or anchor bands. In every anchor band, the spectra
   are scaled to the median value for the corresponding waveband of the median
   spectrum. (See Figure 8.4 vs. Figure 8.5)

2. In order to address flux uncertainties, a boot-strapping algorithm was created to
   produce new spectra based upon the median spectrum found above with the addition of
a Gaussian error term. We are left with a distribution of 20 new spectra to which a
stellar model can be fit.

3. For each boot-strap generated spectrum, $S_{\text{boot}}$, a best-fit Kurucz model atmosphere,
   $K_{\text{best}}$, is determined by fitting the photosphere below $\lambda_{\text{cutoff}}$ to Kurucz models
   interpolated in $T_{\text{eff}}$ and log g space, with steps at every 50 K and 0.1 cm/s$^2$.

4. The relative excess $E_{\text{rel}}$ is determined for each boot-strapped spectrum. According to
   
   $E_{\text{rel},i} = \frac{S_{\text{boot},i}(\lambda) - K_{\text{best},i}(\lambda)}{K_{\text{best},i}(\lambda)}$.

5. From N relative excess fluxes, an averaged dust spectrum with corresponding standard
   deviation is computed.

6. A final excess quantity $E_{\text{rel}} - \delta E_{\text{rel}}$ is determined by computing a weighted average of
   $< E_{\text{rel}}(\lambda) >$ from 2 - 2.3 $\mu$m. The error takes into account correlation between different
   wavelength channels.

For more detail on any of these steps, the reader is referred to Wagner (2015).
Figure 8.4: Ten spectra from $\eta$ Crv prior to anchoring are shown. This shows the variations in flux that occur due to atmospheric and instrumental conditions. The gaps are due to poor transmission windows in the Earth’s atmosphere. (image credit: Wagner 2015)

Figure 8.5: Here are the same spectra from $\eta$ Crv, but now the different orders have been scaled to match at the wavelength anchor points. The vertical dashed lines border each separate order. (image credit: Wagner 2015)
8.1.3 Results

For seven out of the ten stars analyzed, excess quantities consistent with interferometric results were found. Significant excesses were confirmed for three out of five objects: ζ Boo, ζ Aql, and τ Cet (see the examples: Figure 8.6 and Figure 8.7). Figure 8.8 shows an example of a good, non-excess standard star used for telluric correction. The LXD mode observation of known warm/cold debris disk stars did not detect any MIR or FIR excess. However, this analysis is very preliminary and greatly limited by instrumental and reduction errors, significantly more so than the SXD mode observations. Further improvement on the DRS is planned and more robust results are expected from future versions.

8.1.3.1 Example Excess Confirmation

Figure 8.6: Residual excess spectrum for ζ Aql. (image credit: Wagner 2015)
8.1.3.2 Example Non-detection

\[ \langle E_{\text{rel}}(\lambda) \rangle = \frac{(S_{\text{boot}}(\lambda) - K_{\text{best}}(\lambda))}{K_{\text{best}}(\lambda)} \]

Figure 8.7: Residual excess spectrum for \( \eta \) Lep. (image credit: Wagner 2015)

8.1.3.3 Example Standard Star

\[ \langle E_{\text{rel}}(\lambda) \rangle = \frac{(S_{\text{boot}}(\lambda) - K_{\text{best}}(\lambda))}{K_{\text{best}}(\lambda)} \]

\text{\textbullet} \ \langle E_{\text{rel}} \rangle \ \text{\textbullet} \ \text{eta Lep FIV target} \ 
\text{\texttimes} \ \text{chara excess with 3\sigma limit} \ 
\text{\times} \ \langle E_{\text{rel}} \rangle \ \text{HD 20487 AOV standard 198 pc}

Figure 8.8: Residual excess spectrum for the standard star HD20487. (image credit: Wagner 2015)
8.1.4 Discussion

This project is still in its early stages but so far has revealed promising results. With more data and further refinements, it may prove to be a robust detection technique for NIR and MIR excesses, that could be expanded to other infrared spectro-photometric facilities and instruments. The summary of results is shown in Figure 8.9 with Table 8.3 giving a comparison between IRTF, FLUOR, and JouFLU exozodiacal excess detection results.

Figure 8.9: Weighted excess averages obtained from SpeX, $E_{\text{rel}}$, from 2.0 - 2.3 μm. The error bars account for correlation between different wavelength bins. The two G8V excess stars are $\tau$ Cet (upper asterisk) and $\xi$ Boo (lower asterisk). A significant excess is detected for targets $\xi$ Boo, $\zeta$ Aql, $\tau$ Cet. The excesses of control star 30 Mon and standard star HR6977 may be due to non-astrophysical origin, such as poor instrument calibration or passing clouds that were noted in the observing logs. (credit: Wagner 2015)
Table 8.3: Spex results compared to Absil et al. (2013) with FLUOR and recent JouFLU results.

<table>
<thead>
<tr>
<th>Object</th>
<th>Excess SpeX</th>
<th>FLUOR</th>
<th>JouFLU</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\xi$ Boo</td>
<td>y</td>
<td>y</td>
<td>n</td>
</tr>
<tr>
<td>$\zeta$ Aql</td>
<td>y</td>
<td>y</td>
<td>y</td>
</tr>
<tr>
<td>$\tau$ Cet</td>
<td>y</td>
<td>y</td>
<td>-</td>
</tr>
<tr>
<td>54 Psc</td>
<td>n</td>
<td>n</td>
<td>-</td>
</tr>
<tr>
<td>$\theta$ Per</td>
<td>n</td>
<td>n</td>
<td>-</td>
</tr>
<tr>
<td>$\eta$ Lep</td>
<td>n</td>
<td>y</td>
<td>y</td>
</tr>
<tr>
<td>30 Mon</td>
<td>-</td>
<td>n</td>
<td>-</td>
</tr>
<tr>
<td>$\eta$ Crv</td>
<td>n</td>
<td>n</td>
<td>-</td>
</tr>
<tr>
<td>$\beta$ Leo</td>
<td>n</td>
<td>y</td>
<td>y</td>
</tr>
<tr>
<td>$\delta$ Leo</td>
<td>n</td>
<td>n</td>
<td>-</td>
</tr>
</tbody>
</table>

* FLUOR detection was 4 sigma, JouFLU result is 1 sigma.

8.2 Palomar Fiber Nuller

High angular resolution interferometric detections of resolved $\sim 1\%$ NIR excesses by the CHARA Array and the VLTI are at the accuracy limit of current instruments. Little is known about the origin and spatial distribution of the excess source. Proposed scenarios to explain the origin of this excess emission includes extended stellar atmospheres, and chromospheric and coronal emission. However, these explanations are unlikely as FLUOR and PIONIER lack the resolution to resolve such structures at the baselines used to measure NIR disks, recall that these are short, typically 30 m, baselines and the star itself is unresolved. Sub-stellar companions are a possible source of incoherent flux, but are too faint to contribute sufficient NIR light. As discussed previously in Chapter 5, stellar companions could produce enough flux, but are improbable after passing the various checks for companions. Additionally, no companions have been detected around most of these bright, nearby stars that have been observed by high contrast direct imaging, astrometric, and radial velocity
instruments (Absil et al. 2013). If hot exozodiacal dust is responsible for the detected effects, no corresponding MIR excess has been detected by high contrast KIN observations (Millan-Gabet et al. 2011; Mennesson et al. 2014). The exception to this is Fomalhaut (Mennesson et al. 2013; Lebreton et al. 2013), which does show a MIR excess. This suggests that NIR measurements may be a better probe of dust near the habitable zone. The absence of a MIR excess also suggests the dust population is composed of small grains, < 1μm, too small to emit in the MIR. As discussed previously, such small grains should not persist for very long in the region so close to the host star (Wyatt et al. 2007b).

Observations in the K-band (≃ 2.2μm) with the PFN, led by Bertrand Mennesson, can provide complementary information on whether the excess phenomenon is really due to hot exozodiacal light in the habitable zone (Ertel et al. 2014) or to small dust closer to the star (Lebreton et al. 2013). This is a topic of great programmatic interest to NASA with implications for future exo-planet, including exo-Earth, detection missions and for planetary formation models.

The PFN can reach deep contrast levels, a few parts in 10^4, and can confirm or deny previous detection claims, and also search for weaker excesses. The PFN has the advantage of a rotating baseline, giving it the ability to determine if the excesses are due to companions. The presence of a companion would cause a full modulation of the measured excess vs. azimuth, while an extended dust structure would result in a constant or weakly modulated signal vs. azimuth. The PFN is only sensitive to sources located between approximately 30 and 200 mas of the star. This enables the PFN to discriminate between excess emission of thermal
origin, very close to the host star, or scattered light from more extended structures.

### 8.2.1 Method

The PFN, as the name suggests, is a nulling interferometer. It utilizes destructive interference between two sub-apertures of the primary mirror. With its rotating baseline, it can detect faint companions or extended structures near stars (Serabyn et al. 2006; Mennesson et al. 2006; Martin et al. 2008; Serabyn et al. 2010). At the Palomar 200-in telescope, the PFN has a 3.4-m baseline separating the two sub-apertures and any source located between 30 and 250 mas from the central star and contributing more than 0.1% of its flux will be detected. The PFN relies on accurate visibility amplitude, rather than phase, measurements and when compared to coronagraphic or aperture masking instruments the PFN is more sensitive to hot debris disk emission. This is largely due to coronagraph’s dependence upon the stability of the telescope PSF and the difficulty of suppressing diffracted and scattered light from the central star and the removal of residual starlight.

### 8.2.2 Results

This work is still preliminary and the data are still in the process of being reduced. The effects of dispersion and atmospheric refraction must be accounted for carefully. The primary findings so far are that none of the observed CHARA NIR excess targets showed significant null excesses with the PFN. The interpretation, similar to that found for Vega (Mennesson et al. 2011b), is applied more broadly to the other excess targets. If the excesses detected by
CHARA are due to dust and it did not vary strongly between the 2 epochs, the dust is not resolved by the PFN 3.4-m baseline and therefore must reside very close to the star.

Figure 8.10 plots the transmission for the PFN, showing the angular separations where PFN can detect faint companions or extended emission. Nulling interferometer transmission is defined by Serabyn (2007) as:

\[
T(\chi) = \sin^2(\chi/2) = (1 - \cos(\chi))/2, \tag{8.1}
\]

where \( \chi \) is the relative phase between the two telescopes and \( \chi = k B \cdot \theta \).

For this work, JouFLU has provided follow-up observations and added to the PFN target list. In addition, the author assisted on observing runs with the PFN.

A summary paper by Mennesson et al. is in preparation. This will derive constraints on all individual stars as was done for Vega. Vega is the simplest case to interpret, since its outer disk is seen pole-on. So far, the following stars have been observed with PFN: \( \lambda \) Gem, \( \tau \) Cet, Altair, \( \beta \) Leo, \( \kappa \) CrB, Vega, \( \alpha \) Cep, \( \zeta \) Aql, 110 Her, 10 Tau, and \( \gamma \) Ser. None of these stars, except for a possible detection on \( \alpha \) Cep, shows a null excess larger than 0.3%, the typical \( \sim 3\sigma \) upper limit.

All are from the FLUOR excess sample, Absil et al. (2013), except \( \gamma \) Ser, for which JouFLU recently measured an excess. The physical excess could be significantly larger and still remain undetected if it originates from inside of 20 mas, and is thereby nulled by the PFN, or outside of 200 mas, the PFN FoV. If the excess is smaller than 20 mas, then it is starting to
be resolved at the shorter CHARA Array baselines. If the excess is larger than 200 mas, then it could be resolvable with AO.

Figure 8.10: Maximum PFN transmission vs. separation (baseline orientation angle chosen to maximize transmission at any given separation). (image credit: Mennesson et al. 2011a)

8.3 NIR Disks with MIRC

Some stars with RV planet detections have another low-velocity component or trend in the RV fit, possibly from a binary star or high mass planet. These systems may contain cold debris disks or inner dust disks.

At the CHARA Array, our group has begun observation of κ CrB with JouFLU and MIRC, two instruments particularly well-suited to studying systems of this nature. κ CrB is a K subgiant star that RV studies have shown is orbited by at least one planet (Johnson et al. 2008). Herschel images resolve and show excess infrared emission indicating the presence of a disk or rings of dust between ≈20 and ≈220AU (Bonsor et al. 2013b). Absil et al. (2013) have
shown that $\kappa$ CrB has an interferometrically detected incoherent flux excess. The PFN has also detected a 1% level of excess flux; this indicates that $\kappa$ CrB has a warm disk. Bonsor et al. (2013b) resolve a cold debris disk and point to an IR excess that could be due to a companion or the presence of $\mu$m-size dust grains. Measurement of the diameter of the inner dust disk could help constrain the disk temperature, semi-major axis, and separation of the planets in the system. High angular resolution data could distinguish between a single belt, dual narrow belts, or stirring models for the system presented by (Bonsor et al. 2013b).

Multiple observations with JouFLU followed by MIRC six telescope observations could be used to rule out companions and characterize the inner region of the disk. Determining the characteristics of exoplanet host stars, stars with planets detected and showing RV trends, and disk-harboring stars can provide highly valuable information for planetary system formation modeling. Our group has collected preliminary interferometry data on $\kappa$ CrB and proposals have been made to continue this effort, as well as to investigate other exoplanet and dust disk host stars of this nature.

### 8.4 Other Observations Needed

The question of NIR excesses and exozodiacal disks is a challenging one, existing at the edges of current instrument capabilities. Other observations and techniques from existing or future instruments will add invaluable data to this field. The following is a list of instruments or techniques that may bear results related to NIR excesses and exozodiacal disks and, if not them, then their descendants may do so.
Direct Imaging  In an ideal world, all questions could be resolved by exquisitely high-resolution direct imaging. Of course, this turns out to be exceedingly difficult in practice. But the next generation of large telescopes, may deliver on this promise.

- The Gemini Planet Imager (GPI) is a next generation high-contrast imaging instrument currently in the testing phase for the Gemini Telescope with the goal of directly imaging exoplanets (Macintosh et al. 2008).
  http://planetimager.org

- Spectro-Polarimetric High-contrast Exoplanet REsearch (SPHERE) is an adaptive optics and coronographic facility with three instruments that provide imaging, polarimetry, and spectrographic capabilities (Beuzit et al. 2008).
  https://www.eso.org/sci/facilities/paranal/instruments/sphere/inst.html

- The James Webb Space Telescope (JWST) is NASA’s next premier space telescope and is scheduled to launch in late 2018.

- Extremely Large Telescopes (ELTs), such as the upcoming Thirty Meter Telescope (TMT), Giant Magellen Telescope (GMT), and ESO’s ELT

Spectroscopy  could provide the answer to the question of what the dust grains are comprised of through J,H,K spectra and mineralogy. But, can the calibration be good enough to distinguish between various materials?

Interferometry at different wavebands  could continue to improve upon the high angular
resolution dust detections and perhaps constrain or resolve dust disk or ring structure and architecture. Calibration is the primary challenge here.

- **VLTI/GRAVITY (K)** $\sim 1\% V^2$ calibration (Eisenhauer et al. 2011)
  
  https://www.eso.org/sci/facilities/develop/instruments/gravity.html

- **VLTI/Multi AperTure mid-Infrared SpectroScopie Experiment (MATISSE)**
  
  (L,M,N) (Lopez et al. 2014)

  https://www.eso.org/sci/facilities/develop/instruments/matisse.html

- **Large Binocular Telescope Interferometer (LBTI)**
  

  **HOSTS**: 35 stars (68 in sample, (Weinberger et al. 2015)

  LBTI Exoplanet Exozodi Common Hunt (LEECH) (Skemer et al. 2014)

  **ZESTY**: 7 stars with outer dust reservoir

  **Polarimetry** could give some clue to if the dust is trapped in magnetic fields, or if it is reflecting scattered light or producing thermal emission. Sensitivity and calibration are the challenges here.

- **VLT/SPHERE/Zurich Imaging POLarimeter (ZIMPOL)** (Thalmann et al. 2008)

- Subaru/Visible Aperture-Masking Polarimetric Interferometer for Resolving Exoplanetary Signatures (VAMPIRES) (Norris et al. 2015) aperture masking and polarimetry
Temporal monitoring of variability of the dust; what is causing this phenomenon? Can monitoring constrain the possible models? What cadence for observations would be necessary to characterize the variability?

8.4.1 Remaining Observational Questions

Until more observations can be obtained, many questions remain unanswered. These questions may serve to direct current or future observations.

- What is the spatial distribution of the NIR emission source? Does the emitting region extend all the way to the stellar surface? Is there a cut-off or pile-up at the sublimation radius? Can we set a lower limit on the distance from the star from which the NIR emission emanates?

- In the current scheme of modeling, the dust is assumed to be at or beyond the sublimation radius. This sublimation temperature is calculated for some set of lab-determined material properties. How well do we know the properties of these materials? Could a better understanding of the material help us understand the source of the emission, in particular if the source is very near or separated from the stellar surface? A lower limit on the distance of the dust would help determine its composition, lifetime, and hence the required replenishment rate, which is used to test proposed replenishment mechanisms and source scenarios of the dust.
• Is the emission spherical or disk-like? Is there a significant change in visibility with baseline Position Angle (PA)?

• How can IR observations predict appearance at visible wavelengths, in preparation for future visible imaging missions?

• How can magnetic trapping scenarios be tested? Perhaps by observing stars with known magnetic field and searching for correlations.
“Astronomy is one of the sublimest fields of human investigation. The mind that grasps its facts and principles receives something of the enlargement and grandeur belonging to the science itself. It is a quickener of devotion.”

—– Horace Mann
Discussion

This research seeks to study near-infrared dust disks interferometrically, determine their variability, and measure their spectra. Three kinds of observations, making use of different instruments, work together to contribute to our knowledge of the hot/warm inner dust disk region around spectral type A-K stars and their role in planetary formation. A rich data set from multiple sources on a statistically significant sample can feed into models to determine the individual sub-μm grain properties, the life span of the dust, disk morphology, and the relationship to the host stars’ properties. The original FLUOR observations (Coudé du Foresto et al. 1997; Absil et al. 2013), the upgraded JouFLU (Lhomé et al. 2012; Scott et al. 2013), MIRC (Monnier et al. 2006), other instruments at the CHARA Array (ten Brummelaar et al. 2005), the PFN (Martin et al. 2008; Serabyn et al. 2010; Mennesson et al. 2011b), along with NASA’s IRTF and SpeX (Rayner et al. 2003), make obtaining these data possible. The CHARA Array and its associated instruments are unique in the world for their capabilities to resolve and image stellar photospheres and the environments that surround them.
The exact nature of the detected excesses remains elusive, but we can constrain where the dust may reside, the characteristic grain size and temperature. This work on exozodiacal variability has produced similar results for some previously observed stars, new excesses on others, and some known excesses have vanished. The fact that the new DRS reproduces compatible results when used on the original data suggests that some flaw in the DRS is not a source of variability. Likewise, the instrument produces consistent stellar diameter measures on non-excess stars that have been observed by Boyajian et al. (2012) with the CLASSIC beam combiner, so the hardware changes are not suspect. Changes in the DRS have led to larger error bars on many individual data points but the new handling of data must also be considered more robust, as it is required to pass several quality checks before being evaluated. This has led to a change in the level of significance of many exozodiacal detections. Nevertheless, we still make some strong exozodiacal dust detections in both the survey and variability study. Binarity may lead to a false detection of an excess. However, published RV data, astrometric data, and MIRC observations can rule these out. The remaining conclusion is that these excesses represent a physical phenomenon. And for most cases, it seems this phenomenon is variable on short time scales, on the order of years or less.

Where do these excesses come from? Results from the CHARA Array and the PFN show that the flux must be from either very close to the star, < 20 mas, or much further out, > 200 mas. Results from JouFLU and PIONIER provide a lower limit on the excess, but much more dust could lie beyond the FoV of the instruments. Some of our preliminary results from the IRTF
spectrophotometry also support the existence of more dust outside of the interferometric FoV.

If the source of the excess is dust, it may be undergoing multiple destruction factors: including sublimation, radiation pressure, PR drag, and collisions. There are three main models for dust: steady state/continuous replenishment, steady state/trapped nano-grains, and LHB/comet outgassing. A state of collisional equilibrium means the amount of dust would not change with age. The work of some trapping mechanism would cause an increase in the amount of dust with age. The presence of short-term variability favors the LHB or comet infall models, involving active, chaotic dust production and destruction.

Strong stellar photospheric phenomena are one possible explanation for the excess detection. However, A-type stars are not known to have particularly strong stellar winds. In the NIR, the spectral slope of free-free, or bremsstrahlung, emission does not match what would be expected. For free-free excess, the slope should go up from NIR to MIR. Also, one would expect most stellar phenomena to coincide with significant amounts of gas, which would show up in longer wavelength observations.

There is supporting evidence for variability in FIR excesses, which may provide some clues to what is happening in the NIR and MIR regions. Kate Su and her team have found that the presence of a large amount of small grains are related to stochastic, large impacts of \( \sim 500 \) km asteroids, and a short timescale is consistent with the aftermath of a large impact. For example, ID8 (2MASS J08090250-4858172) was flat in 2012, then a brightening event occurred, followed by an exponential decay, with a characteristic decay time of \( \sim 1 \) yr, and subsequent quasi-periodic rising that fits two periods of 26 and 33 days (Meng et al. 2014).
9.1 The Future of the CHARA Array & JouFLU

9.1.1 AO at the Array

The addition of AO at the CHARA Array will enable a greater number of stable, high-quality fringe scans to be recorded. This increase in the quality and quantity of data is expected to enable an additional magnitude of sensitivity in the K-band for JouFLU.

![Graph showing number of stars observable vs K magnitude limit]

Figure 9.1: The number of dwarf star targets observable from the CHARA Array as a function of K-band magnitude.

Figure 9.1 shows the number of targets observable from CHARA by K magnitude. Only main-sequence stars are counted, spectroscopic binaries and pulsating variables are excluded from the list. Note the difference in number of targets available as the limiting magnitude for JouFLU is pushed from $K = 4$ to $K = 5$, made feasible by the addition of AO at the CHARA Array. The drastic increase in sample size will enable strong statistical conclusions to be drawn on the rates of excess detection and its occurrence in relation to spectral type and other stellar parameters.
9.1.2 Integration with Fringe Tracking

Stabilizing the fringes with respect to atmospheric piston could greatly increase the statistical precision of JouFLU. To facilitate this requires the development of a fringe tracker, which works in another wave band from JouFLU. The M0 dichroics were chosen for just such a purpose. They allow K-band light to reflect to JouFLU while transmitting H-band light. It would then be possible to use the CLASSIC beam combiner as a fringe tracker. CLASSIC could operate at a faster rate than JouFLU and measure the fringe central position, sending a corrective offset to the OPLE carts, effectively ‘freezing’ the fringes for JouFLU.

9.1.3 New Combiner Technology

The possibility of upgrading the fiber combining and detector technology of JouFLU is very appealing. Integrated optics and fiber-based combiners that make use of three or more beams, such as VLTI’s GRAVITY (Eisenhauer et al. 2008) or PIONIER (Le Bouquin et al. 2011), are technologies that could be adopted for the next generation FLUOR, not to be named Jou-JouFLU. Newer detector technology has far superior noise characteristics to the current NICMOS-based detector, and is an obvious place for improvement.

9.2 Future Exozodi Work

Work with JouFLU is really just beginning. The system is now robust, both in the sense of hardware and the DRS. The results presented here represent just the first science to come
out of the new instrument. Work on exozodis is expected to continue and precision improve.

Some of our suspected binary detections and very strong exozodiacal detections should be followed up with MIRC observations to confirm or constrain binarity.

The spectral dispersion mode of JouFLU is limited to brighter objects, $K_{\text{mag}} \lesssim 3$, but will be used to characterize the spectral slope for brighter objects, such as Vega and Altair.

### 9.2.1 The Exozodiacal Disk Survey Extension

The exozodiacal disk survey has now reached $\sim 30$ objects, plus the 14 that are part of the variability survey, and incomplete data sets for a handful more. This survey will continue, toward the goal of 100 objects. Statistics from this survey will feed the exozodiacal models with information on exozodiacal rates correlated with stellar properties and spectral type.

### 9.2.2 Variability Monitoring of Hot Exozodis

The exozodiacal variability survey was conceived with the upper limit of a few years in mind as the lifespan of the dust. The goal was simply to see if the excess, or circumstellar flux ratio, varied. The next step for the exozodiacal variability study is to determine the actual period for the variation. For this, we need long-term, high-cadence monitoring of high-excess stars. This would also solidly confirm the existence of excesses and possible variability. A consistent trend for circumstellar excess amount over time would be a very strong indicator of a physical phenomenon. We propose to pick a small number of stars from the variability results that show strong variability and to monitor them with JouFLU approximately once per
month with S1-S2 and E1-E2. This would require only a couple of hours on two telescopes, scattered across the observing season.

### 9.2.3 Future of IRTF Work

The most recent IRTF observing run was unfortunately interrupted by an unprecedented triple hurricanes in the Pacific. However, this work will continue. We plan to propose for more time now that our preliminary tests have yielded some positive results. The customized IRTF DRS will be improved and should lead to an exozodiacal dust confirmation engine. Another option for spectrophotometric monitoring of exozodiacal dust is to move to a space-based approach, utilizing the Spitzer Space Telescope’s warm mission and 3.6 µm capabilities.

### 9.2.4 New Instruments

To best characterize the exozodiacal disk region, we want coverage of the visibility curve at baselines of less than 10 to 30 m where the disk deficit curve departs from the stellar photosphere. This is something that could be achieved on Mount Wilson or elsewhere with existing hardware or telescopes; or an entirely new interferometer could be constructed.

One obvious solution would be to move the Infrared Spatial Interferometer (ISI) telescopes close together to reach ~ 10 m baseline coverage and adapt them for NIR coverage. This would involved a significant time investment, but would give the facility a new purpose as an exozodiacal disk detection machine.
Another option is to build a rotating baseline NIR fiber-nulling interferometer, similar to the PFN, that could be placed on a 8-10 m class telescope, e.g. Subaru or Keck.

My favored approach would be to construct a mini-CHARA-style array out of commercial telescopes of 0.2 to 0.5 m aperture. This could be used to confirm the exozodiagal detections and to monitor their variability. Additional uv-coverage of the < 30 m baseline region would synthesize well with other on-going CHARA science goals.

Concept 1: mini-CHARA (Figure 9.2)

- An array of three or more 20-50 cm telescopes at short baselines, <10-30 m, in a Y-configuration.
- The use of a minimum of 3 telescopes would enable the measurement of closure phase.
- By mounting these telescopes on rails, an infinite number of baseline combinations would be possible.
- The use of off-the-shelf commercial telescopes and components would drastically reduce the cost of the mini-array.
- The telescopes could be connected by fiber patch cords to offset large delay, effectively creating a fiber-PoPs
- Different fibers would allow multiple wavebands.
- The mini-array would benefit from the spatial filtering properties of fibers
• Beam combination could occur after the fibers in open air, as a FLUOR/CLASSIC hybrid, or combination could occur as part of an integrated optical component.

• In the open air combination case, an optical table could be assembled as a mini-OPLE system.

• This could use a JouFLU-like XPS controlled offset and dither stage.

• This setup would be scalable to more telescopes, although packing efficiency becomes a problem at short baselines.

• The gain in sensitivity from the low number of reflections could offset the smaller aperture, which is limited by the characteristic $r_0$.

Figure 9.2: A 30-m Y-configuration interferometer.

The challenge with this setup is how to insert path delay without significant losses resulting
from fiber injection, then out for delay compensation, then back into fibers for detection would include high losses. Single-fiber entry with open air detection would seem preferable. Photometric monitoring could still be implemented in this case. Another concern would be dispersion of the light due to the lengths of fiber.

An alternative setup could be simpler. Consider a pole with two 15-m arms coming off of it, or even a single 30-m rigid beam. Telescopes would be mounted on rails on this very strong, light, and rigid structure. Then the pole could be rotated to take out the bulk delay by directly facing the telescopes to the target object. Small amounts of delay could be taken out by some table-top elements. Different baselines could be achieved by physically moving the telescopes along the beam.

Concept 2: single rotating beam array (Figure 9.3)

- An array of three or more 20-50 cm telescopes at short baselines, <10-30 m, in a rotating linear configuration.
- The use of a minimum of 3 telescopes would enable the measurement of closure phase.
- By mounting these telescopes on rails, an infinite number of baseline combinations would be possible.
- The use of off-the-shelf commercial telescopes and components would drastically reduce cost.
• The rotation of the entire beam would remove the need for large delay compensation, eliminating the need for PoPs or carts.

• Fibers could be used to spatially filter, but combination would occur after the fibers in open air, as a FLUOR/CLASSIC hybrid, or combination could occur in an integrated optic element.

• Scanning delay and detection could occur on a single optical table.

• Scalable to more telescopes.

• This could use a JouFLU-like XPS controlled offset and dither stage, making use of as many off-the-shelf components as possible.

• No metrology system between the telescopes would be needed.

• The gain in sensitivity from the low number of reflections could offset the smaller aperture, which is limited by the characteristic $r_0$.

Table 9.1 gives some approximate cost estimates for such a facility.
Figure 9.3: An alternate 30-m interferometer design. With this array, the telescopes ride a beam which is rotated to face the target, thereby eliminating the need for any large delay compensation.
Table 9.1: Sub-thirty meter array components and back-of-the-envelope cost estimate. Note: RC Optical Systems is no longer in business.

<table>
<thead>
<tr>
<th>Item</th>
<th>Cost (x $1000)</th>
<th>Notes</th>
</tr>
</thead>
<tbody>
<tr>
<td>Telescopes (each)</td>
<td>10</td>
<td>Celestron or Meade 14&quot;</td>
</tr>
<tr>
<td></td>
<td>20/30</td>
<td>0.3/0.4 m Officina Stellare</td>
</tr>
<tr>
<td></td>
<td>15/22/32.5/50</td>
<td>Planewave 14&quot;/17&quot;/20&quot;/24&quot;</td>
</tr>
<tr>
<td></td>
<td>20/25/30/40/65</td>
<td>RC Optical Systems 12.5/16/20/24&quot;</td>
</tr>
<tr>
<td>Mounts (each)</td>
<td>20-27</td>
<td>Astro-Physics 3600GTO</td>
</tr>
<tr>
<td>Auto-Guider (each)</td>
<td>1.2</td>
<td>SBIG SG-4</td>
</tr>
<tr>
<td>Optical table</td>
<td>~10</td>
<td></td>
</tr>
<tr>
<td>Mounts, mechanics</td>
<td>~50</td>
<td></td>
</tr>
<tr>
<td>Optics</td>
<td>~50</td>
<td></td>
</tr>
<tr>
<td>Active delay</td>
<td>~100</td>
<td>Newport XPS+motion stages</td>
</tr>
<tr>
<td>Fibers</td>
<td>~50</td>
<td></td>
</tr>
<tr>
<td>Construction</td>
<td>~150</td>
<td></td>
</tr>
<tr>
<td>Enclosure</td>
<td>~10</td>
<td></td>
</tr>
<tr>
<td>Detector</td>
<td>—</td>
<td></td>
</tr>
<tr>
<td>Total</td>
<td>~500k-700k</td>
<td>not including detector or integrated optics components</td>
</tr>
</tbody>
</table>
9.3 Closing Comments

The overall endeavor for this work has not taken a singular set route from observation to conclusion. Instead the approach has been to start initially with the instrument hardware assisting with its upgrade, replacement, and enhancement. Following this was a ground-up effort to rebuild the FLUOR software, utilizing the CLASSIC software developed by Theo ten Brummelaar, and adapting it into what is now the JouFLU software package. As this developed, it went from on-sky testing and qualification to science-quality data collection, the results of which have been reported herein. As this project grew, ideas for new observations and possible instruments germinated. Resolving the many questions that arise when tackling a seemingly simple problem requires new observations and capabilities, that in turn leads to new results and more questions.

Hardware The work on hardware for JouFLU has been extensive. Prior to installation, major components such as CALI, the XPS, OPD Scan, and The InfraRed camera (IRcam) underwent testing and qualification at Laboratoire d'études spatiales et d'instrumentation en astrophysique (LESIA) in Meudon, France. Working with the Laboratoire d'études spatiales et d'instrumentation en astrophysique (LESIA) engineers on this was the primary purpose of the author's first Chateaubriand Fellowship. For four months, these components were evaluated and optimized before being shipped to Mt. Wilson. Once on the mountain, the components were installed on the optical table with the author assisting Emilie Lhomè. After the install, further testing
by the author determined that CALI could not function as a suitable science camera for JouFLU. The recommendation was made to return to NICMOS for data collection. After this, the problem of differential polarization delay was discovered and measured through the author’s laboratory tests. A potential remedy for this problem is forthcoming with the addition of Lithium Niobate compensators. With the assistance of Judit Sturmann of CHARA, procedures have been developed to align JouFLU carefully within the CHARA Array optical system.

**Software** The author began work at the CHARA Array on porting the FLUOR control software to the C programming language. From that beginning, the JouFLU software has undergone a few incarnations from being a clone of the CLASSIC beam combiner software, to something of its own. Much of the coding was a learning experience for the author, with the majority of the time spent observing and collaborating with the primary architect of the CHARA Array software environment, Theo ten Brummelaar. The author contributed large portions of the code for running the NICMOS motors, XPS, the Zabers and fiber alignment process, reading data from NICMOS, and the GUI.

**Observation** The author has been present for every nighttime operation of JouFLU and a few of the last observations made with FLUOR in 2011. These nights have ranged from the initial testing and debugging of JouFLU to the planning of observing runs and all data collected for these projects. Knowledge gained from these nights frequently fed directly back into improvements on the JouFLU hardware or software. The author also led the project to utilize NASA’s IRTF to obtain independent confirmation of exozodiacal
excess, a project that became the subject of a Master’s thesis for Raphaela Wagner.

**Results of the observations** This work on exozodiacal variability was able to replicate results found for some previously observed stars, finding that the amount of excess flux had not varied over the given time period. However, the stars \( \upsilon \) And, \( \kappa \) CrB, and \( \gamma \) Ser show strong signs of exozodiacal excess variability. Two stars \( \upsilon \) And and \( \kappa \) CrB show an increase in previously detected excesses. The star \( \gamma \) Ser is a new excess detection where previously there was no excess found. The DRS was tested using previous data from FLUOR and gave consistent results, suggesting this variability is not an artifact from the new reduction process. The instrument hardware changes are eliminated as a potential source of the variability by getting consistent results when measuring the stellar diameter of non-excess stars observed independently by Boyajian et al. (2012) with the CLASSIC beam combiner.

**Future goals** Interest in exozodiacal dust and the environments near the habitable zone around stars will increase. The questions raised about the nature, frequency, and location of this dust will remain pertinent and highly relevant in the search for Earth-like exoplanets. The projects described here will need to be continued and expanded including the development of possible alternative and dedicated approaches to tackle this subject fully. The author has proposed the continued use of JouFLU and discussed some of the potential advantages provided by additions such as the installation of AO at the CHARA Array. Some concepts for future dedicated and low-cost facilities to specialize in this problem have been presented.
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Appendices
Before operating the instrument, NICMOS should be pumped down to vacuum and filled with LN$_2$, as per CHARA laboratory procedures. Once cooled, the camera electronics may be turned on. This is done by first switching on the NICMOS power supply, located above the JouFLU optical bench and labeled “1$^{st}$”. Next, the camera interface electronics must be turned on by switching on the power supply to the box labeled “2$^{nd}$”. Finally, the NICMOS PC may be powered on and the JouFLU server started.

A.1 The Server

The JouFLU server is the basic interface between the user, the instrument hardware, and CHARA subsystems. All JouFLU functions may be handled directly from the server. There is a hierarchical menu-based system for JouFLU subsystems such as the XPS and NICMOS. Commands may also be accessed by direct entry on the command-line within the server. All JouFLU commands are listed in the file: “jouflu_functs.c”
Figure A.1: The JouFLU server gives information on the XPS stage positions, the NICMOS camera readout mode and rate, and offers direct control over all JouFLU systems.
Table A.1: Functions that may be entered in the JouFLU server command-line.

<table>
<thead>
<tr>
<th>Command</th>
<th>Function called</th>
<th>Purpose</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td><strong>Standard Function Set</strong></td>
<td></td>
</tr>
<tr>
<td></td>
<td>include std_ui_functs.h</td>
<td>the standard CHARA user interface functions are included</td>
</tr>
<tr>
<td></td>
<td>include std_astro_functs.h</td>
<td>the standard CHARA astronomy functions are included</td>
</tr>
<tr>
<td></td>
<td><strong>Replacements for normal std_rt_functs.h</strong></td>
<td></td>
</tr>
<tr>
<td>settime</td>
<td>set_time</td>
<td>sets the time</td>
</tr>
<tr>
<td>clrcnt</td>
<td>clear_lost_counts</td>
<td>clears lost counts</td>
</tr>
<tr>
<td>time</td>
<td>show_time</td>
<td>shows the current time</td>
</tr>
<tr>
<td></td>
<td><strong>Local Function Set</strong></td>
<td></td>
</tr>
<tr>
<td>ds</td>
<td>call_motors_deselect</td>
<td>deselects one of the motors that move NICMOS</td>
</tr>
<tr>
<td>omp</td>
<td>open_motors_port</td>
<td>opens the port for the NICMOS motors</td>
</tr>
<tr>
<td>cmp</td>
<td>close_motors_port</td>
<td>closes the port for the NICMOS motors</td>
</tr>
<tr>
<td>mm</td>
<td>call_motors_move</td>
<td>used to move one of the NICMOS motors</td>
</tr>
<tr>
<td>ots</td>
<td>open_tracking_socket</td>
<td>opens the tracking socket for OPLE</td>
</tr>
<tr>
<td>cts</td>
<td>close_tracking_socket</td>
<td>closes the OPLE tracking socket</td>
</tr>
<tr>
<td>beams</td>
<td>set_beams_used</td>
<td>sets the beams used to 3&amp;4 or 5&amp;6</td>
</tr>
<tr>
<td>int</td>
<td>set_internal_fringes</td>
<td>used to set for internal (lab) fringes, bypasses OPLE</td>
</tr>
<tr>
<td>internal</td>
<td>set_internal_fringes</td>
<td>—</td>
</tr>
<tr>
<td>pi</td>
<td>set_pi_name</td>
<td>sets PI name for the FITS header</td>
</tr>
<tr>
<td>prog</td>
<td>set_program_name</td>
<td>sets program name for the FITS header</td>
</tr>
<tr>
<td>program</td>
<td>set_program_name</td>
<td>—</td>
</tr>
<tr>
<td></td>
<td><strong>XPS functions</strong></td>
<td></td>
</tr>
<tr>
<td>oxps</td>
<td>open_xps_port</td>
<td>open the port for the XPS</td>
</tr>
<tr>
<td>cxps</td>
<td>close_xps_port</td>
<td>close the XPS port</td>
</tr>
<tr>
<td>vxps</td>
<td>call_get_xps_version</td>
<td>returns the version of the XPS firmware</td>
</tr>
<tr>
<td>sxps</td>
<td>call_status_get</td>
<td>returns the status of the XPS</td>
</tr>
<tr>
<td>Function</td>
<td>Description</td>
<td></td>
</tr>
<tr>
<td>----------</td>
<td>-------------</td>
<td></td>
</tr>
<tr>
<td>hxps</td>
<td>call_kill_init_home</td>
<td>kills XPS stage motion, re-imits, and homes the stage</td>
</tr>
<tr>
<td>rmx</td>
<td>call_move_relative</td>
<td>performs a relative move of an XPS stage</td>
</tr>
<tr>
<td>amx</td>
<td>call_move_absolute</td>
<td>performs an absolute move of an XPS stage</td>
</tr>
<tr>
<td>sgg</td>
<td>call_sgamma_get</td>
<td>returns the current sgamma settings for the XPS</td>
</tr>
<tr>
<td>sgs</td>
<td>call_sgamma_set</td>
<td>sets new sgamma settings</td>
</tr>
<tr>
<td>txp</td>
<td>call_test_XPS</td>
<td>performs a set motion of a stage</td>
</tr>
<tr>
<td>iex</td>
<td>call_init_events</td>
<td>init an XPS event</td>
</tr>
<tr>
<td>rex</td>
<td>call_remove_event</td>
<td>removes an XPS event</td>
</tr>
<tr>
<td>OUT</td>
<td>call_move_OUT</td>
<td>rotate the OUT stage</td>
</tr>
<tr>
<td>FTS</td>
<td>call_move_FTS</td>
<td>move the FTS stage</td>
</tr>
<tr>
<td>ALIU_L2</td>
<td>call_move_ALIU_L2</td>
<td>move the ALIU L2 (viscam focus) stage</td>
</tr>
<tr>
<td>ALIU_A</td>
<td>call_move_ALIU_A</td>
<td>move the ALIU beam A stage</td>
</tr>
<tr>
<td>ALIU_B</td>
<td>call_move_ALIU_B</td>
<td>move the ALIU beam B stage</td>
</tr>
<tr>
<td>OPD_STAT</td>
<td>call_move_OPD_STAT</td>
<td>move OPD STAT</td>
</tr>
<tr>
<td>single_scan</td>
<td>call_single_scan</td>
<td>perform a single scan with OPD SCAN</td>
</tr>
<tr>
<td>getpos</td>
<td>call_get_position</td>
<td>returns the position of a stage</td>
</tr>
<tr>
<td>getposition</td>
<td>call_get_position</td>
<td>—</td>
</tr>
<tr>
<td>sto</td>
<td>call_set_XPS_timeout</td>
<td>set the XPS timeout delay</td>
</tr>
<tr>
<td>abort_move</td>
<td>call_abort_move</td>
<td>abort a movement</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>IR cam functions</td>
<td></td>
<td></td>
</tr>
<tr>
<td>shuttogle</td>
<td>ircam_shutter_toggle</td>
<td>toggle the IRCAM(pupil camera) shutter position</td>
</tr>
<tr>
<td>shutt</td>
<td>ircam_shutter_toggle</td>
<td>—</td>
</tr>
<tr>
<td>irview</td>
<td>call_ircamview</td>
<td>view the CHARA pupil</td>
</tr>
<tr>
<td>sampleviewer</td>
<td>call_run_sampleviewer</td>
<td>run the SampleViewer program to view the pupil</td>
</tr>
<tr>
<td>sv</td>
<td>call_run_sampleviewer</td>
<td>—</td>
</tr>
<tr>
<td>NICMOS control functions</td>
<td></td>
<td></td>
</tr>
<tr>
<td>ons</td>
<td>call_open_nicmos_socket</td>
<td>open the NICMOS socket</td>
</tr>
<tr>
<td>cns</td>
<td>call_close_nicmos_socket</td>
<td>close NICMOS socket</td>
</tr>
<tr>
<td>Mode</td>
<td>Call Nicmos Function</td>
<td>Description</td>
</tr>
<tr>
<td>------</td>
<td>----------------------</td>
<td>-------------</td>
</tr>
<tr>
<td>Mode</td>
<td>call_nicmos_set_mode</td>
<td>Set the NICMOS mode</td>
</tr>
<tr>
<td>Arm</td>
<td>call_nicmos_arm</td>
<td>Arm NICMOS</td>
</tr>
<tr>
<td>Movie</td>
<td>nicmos_movie</td>
<td>Enter “movie” (frame) mode</td>
</tr>
<tr>
<td>Trm</td>
<td>toggle_nicmos_reset_method</td>
<td>Toggle the NICMOS reset method</td>
</tr>
<tr>
<td>Rate</td>
<td>call_nicmos_set_sample_rate</td>
<td>Set NICMOS rate</td>
</tr>
<tr>
<td>Noop</td>
<td>call_nicmos_noop</td>
<td>Getter for NICMOS no operation</td>
</tr>
<tr>
<td>Mess</td>
<td>call_nicmos_message_message</td>
<td>Call a message from NICMOS</td>
</tr>
<tr>
<td>Allnoop</td>
<td>call_nicmos_all_noop</td>
<td>All NICMOS no operation</td>
</tr>
<tr>
<td>Setup</td>
<td>call_nicmos_set_setup</td>
<td>Setup NICMOS</td>
</tr>
</tbody>
</table>

**Controlling printer port driver**

<table>
<thead>
<tr>
<th>Function</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>Nframes</td>
<td>call_set_jouflu_nframes</td>
</tr>
<tr>
<td>Nfwait</td>
<td>call_wait_jouflu_nframes</td>
</tr>
<tr>
<td>Cdwait</td>
<td>call_wait_jouflu_collecting_data</td>
</tr>
</tbody>
</table>

**Doing things with NICMOS data**

<table>
<thead>
<tr>
<th>Function</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>Phot</td>
<td>test_photometry</td>
</tr>
<tr>
<td>Photom</td>
<td>test_photometry</td>
</tr>
<tr>
<td>Frame</td>
<td>call_nicmos_get_frame</td>
</tr>
<tr>
<td>Cgss</td>
<td>call_get_single_scan</td>
</tr>
<tr>
<td>Mss</td>
<td>set_motor_step_size</td>
</tr>
<tr>
<td>Align</td>
<td>align_fiber</td>
</tr>
<tr>
<td>Af</td>
<td>align_fiber</td>
</tr>
<tr>
<td>Data</td>
<td>get_data</td>
</tr>
<tr>
<td>Sds</td>
<td>set_data_scans</td>
</tr>
<tr>
<td>Width</td>
<td>edit_filter_width</td>
</tr>
<tr>
<td>Eem</td>
<td>edit_envelope_mean</td>
</tr>
<tr>
<td>Mean</td>
<td>edit_envelope_mean</td>
</tr>
<tr>
<td>Env</td>
<td>edit_envelope_mean</td>
</tr>
<tr>
<td>Pss</td>
<td>edit_ps_smooth</td>
</tr>
<tr>
<td>Smooth</td>
<td>edit_ps_smooth</td>
</tr>
<tr>
<td>Code</td>
<td>Function</td>
</tr>
<tr>
<td>------</td>
<td>----------</td>
</tr>
<tr>
<td>dcs</td>
<td>edit_dc_suppress</td>
</tr>
<tr>
<td>suppress</td>
<td>edit_dc_suppress</td>
</tr>
<tr>
<td>bls</td>
<td>call_save_baseline_solution_data</td>
</tr>
<tr>
<td>esd</td>
<td>edit_scan_delay</td>
</tr>
<tr>
<td>delay</td>
<td>edit_scan_delay</td>
</tr>
<tr>
<td>smf</td>
<td>call_set_maxframes</td>
</tr>
<tr>
<td>tnds</td>
<td>toggle_use_new_data_sequence</td>
</tr>
</tbody>
</table>

**Zabers**

<table>
<thead>
<tr>
<th>Code</th>
<th>Function</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>zss</td>
<td>call_set_zaber_step_size</td>
<td>set the raster scan step size</td>
</tr>
<tr>
<td>szd</td>
<td>call_set_zaber_position_default</td>
<td>set current as the zaber default position</td>
</tr>
<tr>
<td>szf</td>
<td>call_scan_init_file</td>
<td>read the saved init file for Zaber positions</td>
</tr>
<tr>
<td>wzf</td>
<td>call_write_init_file</td>
<td>write a new Zaber init file with current defaults</td>
</tr>
<tr>
<td>mzd</td>
<td>call_move_zaber_default</td>
<td>move a Zaber</td>
</tr>
</tbody>
</table>

**Fake data control**

<table>
<thead>
<tr>
<th>Code</th>
<th>Function</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>fake</td>
<td>edit_fake</td>
<td>edit fake data parameters</td>
</tr>
</tbody>
</table>
A.2 The GUI

In practice, the JouFLU GUI provides convenient access to functions provided by the server. Not all functions are available via the GUI, but all functions needed for typical setup and normal operation are readily available.

Figure A.2: The JouFLU GUI (bottom) shown while recording white light lab fringes. Clockwise from top left: the server, the fringes, photometry per channel, fringe power spectrum, a waterfall plot of the fringes, and a waterfall plot of the power spectrum.

Figure A.3: The JouFLU GUI - main tab. The number of scans to be recorded during the data, shutter, or dark sequence may be set here. Additionally, the OPLE tracking socket can be changed manually.
Figure A.4: The JouFLU GUI - setup tab. The NICMOS settings may be changed here, including camera read rate, the reset mode (destructive or non-destructive), and which beam the instrument is on. The readout region of NICMOS may be set manually, as can Nreads and Nloops. Finally, spectral dispersion mode may be enabled here, and the number of pixels read from each fiber may be set. Note: spectral dispersion mode also requires the movement of the XPS OUTPUT stage to insert the prism into the beam.

Figure A.5: The JouFLU GUI - XPS tab. In the top left of this tab is a pull-down menu that lists all of the available XPS stages, and the option for all stages to be selected. The “HOME” button homes the stages selected with this pull-down. Next to the home button are the relative move left and relative move right buttons. When clicked, they move the stage selected in the pull-down by the amount (in $\mu$m) entered in for “Step”. The “Step” entry window is also used for the absolute movement button, “abs”. In this instance the stage will move to the absolute position entered in “Step”. Below this region of the tab are the set positions for each stage. Next to the stage name is a pull-down menu with the available default positions. To move the stage to the desired position, click “GO”.
Figure A.6: The JouFLU GUI H-band pupil camera tab. Here the integration time for the pupil viewing, IRCAM can be entered, and the external viewing program can be executed.

Figure A.7: The JouFLU GUI - alignment tab. Here the raster scan parameters may be set. “Pixmult” simply affects the display size of the raster image. “Beam” selects which beam you are aligning. “Size” refers to the size of the raster scan, in number of steps. “Align” starts the alignment. The “GOTO” buttons allow the user to select the best position to center in the window, while “Reject” cancels the scan and returns to the start position. In the next row, the number entry window is the raster scan step size. 40 is the default step size used for the Zabers, but 20-120 can be used, to change the size enter a number, then click “STEP”. “SCAN CONFIG” reads the saved init file for the default positions. The “GOTO” buttons move the Zabers to these default positions for the A and B beams. This is very useful if the spot gets lost. The “SET” buttons set the current Zaber positions as the defaults for beams A and B. Finally, the “WRITE CFG” button writes the currently saved Zaber positions to the init file.
Figure A.8: The JouFLU GUI - a 5x5 raster scan. The red box marks the maximum, while the green cross-hair marks CoG of the surrounding pixel region. Note: CoG is very only accurate in small raster scans.

Figure A.9: The JouFLU GUI - picture tab. This controls the “movie” mode of NICMOS and the motors that move the position of the entire NICMOS stage. “START” begins collecting and displaying frames from NICMOS. The directional arrows move the camera by the amount set with “STEP”. The default is 200 and can only be changed while the camera is not running. The buttons below “<” and “>” all control the display of the frames. They can change the bias or the scaling of the display.
Figure A.10: The JouFLU GUI - picture tab while running. This shows the extra information that is made available while the camera is running. Next to each fiber, they are listed respective to the position they appear in the frame, is the fractional amount of flux in each pixel, relative to the surrounding pixels. This is followed by the total flux detected in that pixel and the uncertainty. The sums for the photometric and interferometric, or signal, channels is also given. "TolFrc/Sum/PixErr" gives the fraction total and total sum for all four pixels, while the last number, pixel error, is representative of the total alignment of the camera readout pixels and the flux. It is optimized when it is very small, $< 0.2$. The last row gives the display range minimum and maximum.
Figure A.11: “movie” mode. This display gives live visual feedback from the NICMOS camera. The amount of flux in each pixel from the fiber bundle is displayed along with alignment cues. The green boxes are the four actively read pixels. The red crosses demarcate the centroid of the light in the region near each pixel. The green cross and red box in the center represent the mean centroid of all four spots. This enables accurate alignment of the camera with respect to the output stage. The three grey pixels are known bad pixels and are forced to a single value.

Figure A.12: The JouFLU GUI - photometry tab. This tab simply has controls to start and stop the photometry collection mode, along with controls to zero the bias, take a new bias, and adjust the display scaling.
Figure A.13: The JouFLU GUI - photometry tab while running. New information becomes available while the photometry mode is running. “Ndata” is the number of data points collected per scan. The Sample rate is the camera read rate in Hz. The plot range is displayed on the next line. For each beam, the flux is plotted in its corresponding color. The mean, standard deviation, variance, and maximum is listed for each channel. In the separate windows, the amplitude and power spectrum of the photometry is plotted.
Figure A.14: The JouFLU GUI - data tab. The data tab for JouFLU should be familiar to anyone that has used CHARA CLASSIC. The dither pitch can be set, as can the number of samples recorded per scan. The ‘<’ to ‘>>’ buttons are used to control the motion of the offset stage. This is the OPLE cart during night observing, or the OPD STAT while testing with the internal WL source. The ‘S<’ and ‘S>’ buttons provide the option for a continuous scan. When depressed, the cart will continue scanning while the offset is increased, so that there are three scans over a given delay space. If fringes are detected, by surpassing a set threshold in the power spectrum, the offset stage will stop. The remaining buttons operate identically to other CHARA instruments and are used to control how the fringes are displayed, to turn off filtering, change the display scale, and to send offsets to “Cosmic Debris”.

Figure A.15: The JouFLU GUI - data collection tab while running. When running, the data tab shows estimates of the limiting magnitude, the interferometric visibility, the interferometric visibility for each signal channel, the SNR. Also, shown is the cart offset position, and the sum of the number of counts in the signal channels. Around the JouFLU GUI, the signal difference channel is shown, here showing the “fringe packet’ (top middle), a waterfall display of the same is shown (bottom middle), as is the power spectrum (bottom right), and a waterfall display of the power spectrum (bottom left). The photometry for the four channels is given in real-time (top right).
Figure A.16: The JouFLU GUI - status tab. This status relays the information displayed in the server terminal.

Figure A.17: The JouFLU GUI - configure tab. Here the setup for the CHARA Array and JouFLU can be confirmed or manually changed. In general, this tab should be automatically populated.
A.3 On-sky Operation

The following checklist assumes a prior laboratory alignment with green laser and general knowledge of how to operate the CHARA Array.

1. Power on NICMOS, its electronics, and its computer. The first switch is labeled 1st with blue tape and is on a silver box on top of the JouFLU table cover. The second switch is a power strip, with a blue electronics box labeled 2nd plugged into it, next to the NICMOS computer on the work table in the lab. This turns on electronics that interface with the NICMOS computer.

2. From the control room, turn on the NICMOS computer using the power GUI. It must be unlocked first, to be able to turn on/green.

3. Start the JouFLU server: bottom-right menu button, click on “Servers-2”, then JouFLU.

4. Use the “tndts” function in the JouFLU server to verify the proper data sequence for your use. The “old sequence” uses the shutters for backgrounds, the “new sequence” moves the OPLE carts “off-fringe” by 2-cm to record backgrounds.

5. Open the GUI: bottom-right menu button, click on “GTK”, then JouFLU.

6. In the server the NICMOS time and CHARA time should be synchronized. If not, cycle the NICMOS power in the power GUI, wait two minutes and then re-open NICMOS in the JouFLU GUI.
7. Set the camera acquisition rate to 500 Hz and the filter to K' using the “Cosmic Debris” GUI.

8. The operator may now acquire a star.

9. Align the fibers using the JouFLU GUI: align tab. Use a raster size of 11 for the first alignment. In the raster display that appears you should see a red box and a green cross-hair overlaid on the raster scan. Once the scan completes you should see a bright region from the star. Based on how this appears, go to MAX (the red box) or CoG (the green cross-hairs) or REJECT. You should see the flux as a spot on the raster scan. If the spot is off to the side of the scan window, do another scan after you have clicked “GOTO MAX” or “GOTO CoG”. The number of counts with the WL source should be a few thousand, or a few hundred on a bright star.

10. Repeat the previous step for the other beam. Future raster scans may use a smaller size such as 5, 7, or 9. The optimum size depends upon seeing conditions. The raster step size may also be adjusted by changing the value next to the “STEP” button, and then clicking “STEP”. Suggested values are 20 or 40 in good conditions. If there is trouble finding the spot in the first raster scan of size 11, a Zaber step size of 80 or 120 may be used to create a larger search scan. A raster must be done for each beam after every slew.

11. Once you have a good alignment for both beams, you may save these positions as a default, and to an init file. This way the Zaber positions will not be lost in the event of a
server crash. This is done by clicking “SET A”, “SET B”, and “WRITE CFG”.

12. Check for flux from all four fibers on the camera by clicking “START” in the “PICTURE” tab. Move the camera with the directional arrows until pixel error is less than $\sim 0.2$.

Press stop when done. The NICMOS motors step size can be changed by adjusting the number next to the “STEP” button and then clicking “STEP” while the camera is not running.

13. At this point you should see flux in all four fibers, and are ready to record data. This is triggered in “Cosmic Debris” and controlled in the JouFLU server in the same way as the CHARA CLASSIC beam combiner. “SCAN FOR FRINGE” is used to find the fringes, once found “HOLD” is clicked in the JouFLU GUI, followed by “SEND”. Next, “RECORD TWO BEAMS” is clicked in “Cosmic Debris” and the “<” and “>” buttons in the JouFLU GUI. Once fringes are found, clicking “SERVO” in the JouFLU GUI followed by “SAVE”, closes the servo loop and begins recording frames.
## A.4 Troubleshooting

Table A.2: Troubleshooting JouFLU, some possible problems and their solutions.

<table>
<thead>
<tr>
<th>Problem</th>
<th>Possible solution</th>
</tr>
</thead>
<tbody>
<tr>
<td>“Timed out waiting for data” error while trying to scan for data</td>
<td>restart NICMOS PC&lt;br&gt;wait 2 mins&lt;br&gt;click “Reopen NICMOS”</td>
</tr>
<tr>
<td>Server fails to start</td>
<td>check that you can bring up the zaber _1 GUI if the Zaber server has died JouFLU will not start&lt;br&gt;sockman may report zaber _1 as alive but the GUI will not start, restart the Zaber server</td>
</tr>
<tr>
<td>No light in the raster scan</td>
<td>check that there is no obstruction in the lab&lt;br&gt;perform a large raster scan, 11x11, with a Zaber step size of 120&lt;br&gt;retro-inject the red laser, and check the M3/Zaber alignment</td>
</tr>
</tbody>
</table>
B

JouFLU Alignment Procedures

The following assumes knowledge of the CHARA laboratory and general alignment procedure. The most common JouFLU alignment procedures are addressed. In general, these alignments are rarely needed, do not begin re-aligning any of the JouFLU optical table setup unless you are very sure it is necessary and you know what you are doing.

B.1 ALIU

In some instances, when checking the internal JouFLU alignment with the CHARA WL source, the Zaber positions when the ALIU stages are in the open position, also called “FTS OUT OUT”, or in the “dichroic” position may differ greatly. This is a sign that the ALIU optics are out of alignment. To correct this: first get the Zabers in the correct position, by using the “align” command to perform a raster scan, so that flux is detected on the four NICMOS spots. The target pixels are marked by white boxes in the frame. Next, set the XPS OUTPUT stage to the LED or “BackLight (BL)” position and the ALIU stage to “dichoic”. Finally, adjust
ALIU_M2 to overlap spots of the CHARA WL and the JouFLU red LED on the Visible band alignment camera (Viscam). Do this for both beams, A (3 or 5) and B (4 or 6).

**Figure B.1: JouFLU in the configuration used when aligning with the ALIU system and the visible camera (VISCAM).** The ALIU stage for the desired beam is moved to the dichroic position, and the light from the external source or star passes through a focusing lens (L1) and another dichroic (D2) to reach a fold mirror (M2), which directs it to the alignment camera (left). To check the CHARA pupil for vignetting or other possible loss of flux, an ALIU stage is moved to the mirror position and light is passed through a the same lens used by the VISCAM, reflects off of dichroic D2, and reaches a focusing lens mounted on a stage (L2). The CHARA pupil is then recorded with the infrared camera (right). (image credit: LESIA - CNRS/Observatoire de Paris)

### B.2 Camera Objective

The objective that forms the image of the four fiber spots on NICMOS is mounted on 3 pico motors, listed on CHARA computers under the pico3 controller. The first two of these motors control the tilt of the objective about the X and Y axes, with the third controls the translation in Z, or focus. The command is: “pico3 -jf-focus”.

To adjust this objective use the CHARA WL source on the engineering beam and run the JouFLU “movie” mode. While viewing the NICMOS movie output, adjust pico3 -jf to optimize
the spots.

**B.3 Fiber Input Stages/OAPs**

The alignment of the JouFLU INput stage (IN) and OAPs is possibly the most tricky alignment to perform. Perfecting the OAP alignment may take several iterations.

In the CHARA laboratory, the addition of a red laser which follows the same optical path as the WL source has greatly improved the ease of JouFLU input alignment. This laser can be used in place of the green laser to align to the JouFLU optical table. At times, it is not possible to verify the visibility of the WL and it may not appear in a raster scan performed by the Zaber mounts. To improve alignment, and to possibly rule poor alignment out as a possible cause of low flux, the JouFLU red fiber laser can be retro-injected through the MONA fiber combiner. The position of the Zabers can then be optimized by adjusting the retro laser spot on the small targets used by CHARA. These small targets are on the telescope optical benches, just after the beam reducing telescopes. Once found, this Zaber position should be recorded and used as the default for future raster scans. Alternatively, the retro-injected JouFLU red laser can be reflected off of the telescope table corner cubes and checked on the JouFLU table targets.

**B.3.1 Fine Alignment of input OAPs**

This method works once you have the JouFLU table aligned and have found spots on both beams with the raster scan. If this alignment is not sufficient then the coarse alignment, listed
below, may be tried. After performing a coarse alignment, then fine alignment should be performed.

1. The corner cubes should be in beams.

2. Align the WL to the laser spot in the WL alignment telescope. Make sure that you have ND 3.0 or greater in the laser before looking through the alignment telescope. Then, adjust the WL with the pico controller above the telescope to overlay the laser spot.

3. Turn on the REF CCD, start the refcam server, and open the ref shutter.

4. Use ESP BC2 to focus the ref on the laser spot. This spot should be small, bright, and round.

5. Set the JouFLU OUTPUT stage to the BL position.

6. Plug in the input stage pico motors to the pico boxes on table. Use the ones labeled for “camera output objective”.

7. Use the picos, output-obj and focus, to adjust the BL spot on the ref cam. Use the Zabers to keep the spot centered.
   a. If the spot is astigmatic and wider in the horizontal direction, make adjustments to the horizontal axis.
   b. If the spot is astigmatic and wider in the vertical direction, make adjustments to the vertical axis.
   c. Adjust the focus pico frequently.
8. The maximum number of counts is displayed in the refcam GUI.

9. Once you have a small, bright, and round spot: move OUTPUT stage to “OUT”, open WL shutter, and run “align”.

B.3.2 Coarse Alignment of input OAPs - recommended method

Only adjust the fiber input stages if needed, this is rare, and after all other alignments are very good. To reach a good alignment of each OAP can easily take a day or more.

B.3.2.1 To find the OAP optical axis

1. Make all normal laboratory alignments up the the to JouFLU OAP using green laser.

2. Unplug the fiber from the input stage.

3. Remove the female fiber connector from its dovetail by removing the two small screws that hold it.

4. Place translucent tape on the fiber side of the vertical plate of the dovetail that holds the fiber connector.

5. Adjust the Zaber to center the laser spot on the tape at the focus.

6. Slide the fiber holder along the dovetail while watching the green laser spot on the tape.

7. Adjust the tip and tilt of the OAP so that the axis of the OAP matches that of the fiber by checking the spot position across the full range of travel along the dovetail.
8. After the axis is matched, adjust the XY translation of the stage to roughly center the spot in the fiber holder.

**B.3.2.2 To align the OAP/fiber injection stage:**

1. Place a target in front of the Zaber along the beam axis and align it using the green laser. This target will block the Zaber mirror.

2. For aligning beam A, place the CHARA laboratory alignment telescope on the MIRC table. NOTE: this is not the surveyor’s theodolite, but instead this is the small telescope with a metal finish that is mounted between two large upright rails. Fit it with the right angle eyepiece for easier viewing. If aligning beam B, the alignment telescope may need to be placed on a different table. Place it on the table holding the CHARA shutters and use a fold mirror placed between the shutters and aperture wheels to redirect the light.

3. Align the alignment telescope using the JouFLU table front targets and the target that was placed in front of the Zaber. Alternate between the two targets to insure both the translation and rotation are correct the alignment telescope is on the optical axis.

4. Place a small flashlight nearly on the beam axis and aim it at the Zaber so that the fiber tip is illuminated.

5. View the fiber tip with the alignment telescope and verify that the OAP is on axis by sliding the holder along the dovetail. Do this with the fiber unplugged.
6. Place the fiber holder at the front of the stage and plug the fiber in. Adjust the XY positions of the input stage to center the fiber tip in the alignment telescope. The hole in the vertical plate takes up most of the FoV.

7. Use the JouFLU software to put the OUTPUT stage at the red LED BL position.

8. Use the Zaber GUI and a step size of $\sim 100$-400 to manually create a raster scan.

    Move the Zaber one step, while noting the direction, and then adjust the XY positions of the fiber stage to re-center the spot in the alignment telescope.

9. The goal is a small, bright, and round spot. A poorly aligned stage will produce an exaggerated “shark fin” shape, due to large amounts of coma and astigmatism.

10. Alternate between stepping the Zaber and centering the spot with the fiber stage.

    During this process you may have to continuously adjust the focus of the alignment telescope. This is fine, just remember to never move the alignment telescope or you will lose the optical axis.

11. Once you have a decent spot, note the position of the Zaber and move to viewing the WL source with NICMOS and making an automatic raster the Zaber using “align” procedure.

12. If you can auto-collimate the alignment telescope, you can use it to optimize the focus of the stage. Focus on the telescope auto-collimate target, then do not adjust the focus of the telescope while you adjust the focus of the stage.
13. When you think the alignment is good, and you have aligned using NICMOS. You can check the counts in the “photom” procedure.

14. Once you have decent counts in “photom”, adjust the focus of the fiber stage to maximize the number of counts.

15. Run “align” again after you change the focus.

16. You can use the CHARA WL alignment telescope, to compare the JouFLU LED in each beam by using the CHARA shutters. Note: For this step it is not the same alignment telescope you used for this procedure, but the one permanently located by the laptop closest to the green laser source.

B.4 M3/Zabers

B.4.1 Fine Alignment

1. Turn on the green CHARA laser with ND 2.0 in place.

2. Look at OUTPUT red LED BL in Viscam with ALIU in the “dichroic” position.

3. Move the Zaber to co-align the BL and the laser spot.

4. Move the ALIU stage to the “OUT” position.

5. Perform a raster scan.
B.4.2 Coarse Alignment

This coarse alignment of M3 and Zabers on the JouFLU table can be done using irises and/or targets in the beam. M3 is the mirror between the OPD stages and the Zabers. This “iris” or new target method, uses irises or new targets on the JouFLU table with the CHARA aperture wheels as the distant target. This alignment is the solution if raster scans do not show a spot and the M3 alignment is suspect.

1. Align two beams to the JouFLU table.

2. Use the CHARA engineering (eng) beam.

3. Place the targets on the JouFLU table using the green laser to align them.

4. Retro-inject the red laser at back of MONA using one of the photometric channels.

5. Adjust the Zaber to center the red light on the near target on JouFLU table. Use the iris to stop down the beam if necessary.

6. Adjust M3 on JouFLU table to center spot at the far target, the one at the beam samplers or closed aperture wheels.

7. Iterate these steps until the adjustments converge and spot is aligned at both targets.

   This alignment should result in the M3 having the correct angle with minimal beam shear.

8. Open the irises.

9. Turn on the WL.
10. Perform a raster scan and you should find the spot.

**B.5 M1/M2 OPD Stage**

M1 and M2 are flat mirrors that make up the dihedral on top of the OPD stages.

1. Align to the JouFLU table targets using the green laser.

2. Place a paper target over M2 and adjust M1 to center the laser spot.

3. Place a target over M3 and adjust M2 to center laser spot.

4. Move OPD Stat over its full range and the spot should not move.

5. If the spot moves as OPD Stat moves, adjust M1 and M2 so the spot remains stable across the full range.

**B.6 Output Fiber Stage**

The JouFLU OUTPUT stage, from the fiber bundle output to the camera aperture, was fully realigned with the assistance of Judit Sturmann. The fold mirror was removed and the fiber output stage was adjusted while viewing the four spots of the fiber bundle illuminated by a flashlight at the input. While viewing the four spots with the theodolite the stage was adjusted to minimize aberrations and to illuminate all four spots evenly. The fold mirror was then replaced and the JouFLU red laser was injected into the fiber bundle. This beam used to adjust the tip/tilt of the fold mirror onto the camera.
1. Carefully remove the OUTPUT fold mirror that directs light to the camera axis.

2. Use the red laser in place of the fiber bundle to position the theodolite.

3. Plug the fibers into the testing stand and shine a flashlight onto them. A very obvious, but perhaps necessary note: make sure the flashlight used for JouFLU tests is one with an incandescent bulb; the fibers and camera filter will not show light from an LED flashlight.

4. Adjust the OUTPUT fiber stage to get four even, round spots. The center-of-focus will have to be located and placed in the middle of the square arrangement of the fiber bundle.

5. Replace the mirror.

6. Perform a raster scan.

7. Make fine adjustments to the fiber stage while watching the counts in “photom”.

**B.7 OUTPUT**

Aligning the OUTPUT OAP is done in a similar manner to the ones at the fiber injection stages. However, due to its crowded location, the process is made slightly more difficult.

1. **Figure B.2** shows the alignment telescope positioned on the NIRO table so that the JouFLU output fiber bundle can be viewed.
Figure B.2: This photograph shows the placement of the CHARA laboratory alignment telescope on the CLASSIC beam combiner optical table. The telescope is oriented such that the fiber bundle can be viewed from the output OAP.

2. A flat mirror is placed so that the alignment telescope can be auto-collimated. See Figure B.3.

3. After auto-collimating the alignment telescope, remove the flat mirror.

4. Place a target, specifically made for this purpose, in place of the OUTPUT fold mirror.

   See Figure B.4. This target is used to adjust the translation of the alignment telescope.

5. Plug the red fiber laser in place of the JouFLU fiber bundle.

6. After adjusting the translation of the telescope, remove the target. Use the laser spot to
adjust rotation of the alignment telescope. Repeat these steps as necessary to align the telescope.

7. If needed, make XYZ translation adjustments to the fiber stage to get a small, bright, and round laser spot from the fiber bundle. This means that the fiber is at the focus of the OAP.

8. Replace the OUTPUT flat mirror.

9. The height and position of the beam should be checked. If the beam is straight, NICMOS is positioned so that the four spots hit the readout region.
Figure B.4: Finally, this photograph shows the OUTPUT flat mirror that redirects the light after the OAP removed and replaced with a target. This target is used to adjust the alignment telescope onto the beam axis.

B.8 FTS

To align the FTS, use the green laser and the red CHARA alignment laser in order to view the beam transmitted by the FTS Beam Splitter (BS) cube. The beam splitter works in K-band. The green and red lasers are co-aligned before passing through the beam splitter. The procedure is as follows:

1. Perform a laboratory alignment to the JouFLU table using the green laser.

2. Turn on the CHARA WL, insert the corner cubes in the beams you aligned on, and
perform a JouFLU raster scan. At this point you should have a completely aligned JouFLU system.

3. The FTS stage holds a fold mirror and the beam splitter. Move it out from the optical path using the JouFLU software.

4. Use the alignment lasers to center two targets in beam B.

5. Use the JouFLU software to move the FTS stage to the “FTS In” position.

6. To align the reflected beam, the beam splitter must first be translated so that the reflected beam reaches the center of the mirror. This step requires a lot of caution because the cube is extremely fragile.

7. After the precise translation of the cube, the path of the reflected beam has to be corrected by using two rotations. Rotate the beam splitter to align it on the first target. Then adjust the tip/tilt of the FTS fold mirror to align the beam on the second target. Iterate this procedure until it converges. After this step, the reflected beam is aligned on the two targets, and so with the table and the camera.

8. It may be useful to verify the alignment of the transmitted beam afterwards.