Annu. Rev. Astron. Astrophys. 2001. 39:353–401

OPTICAL INTERFEROMETRY

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Key Words astronomical instrumentation, astrometry, stellar parameters, binary stars, circumstellar matter

■ **Abstract** The field of optical and infrared (IR) interferometry has seen rapid technical and scientific progress over the past few years. A number of instruments capable of precise visibility measurements have been built, and closure-phase imaging with multitelescope arrays has been demonstrated. Astronomical results from these instruments include measurements of stellar diameters and their wavelength dependence, limb darkening, stellar surface structure, and distances of Cepheids and of Nova Cygni 1992. Precise stellar masses have been obtained from interferometric observations of spectroscopic binaries, and circumstellar disks and shells have been resolved. Searches for substellar companions and extrasolar planets with interferometric astrometry will begin soon. Nulling interferometry will enable studies of exozodiacal disks from the ground and the detection and characterization of terrestrial extrasolar planets from space. These developments are reviewed, as well as progress in some key technological areas.

1. INTRODUCTION

The technique of optical[1] interferometry is rapidly coming of age as a productive tool of astrophysics. Whereas much effort will still be needed to develop the technological basis of the advanced ground-based and space-based arrays that are expected to revolutionize fields such as astrometry and high-contrast high-resolution imaging, past and current instruments have produced a fair number of results on stellar diameters, limb darkening, stellar surface structure, binary stars, and circumstellar matter. The first aperture-synthesis images from Michelson instruments were obtained a few years ago, using phaseless reconstruction from single-baseline data (Quirrenbach et al. 1994a) and closure-phase imaging with multielement arrays shortly thereafter (Baldwin et al. 1996, Benson et al. 1997).

[1]The word "optical" is used in this article in a broad sense, including both the visible and IR spectral ranges.

The imminent commissioning of baselines longer than 100 m for observations at visible wavelengths will push the resolution into the submilliarcsecond regime, and the first fringes between giant telescopes—8 m for the European Southern Observatory's Very Large Telescope Interferometer (VLTI), 10 m in the case of the Keck Interferometer on Mauna Kea—have been obtained in early 2001. These large facilities will be open to guest observers, which will bring about a profound change in the way optical interferometry is practiced.

This article is intended to be a review of the technical and astronomical progress over the past few years in optical interferometry. This field was previously reviewed in this series by Shao & Colavita (1992b), who gave an excellent overview of the foundations and early history of interferometry, as well as the state of the art in the early 1990s. I have tried to concentrate on the developments that have occurred since that article was written. The material is organized in four sections, covering interferometer theory, basic technology, actual instruments, and astronomical results, respectively. The application of optical and IR interferometry to the detection and characterization of extrasolar planets has been reviewed recently by Marcy & Butler (1998) and Woolf & Angel (1998); these topics are therefore not covered here with a depth commensurate with their importance. In addition to the reviews already mentioned, the reader may wish to consult the following books and proceedings volumes for more information: the proceedings from the 1999 Michelson Summer School (Lawson 2000a) provide a tutorial overview; many significant papers have been collected by Lawson (1997); the SPIE conference volumes 3350 (Reasenberg 1998) and 4006 (Léna & Quirrenbach 2000) contain mostly (but not exclusively) material on technology and instrumentation; the conference "Science with the VLTI" (Paresce 1997) focused on astronomical results already obtained with, or expected from, ground-based interferometry; the proceedings from the "Working on the Fringe" workshop (Unwin & Stachnik 1999) contain papers on technology and instruments along with discussions of their scientific potential; the proceedings "Infrared Space Interferometry" (Eiroa et al. 1997) and the booklets on the Space Interferometry Mission and Terrestrial Planet Finder (Danner & Unwin 1999, Beichman et al. 1999) provide an overview of the goals and technological challenges of future space missions; and a wealth of information on optical fibers and planar waveguides can be found in "Integrated Optics for Astronomical Interferometry" (Kern & Malbet 1996).

2. INTERFEROMETER THEORY

2.1. Interferometric Observables

2.1.1. OBSERVABLES FOR IMAGING The primary observable in a Michelson interferometer is the complex visibility $\Gamma = Ve^{i\phi}$, which is related to the sky brightness distribution $B(\xi, \eta)$ through a Fourier transform (the van Cittert–Zernike Theorem) (see e.g., Shao & Colavita 1992b, Thompson et al. 1986):

$$\Gamma(u, v) = \iint B(\xi, \eta)\, e^{-2\pi i(u\xi + v\eta)} d\xi\, d\eta, \tag{1}$$

where ξ and η are the coordinates in the tangent plane of the sky, and u and v the two components of the baseline vector in the Fourier plane, measured in units of the observing wavelength. The visibility can be observed either as the fringe contrast in an image plane or by modulating the internal delay and detecting the consequent temporal variations of the intensity in a pupil plane. The phase of the visibility is frequently corrupted by atmospheric or internal fluctuations of the optical pathlength difference (OPD). Under these circumstances, it is usually better to work with the square of the visibility V^2 rather than V itself, because V^2 estimators can be constructed that take the noise bias into account properly (Tango & Twiss 1980, Shao et al. 1988, Colavita 1999). For delay modulation and synchronous detection of the photon count rate in n bins per wavelength of modulation, such an unbiased estimator is given by

$$V^2 = \left(\frac{4\pi^2}{n^2 \sin^2(\pi/n)}\right) \frac{\left\langle X^2 + Y^2 - \sigma_N^2 \right\rangle}{\langle N - D\rangle^2}, \tag{2}$$

where X and Y are the real and imaginary parts of the visibility constructed from the bin counts, N the total number of counts in all bins, σ_N^2 the variance of N owing to noise, and D the dark count measured separately (Benson et al. 1998). The analysis of data from single-baseline interferometers normally proceeds by fitting models to V^2 values after these have been properly calibrated. [The calibration procedure is a serious issue by itself (see e.g., Mozurkewich et al. 1991, von der Lühe & Quirrenbach 1995).] In the data from multibaseline instruments, useful phase information is preserved in the form of closure phases or triple products, which can be used together with the visibility amplitudes for image reconstruction purposes as in the radio case (e.g., Baldwin et al. 1996, Benson et al. 1997; see also Thompson et al. 1986). In contrast, closure amplitudes are normally not useful in optical interferometry, because the dominant source of amplitude errors are most frequently wavefront aberrations across the pupil, which are nonclosing errors.

It should be recalled here that the signal-to-noise ratio (SNR) of interferometric observables depends not only on the photon count N, but on NV^2 for the photon-noise limited regime and on NV for the background-limited regime, and that the SNR drops precipitously in the photon-starved regime (Shao & Colavita 1992b).

So far we have assumed that the phase is constant across each individual telescope. In a ground-based interferometer this is true only if the aperture D is sufficiently small or if adaptive optics is used to flatten the wavefront. If telescopes larger than the atmospheric coherence length r_0 are used in an interferometer, fringed speckles can be observed in short exposures. The power spectrum of such data contains information at low spatial frequencies ($\leq r_0/\lambda$ and $\leq D/\lambda$) owing to seeing and the diffraction limit of the individual telescopes, respectively, and at high spatial frequencies ($\sim B/\lambda$) owing to the fringes. V^2 can in this case be estimated from the ratio of the power in the fringe peak to that in the low-frequency peaks (Mourard et al. 1994b).

2.1.2. ASTROMETRIC OBSERVABLES The basic astrometric observable in an interferometer is the delay $d = d_{\rm int} + (\lambda/2\pi)\phi$, where $d_{\rm int}$ is the internal delay measured by a metrology system and ϕ is the observed fringe phase, which has to be unwrapped, i.e., not restricted to the interval $[0, 2\pi)$. Because d is related to the baseline $\vec{B}$ by

$$d = \vec{B} \cdot \vec{s}, \tag{3}$$

where $\vec{s}$ is a unit vector in the direction towards the star, the angle θ between $\vec{B}$ and $\vec{s}$ can be derived from d; this corresponds to a one-dimensional measurement of the source position. The second source coordinate can be determined by changing the baseline orientation. The astrometric accuracy σ is given by

$$\sigma = \frac{1}{\rm SNR} \cdot \frac{\lambda}{2\pi B}. \tag{4}$$

Because high SNR can be obtained for bright stars on long baselines, optical interferometry can provide extremely accurate astrometry, provided that the instrumental and atmospheric errors can be adequately controlled or calibrated (see Shao & Colavita 1992b).

Various differential schemes have been devised to overcome the limitations of astrometric accuracy imposed by atmospheric turbulence. Dual-star narrow-angle astrometry can be used to determine relative positions and motions with $\sim$10 μas precision (Shao & Colavita 1992a; see Section 5.8.2). If the position of the photocenter of an object depends on wavelength, the observed phase will also be wavelength-dependent according to Equation 3. This effect can be used to resolve structure on scales smaller than the resolution limit of the interferometer. For example, the photocenter of a star-planet system shifts with wavelength across molecular absorption bands in the planetary atmosphere; phase measurements with $\sim$0.1 milliradian accuracy could therefore be used to perform IR spectroscopy of "hot Jupiters" (Akeson & Swain 1999, Quirrenbach 2000a). Differential phase measurements with high spectral resolution across stellar absorption lines could also be used to determine angular diameter, equatorial rotation speed, and orientation of the stellar rotation axis, as well as size and location of starspots (Petrov 1989, Chelli & Petrov 1995).

2.1.3. DOUBLE-FOURIER INTERFEROMETRY In Section 2.1.1 we have tacitly assumed that fringe measurements are made at zero OPD. It is also possible to scan through the whole fringe packet and measure the complex degree of coherence $\Gamma(u, v, \tau)$ as a function of delay τ; this method is known as double-Fourier interferometry (Mariotti & Ridgway 1988, Mékarnia & Gay 1990). In generalization of

Equation 1, one can write

$$\Gamma(u, v, \tau) = \iiint B(\xi, \eta, \nu) e^{-2\pi i(u\xi + v\eta + \tau\nu)} d\xi\, d\eta\, d\nu, \tag{5}$$

which shows that the brightness B as a function of position and frequency ν can be obtained from the interferogram. The special case $u = 0, v = 0$ corresponds to a classical Fourier transform spectrometer (FTS). As in the FTS, the spectral resolution depends on the length of the OPD scan. It should be pointed out that for a given baseline length, u and v scale with wavenumber, so that each interferogram provides data on a diagonal in (u, v, ν) space. Atmospheric phase variations during the OPD scan corrupt the interferogram, but they can be measured simultaneously in a separate channel, so that a real-time or off-line correction is possible.

2.2. Beam Combination and Field of View

2.2.1. BEAM COMBINATION CONCEPTS The various beam combination schemes that can be employed in astronomical interferometry may be classified according to several criteria (Eur. South. Obs. 1992): the beam étendue (single-mode or multimode), the beam direction (co-axial or multiaxial), the combination plane (image plane or pupil plane), and the relation between input and output pupils (Michelson or Fizeau configuration, see below). For N telescopes in an array, there are $N(N-1)/2$ baselines. The $N(N-1)/2$ visibilities can be measured either by pairwise beam combination or by bringing the light from all telescopes together on one detector. In the latter "all-on-one" techniques the fringes from the different baselines have to be encoded either spatially (by using a nonredundant output pupil) or temporally (by using different dither frequencies for the beams from individual telescopes).

Unlike in radio astronomy, where the radiation is detected and amplified before correlation, in an optical interferometer the beam combination occurs before detection. For pairwise beam combination the light from each telescope has to be divided in $(N-1)$ beams; in all-on-one schemes the visibility measurement for each baseline is affected by noise contributed by the $(N-2)$ other telescopes. This means a baseline that is part of an N-element array is always less sensitive than an equivalent two-telescope interferometer. The detailed trade-offs between different beam combination schemes depend on the predominant noise source (background, detector, photon noise), detector cost and availability, and other technical considerations. Armstrong et al. (1998) discussed the case of pupil-plane combination with temporal encoding of the fringes, for which in the photon-rich regime

$$\mathrm{SNR} \propto \left(\frac{n_{\mathrm{phot}} N}{N_{\mathrm{out}}}\right)^{1/2} \frac{V}{N_{\mathrm{corr}}}, \tag{6}$$

where n_{phot} is the photon rate from each telescope, N the number of array elements (equal to the number of input beams to the combiner), N_{out} the number of output beams from the combiner, and N_{corr} the number of input beams combined to produce each output beam. For pairwise combination $N_{\rm corr} = 2$ and $N_{\rm out} = N(N-1)$. (Note that combining beams at beamsplitting surfaces produces two output beams.) For all-on-one combination $N_{\rm corr} = N$ and $N_{\rm out} = 2$. In both cases SNR $\propto N^{-1/2}$, which demonstrates that multielement arrays are indeed less sensitive than single-baseline instruments. Equation 6 gives a $\sim\sqrt{2}$ advantage of the all-on-one technique over pairwise beam combination, but the required temporal encoding of the fringes is difficult to realize technically.

2.2.2. MICHELSON AND FIZEAU INTERFEROMETERS In a Fizeau interferometer the output pupil is an exact replica of the input pupil, scaled only by a constant factor. This is also known as homothetic mapping between input and output pupil. In contrast, in a Michelson interferometer there is no homothetic relation between the input and output pupils.[2] This means the object-image relationship can no longer be described as a convolution, because the rearrangement of the apertures rearranges the high-spatial frequency part of the object spectrum in the Fourier plane (Tallon & Tallon-Bosc 1992). This has an important consequence for off-axis objects: The image position does not coincide with the white-light fringe position (see Figure 1 in Tallon & Tallon-Bosc 1992). For a finite spectral bandwidth this means the fringe contrast decreases with field angle and the field of view is limited; the maximum size of an image from a Michelson interferometer is $\sim R \equiv \lambda/\Delta\lambda$ resolution elements in diameter. This effect is known as bandwidth smearing in radio astronomy (Thompson et al. 1986). If a Michelson interferometer is used with image plane beam combination, and the visibilities are estimated by integration over each fringe peak, the field of view is additionally restricted to the size of one Airy disk of the individual telescopes.

2.2.3. THE DENSIFIED PUPIL INTERFEROMETER It is interesting to consider interferometers in which the output pupil is not a homothetic image of the input pupil, but the pattern of the subaperture centers is conserved (Labeyrie 1996). Such an arrangement has a limited field of view, but the high-spatial-frequency content of the object spectrum remains intact, i.e., the interferometer forms a true image of small sources at the combined focus. This is particularly attractive for arrays consisting of many telescopes, because the desired information is obtained directly (or after simple deconvolution with the known point spread function), rather than indirectly through the measurement of $N(N-1)/2$ visibilities. In the output pupil, the subapertures can almost touch each other. Such a "densified pupil interferometer" has a substantial signal-to-noise advantage over a Fizeau instrument (Roddier & Ridgway 1999). It is therefore possible to trade field of view for sensitivity in interferometric arrays if direct imaging is required.

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[2] Sometimes the terms Fizeau and Michelson are also used to mean "image plane" and "pupil plane" interferometer, respectively. In my nomenclature it is possible to build an image plane Michelson interferometer.

2.2.4. ASTROMETRY WITH FINITE FIELD OF VIEW For imaging applications fringe contrast at off-axis angles is the most important parameter that determines the usable field of view. For astrometric applications, distortions that cannot be calibrated are of equal or even larger importance (Quirrenbach et al. 1998). In dual-star interferometers such as the Palomar Testbed Interferometer (PTI) and the Keck Interferometer (see Section 4), the star separator ("dual-star module") splits the image plane and sends the light from two stars on separate paths to two different beam combiners. The lengths of these "on-axis" paths are measured by the metrology system. Light arriving from a nonzero field angle in one of the two fields travels down the delay lines off-axis and has a different "footprint" on mirrors that are not in a pupil plane. A 5 nm goal on the OPD accuracy therefore implies that surface imperfections on these mirrors have to be calibrated with that same precision. A particular problem for the VLTI in this respect is the variable curvature mirror (see Section 3.5), which is located in an image plane at the focal position of the delay line trolley's Cassegrain telescope. The variable curvature mirror allows for proper pupil relay, but any slight deviations from its nominal shape cause field distortions, which are problematic for astrometric observations over a "wide" ($\sim 1''$) field.

2.3. Fringe Detection and Sensitivity

2.3.1. COPHASING AND PHASE REFERENCING Fringe detection in optical interferometry can be separated into two distinct tasks: fringe tracking, i.e., measuring the fringe phase and driving a servo loop that keeps the interferometer cophased, and obtaining the visibility measurements at a desired wavelength and on a specific baseline. Whereas it is possible to do "passive" interferometry, i.e., measuring visibilities or triple products without real-time fringe tracking, "active" interferometry gives higher signal-to-noise and makes calibration easier (Shao & Colavita 1992b). Passive interferometry can potentially allow reaching better magnitude limits, but the SNR will normally remain poor.

Usually the sensitivity of an active interferometer is limited by the fringe-tracking, which has to be done with integration times that are short compared with the atmospheric coherence time. Once the interferometer is cophased, the integration time for data-taking may be arbitrarily long; the point-source sensitivity is then comparable to the sensitivity of a single telescope with the same aperture (apart from the notoriously low transmission due to the many mirrors in interferometers). The cophasing can also be done off-line in the data reduction. If the visibilities in the tracking and data channels are recorded together, the phase information can be transferred from the tracking to the data channel, so that the corrected visibilities in the data channel can be integrated coherently (Colavita 1992, Quirrenbach et al. 1994b, Meisner 2000). This phase-referencing technique can substantially increase the SNR in the data channel, but all processes that may lead to a phase decorrelation between the two channels have to be understood thoroughly.

2.3.2. WAVELENGTH AND BASELINE BOOTSTRAPPING It is important to realize that fringe tracking does not necessarily have to be performed at the same wavelength, on the same baseline, or even on the same star as the observation. For example, the Mark III (Mk III, discussed in Section 4.2) used a wide-band channel to track the fringes and narrower channels to take data (Shao et al. 1988). This has the dual advantages of optimizing the fringe-tracking sensitivity and of having a relatively narrow envelope of the fringe packet in the tracking channel, which helps the fringe tracker identify the central white-light fringe.

Another common task is the need to take data at high spatial frequencies, where V^2 may be too low for fringe tracking. If V^2 and therefore the SNR is high at a wavelength λ_1 but low at another wavelength λ_2, it is possible to take data at λ_2 with the fringe tracker working at λ_1. The spatial frequency leverage afforded by this wavelength bootstrapping technique is useful, for instance, for imaging stellar photospheres, where λ_1 can be in the IR and λ_2 in the visible, or for observations of circumstellar matter, where λ_1 can be in the continuum (where the small stellar photosphere dominates) and λ_2 in a line emitted by the extended material. In arrays with more than two telescopes it is possible to employ baseline bootstrapping, which uses the signal on the short baselines for the fringe tracking servo, while data are taken on the long baselines. The array geometry of the Navy Prototype Interferometer (NPOI) has been specifically optimized for baseline bootstrapping (Armstrong et al. 1998). A related idea has been developed for arrays with telescopes of different sizes, such as the VLTI and the Keck Interferometer. Fringe tracking is required only on the more sensitive baselines involving at least one large telescope, while bootstrapping enables observations also on the baselines between two small telescopes. Whereas these techniques rely on tracking the fringes on a baseline-by-baseline basis, it might be interesting to explore global fringe detection methods that consider all baselines simultaneously, as is commonly done in radio interferometry (e.g., Rogers et al. 1995).

2.3.3. DISPERSED FRINGES AND GROUP DELAY TRACKING Spectrally dispersing the output of a two-element pupil plane interferometer produces the so-called channeled spectrum

$$I(\nu) = I_0(\nu)(1 + \Gamma(\nu)e^{-2\pi i\tau\nu}), \tag{7}$$

where $I_0(\nu)$ is the stellar spectrum as a function of frequency and τ is the total delay. From Equation 7 we see that fringes with period $\Delta\nu = 1/\tau$ appear in the channeled spectrum, provided that I_0 and Γ do not vary strongly with wavelength. If the dispersion of the spectrometer produces a linear mapping of wavenumber to the detector pixels, a fast Fourier transform of the channeled spectrum yields the delay (e.g., Lawson 1995). This technique suffers from the fact that the sign of the delay is not measured, which complicates the implementation of a zero-seeking servo loop. This problem can be fixed by taking a second measurement at a known delay offset $\Delta\tau$ and observing whether the fringe period increases or decreases. A generalization of this idea is the combination of spectral dispersion with delay dithering, i.e., recording the photon counts as a two-dimensional function of ν and τ, or of wavenumber and modulation bin. The delay can be found with a two-dimensional Fourier transform of this array, provided that the detector electronics are set such that the bin lengths are proportional to the wavelength (Benson et al. 1998). This method estimates the variation of phase with frequency, i.e., the group delay

$$\tau_g \equiv \frac{1}{2\pi}\frac{\partial\phi}{\partial\nu}. \tag{8}$$

This is a robust way to find and track the white-light fringe, because it provides a straightforward error signal for the fringe servo and because the large coherence length (inversely proportional to the width of the individual spectral channels) is helpful for the initial acquisition of the fringe. Because the signal is distributed over a potentially large number of spectral channels, the use of essentially noiseless detectors such as avalanche photo diodes (APDs) is highly advantageous, and read noise is a critical parameter if CCDs are to be used (ten Brummelaar 1997).

It is possible to use the group delay for astrometric measurements, but a comparison of the variance of the group delay σ_g^2 with the variance of the phase delay σ_ϕ^2 shows that

$$\sigma_g^2 = 12\left(\frac{\lambda_c}{\Delta\lambda}\right)^2\sigma_\phi^2, \tag{9}$$

where λ_c and $\Delta\lambda$ are the central wavelength and total bandwidth covered by the spectrograph, respectively (Thompson et al. 1986). For the atmospheric K-band ($\lambda_c = 2.2\,\mu$m, $\Delta\lambda = 0.4\,\mu$m) this means that the rms noise is 19 times smaller if the delay is estimated from phases rather than from the group delay.

For the performance of a fringe tracker the probability of tracking loss is an even more important metric than the delay variance. The tracking fails when a noise peak is higher than the fringe peak, either because the light level is too low or because the fringe visibility is too small. For interferometers with apertures comparable to r_0, it is therefore important to consider not only photon and detector noise, but also fluctuations of the visibility due to seeing variations (Lawson et al. 1999). The correct identification of the fringe peak in the presence of noise is more difficult if the search space is larger (Thompson et al. 1986); it is therefore advantageous to include a priori knowledge about the fringe motion in the delay estimation algorithm (Morel & Koechlin 1998, Padilla et al. 1998).

If the telescopes operate in the multispeckle regime, the two-dimensional speckle pattern is usually recorded in the image plane. Observations of dispersed fringes require use of a slit with a width comparable to the size of individual speckles. This uses only a fraction $\sim r_0/D$ of the light, unless an image slicer is employed. Across each speckle the spatial variations of the phase are small; dispersion perpendicular to the slit therefore gives a channeled spectrum for each speckle, which can be used to estimate the delay as described above. Because the phase is not coherent between speckles, they must be processed individually; the individual delay estimates can then be averaged (Koechlin et al. 1996, Berio et al. 1999a).

2.4. Optical Synthesis Imaging

Because the complex visibility is related to the sky brightness distribution by a Fourier relation (Equation 1), it is in principle possible to obtain images from interferometer data by an inverse Fourier transform. The principal difficulties associated with this task are incomplete coverage of the *uv* plane and corruption of the phases by atmospheric fluctuations and instrumental effects. These are well known from radio interferometry, and the methods and software developed for radio synthesis imaging (Thompson et al. 1986, Perley et al. 1989) can to a large extent also be applied at optical and IR wavelengths (Baldwin et al. 1996, Benson et al. 1997, Tuthill et al. 1999b, 2000b). Because of the small number of elements in current optical synthesis arrays, Earth (or spacecraft) rotation synthesis is important to achieve reasonable *uv* coverage. This technique can frequently be combined with reconfigurations of the array, because several instruments (Cambridge Optical Aperture Synthesis Telescope, NPOI, VLTI) use movable telescopes. However, variations of the source structure due to intrinsic variability or stellar rotation limit the time available for the aperture synthesis.

The combination of data taken in multiple spectral channels can further increase the *uv* plane coverage, because the spatial frequency is inversely proportional to the wavelength. This technique obviously works best if the source structure is independent of wavelength, as for example in binary systems with two components of equal spectral type (Benson et al. 1997). A generalization is possible if the wavelength dependence can be easily parameterized, such as a spotted stellar surface, which can be described by brightness and temperature for each point.

There are a number of important differences between optical and radio synthesis arrays, however. The decrease of SNR with increasing number of elements (Equation 6) favors combining light from only a few telescopes at a time and switching between subarrays for the full *uv* synthesis. Cophasing the array is more challenging in the visible, which drives the layout towards optimizing bootstrapping (see Section 2.3.2) rather than instantaneous *uv* coverage. Finally, the observables in optical and radio interferometry are different (V^2 versus V) (see Section 2.1.1). The current practice of using data analysis packages developed for radio interferometry with optical data is therefore not optimal, as this requires assigning $\langle V^2\rangle^{1/2}$ to the visibility amplitude and prevents statistically proper use of the closure phases. For imaging of faint sources, an algorithm would be needed that uses all information in the bispectrum phasors (Kulkarni et al. 1991).

Many properties of an interferometer—aperture size, geometric configuration, wavefront quality, beam combination technique, detector characteristics, data analysis methods—contribute to its imaging performance, although an analytical computation suggests that the noise in optical synthesis images depends only weakly on the beam combination scheme chosen (Prasad & Kulkarni 1989). Nevertheless, the assessment of image fidelity requires realistic simulations that include the effects of noise, incomplete *uv* coverage, and image restoration algorithms. Simulated observations of Hα emission from a rotating disk around the black hole at the center of M87 and of the β Pic disk have been used to demonstrate the imaging capabilities of the Space Interferometry Mission (Böker & Allen 1999). The most complete modeling tools perform "end-to-end" simulations of the interferometer. They use dynamic optomechanical models of telescopes, delay lines, beam combination optics, and detectors to compute realistic time-dependent interferograms (Schöller et al. 2000). End-to-end models that include the effects of atmospheric turbulence and adaptive optics will be an invaluable tool for the design of future synthesis arrays with large apertures.

2.5. Atmospheric Turbulence

The sensitivity of ground-based interferometers depends strongly on the properties of atmospheric turbulence, because the coherence length r_0 determines the maximum aperture size that can be used without adaptive optics, and the maximum integration time for fringe tracking is limited by the coherence time τ_0 (Buscher 1988). Scintillation noise can bias visibility measurements (Tango 1998), although this effect is experimentally found to be small (Quirrenbach et al. 1994b). The standard Kolmogorov-Taylor model (for a brief summary, see e.g., Roddier 1989 or Quirrenbach 2000c) provides the framework in which seeing measurements are usually interpreted. Potential deviations from this model are important, however: an outer scale of the turbulence $L_0 \lesssim B$ would reduce the fringe motion dramatically and give much-reduced astrometric errors (Shao & Colavita 1992a), and a non-Kolmogorov power-law slope of the turbulence spectrum would affect the wavelength scaling of r_0 and τ_0.

Measurements of L_0 have been undertaken with various techniques. An example is the campaign on Cerro Paranal carried out with the Generalized Seeing Monitor (GSM), from which a mean value $L_0 = 24$ m was derived for a von Karman spectrum (Conan et al. 2000). Although such data are extremely useful, one may feel a bit uneasy about values of L_0 inferred by extrapolations from measurements at small spacings ($\sim$1 m in the case of the Generalized Seeing Monitor), which are necessarily model dependent. It is therefore important to note that measurements of pathlength fluctuations with different baselines of the Mk III ($B = 12$ m and $B = 31.5$ m) and the Sydney University Stellar Interferometer ($B = 5$ m, 20 m and 80 m) also support an outer scale in the range 10 m to 100 m (Buscher et al. 1995, Davis et al. 1995).

In the regime $t \ll B/V$, where V is the wind speed, the Kolmogorov model predicts a delay structure function $D_\tau(t) \propto t^\beta$ with $\beta = 5/3$, corresponding to a power spectrum $P \propto \nu^{-8/3}$. Measurements with the Sydney University Stellar Interferometer (Davis et al. 1995) and an extensive set of Mk III data (Buscher et al. 1995) support power law slopes very close to the Kolmogorov value. In contrast, the Infrared Spatial Interferometer (ISI) data suggest a sub-Kolmogorov slope (Bester et al. 1992), and Linfield et al. (1999) found $\beta \approx 1.3$ from observations with PTI. Because $r_0 \propto \lambda^{2/\beta}$, this would imply a much larger difference between the visible and IR seeing than usually assumed. Resolving the discrepancy between the data from the different instruments and gaining a more accurate understanding of the turbulence spectrum on scales of tens and hundreds of meters will require simultaneous measurements with a larger range of baseline lengths.

2.6. Polarization

The theory of optical interferometry sketched in the preceding sections ignores the vector properties of the electric radiation field. This approach is valid only if the source is not highly polarized and if the individual arms of the interferometer have identical polarization characteristics. Any violations of the latter condition, i.e., differential field rotation between the interferometer arms or phase shifts between the two polarization states, lead to a reduction of the fringe contrast (e.g., Eur. South. Obs. 1989). Because reflections at nonnormal incidence angles lead to phase shifts between the *s* and *p* polarization states, optical interferometers are usually designed such that the number of reflections in each arm is the same, that they occur at identical angles of incidence, and that the same coatings are used on corresponding surfaces. The loss of coherence due to misalignment of optical elements, aging of coatings, and accumulation of dust can be analyzed with a full vector treatment of light propagation (Elias 2001).

The requirement of having arms with matched reflections can in practice only be met in array configurations with just a few preferred directions of light propagation. This can relatively easily be achieved in Y-shaped or T-shaped arrays. If the telescopes are located at "random" positions as in the VLTI or the Keck Interferometer, the light pipes must still be laid out such that they follow a few fixed directions. Matched layouts also require a fairly large number of reflecting surfaces, which reduces the throughput. It has therefore been suggested that configurations with unmatched reflections—such as a circular arrangement of the telescopes with radial light propagation to a central beam combiner and a standard field-derotator in each arm—might be advantageous (Rousselet-Perraut et al. 1997a). The visibility loss due to the different incidence angles on corresponding optical elements in each arm would be tolerable at least over part of the sky, so that the advantage in throughput gives a higher SNR. The downside of this approach, which accepts large polarization losses, is likely an increased sensitivity to additional "real world" effects (aging of coatings, dust etc.), which may make a reliable calibration impossible.

Separate detection of the two polarization states can be used to eliminate (for pairwise beam combination) or at least reduce (for multiway beam combination) the loss of fringe contrast due to phase shifts between the two polarizations. The relative phase between the two linear polarization states can be controlled with a rotating half-wave plate sandwiched between two quarter-wave plates; this effect can be used to implement an achromatic phase modulator (Tango & Davis 1996). Furthermore, interferometric polarimetry may in the future be used to obtain high-resolution information on polarized objects, e.g., Be stars and magnetic stars (Rousselet-Perraut et al. 1997b, 2000).

3. INTERFEROMETER TECHNOLOGY

3.1. Optical Fibers and Spatial Filtering

Single-mode optical fibers can be used for many of the functions required in an optical interferometer: beam transport, beam combination (in **X** couplers), OPD modulation (by physically stretching a fiber) (Shaklan 1990), polarization control, and spatial filtering. The last capability is particularly attractive; single-mode fibers can eliminate the decrease in fringe visibility caused by atmospheric turbulence and thus alleviate the calibration difficulties of ground-based interferometers. Because the coupling efficiency into single-mode fibers depends on the wavefront shape (Shaklan & Roddier 1988), changes of atmospheric wavefront aberrations are converted into intensity fluctuations. When the light from two telescopes is combined in an **X** coupler, the observed interferogram I is therefore given by (Coudé du Foresto et al. 1997):

$$I = P_1 + P_2 + 2\sqrt{P_1 P_2}\Gamma e^{i\phi}, \tag{10}$$

where P_1 and P_2 are the intensities in the two input fibers, Γ the complex visibility, and ϕ the internal phase. Splitting off some of the light from each telescope in a **Y** coupler before beam combination allows monitoring of P_1 and P_2. The corrected interferogram

$$I_{\text{cor}} \equiv \frac{I - P_1 - P_2}{2\sqrt{P_1 P_2}} = \Gamma e^{i\phi} \tag{11}$$

is then independent of atmospheric wavefront degradation. By definition, the étendue of a single-mode fiber is λ^2, i.e., the field of view is limited to one Airy disk. The case in which the source is partially resolved by the individual telescopes has been analyzed by Dyer & Christensen (1999). Spatial filtering with single-mode fibers is also a promising technique for producing the ultraflat wavefronts required for high-contrast nulling (Ollivier & Mariotti 1997, Mennesson et al. 2001).

Dispersion is a serious concern for the use of optical fibers in interferometry, because it requires very careful matching of the fiber lengths for wide-band operation (Shaklan & Roddier 1987, Rohloff & Leinert 1991). For example, the lengths of standard silica fibers have to be equalized to $\lesssim$50 μm for use over the visible wavelength range (0.4 μm $\leq \lambda \leq$ 0.7 μm). This requirement can be relaxed by optimizing the waveguide structure of the fiber, so that the material dispersion and the waveguide dispersion cancel each other (Coudé du Foresto et al. 1995). For applications at longer wavelengths, transmission also becomes an important issue. A fluoride glass fiber coupler has been used successfully for L band (3.75 μm) observations (Mennesson et al. 1999), and a 1 mm-long fiber of the same material has shown 40% transmission at 10 μm, but the transparency becomes negligible at 12 μm (Perrin et al. 2000). Spatial filtering at mid-IR wavelengths calls for the development of single-mode fibers from new materials, such as chalcogenide glasses or halides.

3.2. Integrated Optics

An alternative technology to the use of optical fibers is integrated optics on a planar substrate (Kern et al. 2000). Single-mode planar waveguides can be produced by etching (Mottier 1996) or ion exchange (Schanen-Duport et al. 1996) processes on glass or silicon substrates. This technology is well developed for applications at visible wavelengths and in the near-IR, where transmission losses of 0.1 dB/cm or better can be achieved; in the thermal IR new materials such as chalcogenide glasses will be needed (Laurent et al. 2000). It is possible to implement elements such as straight and curved waveguides, direct and reverse **Y** junctions, **X** junctions, directional couplers, mirrors, and OPD modulators; this creates the potential to integrate the functions of fringe tracking, polarization control, spatial filtering, and beam combination on a single chip (Malbet et al. 1999). A prototype two-telescope beam combiner implemented on a 5 mm $\times$ 40 mm glass substrate has been tested in the laboratory; a transmission of 43% (dominated by a 50% loss in the **Y** coupler, which keeps only one interferometric output) and fringe contrast of 92% were obtained with a bandpass similar to the astronomical *H* band in a single polarization (Berger et al. 1999, Haguenauer et al. 2000a). Similar results have recently been achieved for three-telescope beam combiners (Haguenauer et al. 2000b), and designs for combining the light from up to eight telescopes on one chip are already under consideration (Berger et al. 2000a). First fringes on stars with the IONIC integrated optics beam combiner were obtained at the Infrared Optical Telescope Array (IOTA) in November 2000 (J.P. Berger, P. Haguenauer, P. Kern, F. Malbet, R. Millan-Gabet, K. Perraut, personal communication).

3.3. Nulling Interferometry

3.3.1. NULLING INTERFEROMETER CONCEPTS The introduction of an achromatic 180° phase shift in one arm of a two-element Michelson interferometer produces a dark central fringe. This technique, nulling interferometry, holds great promise for the detection and characterization of Earth-like extrasolar planets at mid-IR wavelengths, because the light from the parent star arriving on axis is completely rejected (Bracewell 1978, Léger et al. 1996, Beichman et al. 1999, Woolf & Angel 1998). In a two-element nulling interferometer the transmission of light arriving at a small off-axis angle θ is $\propto\theta^2$. The "leakage" of 10 μm flux from the edge of the stellar disk is thus ~100 times brighter than an Earth-like planet, for a baseline length optimized for the detection of planets in ~1 AU orbits. This problem can be solved by combining the output of three or more telescopes. Possible geometries include linear arrays of four telescopes producing a θ^6 null (Angel & Woolf 1997) and elliptical configurations of five telescopes, which give only a θ^4 null and require 72° phase shifts, but provide near-uniform sensitivity over a fairly large range of spatial frequencies (Mennesson & Mariotti 1997). The latter arrangement, which does not have a mirror symmetry, also enables the discrimination of a planet from a symmetric exozodiacal dust cloud. In these concepts the temporal modulation of the output, required in any Michelson interferometer, is produced by rotation of the array around the optical axis. This is somewhat problematic because the instrument has to be stable on time scales of hours. It has therefore been proposed to combine the output of two nulling arrays with an adjustable phase delay, so that fast internal modulation of the fringes is possible (Woolf & Angel 1997, Mennesson & Léger 2001).

The ideal mid-IR nulling array should satisfy the following conditions:

- All telescopes are at equal distance from the beam combiner, so that compensation of large pathlength differences is not required.
- All telescopes and the beam combiner are in one plane (perpendicular to the optical axis); this simplifies the design of thermal shields for radiative cooling.
- Starlight suppression is $\propto \theta^4$ or $\propto \theta^6$.
- The only applied phase shift is 180°, which is easier to control than shifts by arbitrary phase angles.
- All telescopes have the same diameter.
- Internal phase modulation is possible.

Several hexagon and pentagon configurations (with the beam combiner at the center) that fulfill these conditions do actually exist (Karlsson & Mennesson 2000). The beam combination scheme for most of these arrays is quite complicated and requires asymmetric beam splitters, but a hexagonal arrangement that uses only 50/50 beam splitters has also been found.

An alternative to the nulling Michelson interferometer is the use of a transparent phase-shifting mask in conjunction with a densified pupil interferometer (Boccaletti et al. 2000, Guyon & Roddier 2000). The basic idea of this concept is to introduce a 180° phase shift over the central part of the point spread function produced by the densified pupil; careful adjustment of the size of the phase mask together with pupil apodization in a coronographic setup can in principle provide excellent nulling performance. It is difficult to achieve an achromatic 180° phase shift over a wide wavelength band, however, and the scaling of the point spread function size with wavelength introduces additional complications for wide-band operation. If these can be overcome, interferometric nulling coronography may offer a sensitivity advantage over Bracewell nulling, because the former method produces an image on a multipixel detector so that the photons from planets fall on different detector pixels than most of the background emission by an exozodiacal disk or confusing background sources.

3.3.2. NULLING TECHNOLOGY Three different ways of introducing an achromatic phase shift have been proposed: rotational shearing with rooftop mirrors (Shao &

Colavita 1992b), passage through focus in one interferometer arm (Gay & Rabbia 1996), and phase retardation by dielectric plates with carefully chosen properties (Morgan et al. 2000). The first two methods allow only a 180° phase shift, whereas the third can produce any desired phase angle. For the study of extrasolar planets in the mid-IR a very deep ($\sim 10^{-6}$) null, i.e., a fringe contrast of 99.9999%, and excellent control of the phase are required. This means that instrumental errors such as amplitude imbalance between the two beams, polarization effects, field rotation, phase fluctuations, and wavefront aberration have to be carefully controlled (Serabyn 2000).

The rooftop-mirror nulling scheme uses the beam splitter in double pass so that the output beams are perfectly balanced even if the beam splitter deviates from the 50/50 ratio (Shao & Colavita 1992b). Furthermore, internal dithering of one of the rooftops produces an opposite phase in the two nulled outputs. By modulating one of the inputs simultaneously with the internal dither it is thus possible to dither one of the outputs around zero while maintaining an achromatic null in the other output; this property can be exploited to produce an error signal for a nulling servo loop (Serabyn 1999). In a laboratory demonstration experiment based on this principle, null depths of $\sim 10^{-5}$ for laser light (Serabyn et al. 1999) and $\sim 10^{-4}$ for broadband (18% bandpass) visible light (Wallace et al. 2000) have been achieved. In the thermal IR, 10^{-3} rejection of 9.6 μm laser light has been demonstrated, but in a setup that does not include an achromatic phase shifter and can thus not easily be extended to broadband operation (Ollivier et al. 2000).

First demonstrations of nulling on the sky have been carried out at 10.3 μm at the Multiple Mirror Telescope (Hinz et al. 1998) and at 2.2 μm at the 1.52 m telescope of the Observatoire de Haute Provence, equipped with an 88-actuator adaptive optics system (Baudoz et al. 2000a, 2000b). The performance of these experiments depended strongly on the seeing conditions; under the best conditions both achieved a ~ 0.06 null, as measured by the total transmitted flux of an unresolved star. The successful use of nulling techniques from the ground will clearly depend on excellent performance of adaptive optics and pathlength control systems.

3.4. Delay Lines and Metrology

The development of precise computer-controlled delay lines has arguably been the most important technical breakthrough for the success of modern optical interferometry (Shao & Colavita 1992b). The successful principle of the Mk III delay lines (Shao et al. 1988), which used laser interferometers to measure the positions of the delay line carts and nested servo loops for fine control of the OPD, has been widely used in the more recent instruments (Armstrong et al. 1998, Colavita et al. 1999, Colavita & Wizinowich 2000, Derie 2000, Hogenhuis et al. 2000, ten Brummelaar et al. 2000).

One major difference that sets the Mk III, NPOI, and IOTA apart from the other ground-based instruments is the implementation of vacuum delay lines. This is expensive but necessary for two-color astrometry (Colavita et al. 1987) and for observations on long baselines spanning a wide wavelength range in the visible (Benson et al. 1997). The alternative, compensation of the OPD occurring in vacuo with an air delay line, leads to longitudinal dispersion, i.e., a phase shift with wavelength. The corresponding loss of fringe contrast limits the usable bandwidth to $\lesssim$30 nm at 950 nm for a 100 m baseline, and even less for shorter wavelengths or longer baselines (Lawson 1996). Air delay lines are therefore usually implemented in conjunction with a dispersion corrector, which can be built from movable wedges of crown and flint glass (Tango 1990). Solutions with a single glass are also possible (ten Brummelaar 1995). The dispersion corrector can be optimized to provide both longitudinal dispersion correction at the observing wavelength, and linear mapping of fringes to detector pixels in a channeled spectrum; this simplifies the implementation of a group-delay fringe tracker (Lawson & Davis 1996, Davis et al. 1998).

Whereas the metrology systems necessary for the control of delay lines in imaging interferometers can be considered fairly routine affairs today, the requirements of astrometric instruments remain extremely challenging. Two types of metrology systems can be distinguished: internal (or constant-term) metrology measuring the internal path from the telescope to the beam combiner and global metrology systems that monitor the three-dimensional geometry of the interferometer with respect to bedrock (NPOI) (Hutter et al. 1998) or to guide interferometers locked on stars [the Space Interferometry Mission (SIM)] (Milman & Turyshev 2000). An important lesson from the SIM technology development program is that building three-dimensional "optical trusses" is far more difficult than just putting a number of one-dimensional metrology systems together; the optical surface quality of elements such as cat's eye retroreflectors, thermal stability, coupling to spacecraft vibrations, and control of a large number of degrees of freedom all have to be considered (Gürsel 1998, Laskin & Yu 2000).

Internal metrology is normally implemented in a double-pass setup by sending a laser beam backwards through the beam combiner along the optical path to a corner cube mounted at the telescope. The main challenge in meeting the exquisite precision requirements ($\sim$5 nm for differential astrometry on the ground, $\sim$50 pm for SIM) are noncommon path effects: The laser beam probes only a small patch of the optical surface, so that beam walk combined with imperfect surface figuring leads to metrology errors. The alternative, full-aperture metrology is almost impossible to implement because a large beam splitter would have to be used as the retroreflecting element. Dispersion effects between the wavelength of the metrology laser and the observing wavelengths in the long air delay lines of ground-based interferometers also have to be controlled precisely. This difficulty can be alleviated by multiplexing light from the target and reference star (e.g., by splitting and multiplexing the two polarization states) and sending light from both stars through the delay lines on a common path (Quirrenbach et al. 1998).

3.5. Pupil Transfer

Obtaining a "wide" interferometric field of view (i.e., a few arcseconds) requires accurate relay of the telescope pupils to the beam combiner; the variation of the distance from the telescope to the beam combination laboratory as the delay lines slew back and forth has to be compensated by a "zoom" capability (Ferrari & Derie 1998). In the VLTI, correct transfer of the pupils is achieved by variable curvature mirrors used as the tertiary mirrors for the delay lines' cat's eye (Ferrari et al. 2000). The variable curvature mirrors are thin (300 μm central thickness) stainless steel membranes that form part of a sealed chamber, to which a pressure of up to 10 bars can be applied. Varying the pressure from 0 to 8 bars changes the radius of curvature from 2800 mm to 84 mm corresponding to a change from f/∞ to $f/2.625$, with little hysteresis (Lemaître et al. 2000). Continuous control of the pressure during the observations will ensure the relay of the pupil to a fixed position on the instrument tables. In addition to this longitudinal relay, control of the lateral pupil position, pupil rotation, and beam demagnification is also required (Eur. South. Obs. 1989). The alignment tolerances for astrometry in a $\sim 1''$ field are even more stringent than for wide field imaging (Quirrenbach et al. 1998).

4. LONG-BASELINE OPTICAL AND INFRARED INTERFEROMETERS

4.1. Historical Instruments

Following Fizeau's (1868) suggestion that stellar diameters could be measured interferometrically, Stéphan (1874) carried out such an experiment and found that stars were unresolved by his 80 cm reflector. Michelson (1890) seems to have invented stellar interferometry independently, although it is possible that he might have been influenced by Fizeau (see Lawson 2000b). In any case, Michelson and collaborators are usually given credit for having performed the first useful measurements of a stellar diameter (Michelson & Pease 1921) and a binary orbit (Anderson 1920, Merrill 1922) with the 20-foot interferometer mounted on the 100 inch Hooker telescope on Mt. Wilson.

Because of the technical difficulties with stabilizing the OPD, the first instrument using separate telescopes performed intensity interferometry (Hanbury Brown & Twiss 1956). The stellar interferometer at Narrabri Observatory (Hanbury Brown et al. 1967) was quite productive for measurements of stellar diameters (Hanbury Brown et al. 1974a), but because of its limited sensitivity this technique has not been pursued further for astronomical applications.

4.2. Modern Interferometers

The first direct interference fringes between separate telescopes were reported by Labeyrie (1975). His instrument at the Observatoire de la Côte d'Azur, called Interféromètre à Deux Télescopes (I2T), made observations in the visible (e.g., Thom et al. 1986) and pioneered interferometry at near-IR wavelengths (Di Benedetto & Rabbia 1987). The Grand Interféromètre à Deux Télescopes (GI2T) (Mourard et al. 1994a), a successor of the I2T, uses two innovative "boule" telescopes with 1.5 m apertures on a north-south baseline that can be reconfigured from 12 m to 65 m. The GI2T has the capability of performing spectrally resolved interferometry (see Section 5.6.1). It has recently been upgraded with a new beam combination table, including a versatile visible spectrograph and an IR focus (Mourard et al. 2000). A third telescope may be added to the GI2T in the future.

Active fringe tracking was first demonstrated by the Mark I interferometer on Mt. Wilson, California (Shao & Staelin 1980), which evolved into the Mark II and Mark III (Mk III) instruments (Shao et al. 1988, Shao & Colavita 1992b). The Mk III was specifically designed to perform wide-angle astrometry, but a variable baseline that could be configured from 3 m to 31.5 m provided the flexibility needed for a variety of astronomical programs (see Section 5). Because full computer control of the siderostats and delay lines allowed almost autonomous acquisition of stars and data taking, the Mk III could observe up to 200 stars in a single night. This capability was an important factor for the calibration of instrumental effects and for the scientific productivity of the instrument.

The Navy Prototype Interferometer (NPOI), located on Anderson Mesa near Flagstaff, Arizona, combines an imaging array and an astrometric facility (Armstrong et al. 1998). It operates in the visible with 32 spectral channels covering the wavelength range from 450 nm to 850 nm. The imaging subarray consists of six movable siderostats with baseline lengths from 2.0 m to 437 m. The array geometry has been optimized for baseline bootstrapping to facilitate the imaging of stellar surface structure. Like the Mk III, the NPOI uses vacuum delay lines for pathlength compensation. The four-element astrometric subarray of the NPOI includes an extensive site metrology system that monitors the motions of the siderostats with respect to one another and to the bedrock. It is anticipated that the NPOI will perform wide-angle astrometry with $\sim$2 mas precision.

The Palomar Testbed Interferometer (PTI) (Colavita et al. 1999, Lane et al. 2000a) was built as a precursor to the Keck Interferometer to demonstrate narrow-angle astrometry with dual-star operation. Two 40 cm siderostats on a 110 m baseline provide $\sim$3 mas resolution in the near-IR (*H* and *K* bands). Aside from its role for the technical development of dual-star astrometry, the PTI is used mainly for stellar diameter measurements and binary star work (see Section 5). Both NPOI and PTI draw extensively on the experience gained with the Mk III interferometer.

The distinction of having produced the first closure-phase images from an optical synthesis array belongs to the Cambridge Optical Aperture Synthesis Telescope (COAST) (Baldwin et al. 1996). The number of telescopes has now been brought up to five, and light from four of them can be interfered simultaneously (Haniff et al. 2000). Switching between two four-telescope arrays allows taking data on nine baselines and seven closure triangles during a single night. Observations can be carried out in the visible or near-IR on relatively short baselines (up to 22 m), but longer baselines will be put into service soon.

IOTA (Infrared Optical Telescope Array) (Traub 1998) is a two-telescope interferometer on Mt. Hopkins, Arizona, with 0.45 m telescopes and baselines of up to 38 m. Observations can be carried out in the visible or near-IR. Visibilities with excellent calibration can be obtained with the single-mode fiber system FLUOR (Coudé du Foresto et al. 1998; see also Section 3.1), which accounts for a large fraction of the astronomical results obtained with IOTA. A third telescope enabling closure phase observations has recently been added (Traub et al. 2000).

The Sydney University Stellar Interferometer (SUSI) (Davis et al. 1999a,b), located near Narrabri in New South Wales, makes observations on a single baseline selected from a set of fixed north-south baselines with lengths ranging from 5 m to 640 m. (Baselines longer than 80 m have not yet been commissioned.) The 640 m baseline length, the longest of all instruments currently operational or under construction, has been chosen to resolve a sample of O stars at a wavelength of 450 nm.

The Mitaka Optical-Infrared Array (MIRA) is an ambitious plan to build a series of interferometers with increasing capabilities (Nishikawa et al. 1998). The first phase of this project (MIRA-I) (Machida et al. 1998), which consists of two telescopes with coudé optics on a 4 m baseline in Tokyo has successfully been completed and has acquired stellar fringes (Nishikawa et al. 2000). The next step will be an instrument with a 30 m baseline, and construction of a multi-telescope array is planned for the future.

The tradition of optical interferometry on Mt. Wilson continues with the CHARA array (named after Georgia State University's Center for High Angular Resolution Astronomy) (McAlister et al. 2000). It consists of six 1 m telescopes arranged in a Y-shaped configuration with a maximum baseline of $\sim$350 m. First fringes on a single baseline have been obtained, and commissioning of the full array continues. The CHARA array will be a powerful instrument for imaging of binary stars and other applications in stellar astronomy.

Whereas direct Michelson or Fizeau interferometry is by far the most sensitive technique in the visible and near-IR, heterodyne methods familiar from radio interferometry can also be used in the mid-IR. The Berkeley Infrared Spatial Interferometer (ISI) (Hale et al. 2000), located close to the CHARA array and the former site of the Mk III on Mt. Wilson, operates at wavelengths between 9 μm and 12 μm. The stellar radiation is mixed with the output of a CO_2 laser, which acts as the local oscillator. Observations have been carried out with baseline lengths up to 56 m; the commissioning of a third telescope is currently under way. The ISI has mostly been used for observations of dust around late-type stars (Section 5.5).

4.3. The Next Generation of Ground-Based Interferometers

The Keck Interferometer on Mauna Kea, HI, will consist of the two 10 m Keck telescopes and four new 1.8 m "outrigger" telescopes (Colavita et al. 1998, Colavita & Wizinowich 2000). The 10 m telescopes are equipped with high-order adaptive optics systems, which provide good correction in the near-IR and excellent wavefront quality at 10 μm. Together with spatial filtering this enables the implementation of a nulling beam combiner, which will be used to characterize exozodiacal emission around nearby main-sequence stars. A large fraction of the observing time available with the outriggers will be devoted to an astrometric search for extrasolar planets (see Section 5.8.2). The combination of all six telescopes will result in a sensitive imaging array, as it is expected that a star as faint as $K = 13.7$ will be sufficient for cophasing (Vasisht et al. 1998).

The interferometric combination of the four 8 m telescopes, augmented by movable auxiliary telescopes, has always been part of the European Southern Observatory's Very Large Telescope project. The Very Large Telescope Interferometer (VLTI) has been designed for astronomical applications such as imaging of the star cluster at Galactic Center that require a large field-of-view. It is therefore the only current instrument that implements a proper pupil relay (Section 3.5), although the general concept has changed from a contiguous 8$''$ field and homothetic pupil mapping (Eur. South. Obs. 1989) to a dual-star design (Glindemann et al. 2000). The instrumentation of the VLTI will enable a wide variety of astrophysical programs. The first generation of instruments will include a commissioning instrument based on a two-telescope near-IR fiber beam combiner (VINCI) (Kervella et al. 2000), a 10 μm instrument (MIDI, Leinert et al. 2000), a near-IR instrument with spectral resolution up to $R = 10000$ and capable of combining light from three telescopes (AMBER) (Petrov et al. 2000), and a dual-star astrometric facility (PRIMA) (Quirrenbach et al. 1998).

The Large Binocular Telescope under construction on Mt. Graham, Arizona will consist of two 8.4 m telescopes mounted side by side in a single alt-azimuth mount, with a 14.4 m center-to-center spacing. This configuration offers some unique capabilities for interferometry, as it lends itself to Fizeau beam combination with a wide field of view and low thermal background (Angel et al. 1998). Images taken at several parallactic angles can be combined with tomographic methods to synthesize a nearly round 22.8 m pupil (Bertero et al. 2000, Correia & Richichi 2000).

The concentration of large apertures on Mauna Kea ($2 \times$ Keck, Gemini, Subaru, CFHT, UKIRT) lends itself to ideas for interferometric combination of these six telescopes (Mariotti et al. 1996). By taking advantage of the existing adaptive optics systems, such a Mauna Kea array would provide excellent sensitivity and unprecedented imaging capabilities with baselines ranging from 75 m to $\sim$800 m. Implementation as an all-fiber interferometer would simplify the interfaces to the telescopes and keep costs at a minimum.

Several studies have been made of arrays consisting of a fairly large number ($\geq$15) of medium-size telescopes (Labeyrie 1998, Buscher et al. 2000). Apertures of $\sim$1.5 m are small enough to be phased with relatively simple and inexpensive low-order adaptive optics systems, yet sufficiently large to allow fringe tracking at $K \sim 12$. A maximum baseline of order 1 km would be needed for most applications in stellar astronomy. A Y-shaped configuration would be favorable for baseline bootstrapping and closure phase imaging. A more ambitious array comprised of 27 telescopes with 4 m diameter (Ridgway & Roddier 2000) would be able to address additional astronomical goals, such as resolving the broad line region in nearby quasars; the limiting magnitude for fringe tracking would be $K \sim 14$. Objects much fainter than this can only be observed with arrays that can be externally cophased. It is currently believed that giant filled-aperture telescopes can be efficiently phased with laser guide stars and tomographic wavefront sensing, but these techniques are not easily applicable to diluted arrays. One may thus plausibly expect that stellar astrophysics (including extrasolar planets) and active galactic nuclei will remain the scientific drivers of ground-based interferometry for the foreseeable future.

4.4. Space Interferometry

Interferometry from space offers access to the full IR spectral range, astrometric capabilities well beyond the atmospheric limit, and excellent sensitivity through long coherent integrations and cool optics (for the thermal IR). Several interferometric missions are therefore included in the plans of the European and US space agencies for the next decade.

The Space Interferometry Mission (SIM) (Danner & Unwin 1999) will perform astrometry of selected targets with a limiting magnitude $V = 19$. The design goal is to achieve $\sim 4\ \mu$as precision for global astrometry, and $\sim 1\ \mu$as for narrow-angle measurements (over $\sim 1^\circ$) with a 10 m baseline. SIM will address a large number of problems in Galactic dynamics, perform distance measurements of Cepheids, RR Lyrae stars, globular clusters, and X-ray binaries and detect substellar companions down to the mass range of terrestrial planets. The spacecraft geometry is monitored with a three-dimensional laser metrology system, and the attitude is determined with two guide interferometers while a third interferometer performs the actual measurements (Milman & Turyshev 2000). SIM needs a grid of ~ 3000 reference stars distributed evenly over the sky that are stable on the $\sim 4\ \mu$as level (aside from parallax and proper motion). The identification of suitable grid stars before launch is challenging because even planetary companions pose a severe problem unless stars at distances ≥ 1 kpc are chosen for the grid; K giants are therefore the best option (Patterson et al. 1999, Frink et al. 2000a,b, 2001).

A mid-IR nulling interferometer for the detection and characterization of Earth-like planets around nearby stars is being studied under the name DARWIN/IRSI by the European Space Agency (2000) and under the name Terrestrial Planet Finder (TPF) by NASA (Beichman et al. 1999) (see Section 3.3). DARWIN/TPF will operate in the 6 μm to 20 μm wavelength range, in which molecular absorption bands of carbon dioxide, water, methane, and ozone can be observed (Schindler & Kasting 2000). Early plans for a rigid ~ 50 m structure have now been largely abandoned in favor of a free-flyer architecture, with separate spacecraft for the individual telescopes. A study conducted by the European Space Agency (1996) indicated the feasibility of this approach and showed that in general, free-flying interferometers are superior to facilities on the Moon. As a technology precursor to TPF, the Space Technology 3 mission (Lay et al. 1999) will demonstrate separate-spacecraft interferometry with a baseline of up to 200 m. With a sensitivity of $\sim$1 μJy, and a few tens of mas angular resolution, DARWIN/TPF will also have a tremendous potential for general astrophysics, including detailed studies of high-redshift galaxies, which cannot be resolved by the Next Generation Space Telescope (Röttgering et al. 2000).

Many more future applications for separate-spacecraft interferometry can be imagined. An array with $\sim$150 3 m telescopes on baselines up to $\sim$150 km could deliver resolved images of an Earth twin at 3 pc, if such a planet exists (Labeyrie 1999). Direct-detection interferometry in the 40 μm $\leq \lambda \leq$ 500 μm range could revolutionize mid-IR astronomy by providing subarcsecond resolution and sufficient sensitivity to observe key lines of H_2, HD, and other species at high redshift (Shao et al. 2000). The successful detection of X-ray fringes with a grazing-incidence interferometer in the laboratory (Cash et al. 2000) suggests that astronomical interferometry may even be possible in the keV range, with the exciting perspective of imaging studies of X-ray binaries and black hole candidates.

5. ASTRONOMICAL RESULTS FROM OPTICAL INTERFEROMETRY

5.1. Stellar Diameters

The visibility function of a single star is the Fourier transform of the center-to-limb surface brightness profile $I(\rho)$:

$$V(kB) = \int_0^r \int_0^{2\pi} \cos(kB\rho \cos\phi) I(\rho)\rho \,\mathrm{d}\rho \,\mathrm{d}\phi$$
$$= 2\pi \int_0^r J_0(kB\rho) I(\rho)\rho \,\mathrm{d}\rho, \tag{12}$$

where k, B, and r denote the wave number, projected baseline length, and angular radius of the star, respectively, and J_0 is the Bessel function of zeroth order. For stars the intensity distribution is customarily given as a function of $\mu = \sqrt{1-(\rho/r)^2}$, rather than ρ itself. For polynomials

$$I(\mu) = \sum_\nu a_\nu \mu^\nu, \tag{13}$$

Equation 12 leads to (Quirrenbach et al. 1996)

$$V(kB) = \frac{1}{C} \sum_\nu a_\nu 2^{\nu/2} \Gamma\left(\frac{\nu}{2}+1\right) \frac{J_{\nu/2+1}(kBr)}{(kBr)^{\nu/2+1}} \tag{14}$$

with

$$C = \sum_{\nu} \frac{a_\nu}{\nu + 2}. \tag{15}$$

For a disk of uniform brightness $I \equiv 1$, this formula reduces to the familiar Airy pattern. (For a discussion of nonpolynomial limb darkening laws and corresponding visibility functions see Hestroffer 1997.) Equation 14 provides a direct relation between the stellar properties (angular radius r, limb darkening coefficients a_ν) and the visibility, as a function of wavenumber and baseline length. It could thus be used in a least-squares procedure to obtain the stellar radius and limb darkening coefficients from interferometer measurements. However, interferometric measurements normally cannot distinguish between a limb-darkened disk and a somewhat smaller uniform disk (except when data are taken beyond the first null of the visibility function; see below). The measurement of stellar diameters is therefore usually done in two steps (Hanbury Brown et al. 1974b): First, the diameter $\theta_{\rm UD}$ of the uniform disk model (Airy function)

$$V^2 = \left(\frac{2J_1(x)}{x}\right)^2, \quad x \equiv kBr_{\rm UD} = \frac{\pi B\theta_{\rm UD}}{\lambda}, \tag{16}$$

is fitted to the visibility data. In the second step $\theta_{\rm UD}$ is multiplied with a correction factor derived from model atmospheres to obtain the limb-darkened diameter $\theta_{\rm LD}$. This procedure keeps the observed and theoretical contributions to $\theta_{\rm LD}$ cleanly separate in two independent factors.

The limb-darkening correction usually increases with decreasing effective temperature and gravity, and for the same star it is larger at shorter wavelengths. The latter trend leads to a wavelength dependence of the observed uniform disk diameter (Mozurkewich et al. 1991, Quirrenbach et al. 1996), and it makes diameter measurements in the near-IR less susceptible to systematic errors owing to problems with the limb-darkening correction. For example, the correction factors applicable to Arcturus range from $\sim$1.03 at 2.2 μm to $\sim$1.12 at 450 nm (Quirrenbach et al. 1996). Limb-darkening laws, which can be used to compute interferometric correction factors, have been tabulated for Bell et al. models of cool giant and supergiant atmospheres (Manduca et al. 1977, Manduca 1979), for Kurucz model atmospheres (Van Hamme 1993, Díaz-Cordovés et al. 1995, Claret et al. 1995), and for non-Mira M giant and M-type Mira models by Bessell et al. (Hofmann et al. 1998, Hofmann & Scholz 1998).

5.1.1. STELLAR EFFECTIVE TEMPERATURES The effective temperature of stars is defined by

$$T_{\rm eff} \equiv \left(\frac{L}{4\pi\sigma R^2}\right)^{1/4} = \left(\frac{4 f_{\rm bol}}{\sigma\theta_{\rm LD}^2}\right)^{1/4}, \tag{17}$$

where L is the luminosity, R the stellar radius, $f_{\rm bol}$ the bolometric flux, and σ the Stefan-Boltzmann constant. The most direct and model-independent way of measuring effective temperatures is thus the combination of bolometric fluxes with angular diameters. More indirect methods such as the infrared flux method (IRFM) (Blackwell & Shallis 1977), which uses the ratio of total integrated flux to K-band flux as temperature indicator, can be validated by comparison with directly determined effective temperatures.

The first extensive set of angular diameter measurements of 32 stars was obtained with the Narrabri Intensity Interferometer (Hanbury Brown et al. 1974a). Because of the very bright limiting magnitude of this instrument ($B = 2.5$) and its blue-sensitive detectors, stars with early spectral types were observed. Because the observations were carried out with baseline lengths up to 188 m, these are still the best diameters for A and B stars. Fairly large samples of cooler stars, mostly G, K, and M giants, have been observed in the visible with the Mk III (Mozurkewich et al. 1991, 2001) and the NPOI (Nordgren et al. 1999). The Mk III sample consists of 82 stars; more than half of them have angular diameters at 800 nm with formal errors $\leq$1%. These data have been used to compute average relations between the surface brightness and $V - R$ or $V - K$ colors. These relations are useful for predictions of expected angular sizes, and they can aid intercomparisons of different instruments or methods. A star-by-star comparison between the Mk III and NPOI data sets (40 stars in common) indicates good agreement between the two instruments; a comparison between Mk III diameters and the IRFM (34 stars in common) gives a median systematic offset of only 0.3% (Mozurkewich et al. 2001).

Angular diameter data in the near-IR have been collected with the I2T (Di Benedetto & Rabbia 1987), IOTA (Dyck et al. 1996a, 1998), and PTI (van Belle et al. 1999). There are few stars in common between these data sets, but the IOTA and PTI data on the effective temperature scale for K and M giants as a function of spectral type agree with each other. The average standard deviation of the temperatures in each spectral subtype bin for the combined sample is $\Delta T \approx 270$ K (van Belle et al. 1999). The mean error of the diameter data, 3.9%, and the bolometric flux, 17%, correspond to only $\pm$60 K and $\pm$130 K. The remainder of the temperature scatter could be due to errors in the spectral classification by two subtypes or to genuine star-by-star variations. The latter interpretation is consistent with the large χ^2 in the surface brightness–color relation from the Mk III data, which has been attributed to an intrinsic scatter (Mozurkewich et al. 2001). There also seems to be a systematic discrepancy between the PTI diameters and the Mk III surface brightness versus $V - K$ relation; more data on a common sample of stars with properly scaled baselines are needed to resolve this issue. On the other hand, the photospheric diameters of α Sco and α Ori measured at 11.15 μm agree remarkably well with visible determinations (Bester et al. 1996).

A sample of very cool stars, including spectral types as late as M8III, has been observed with the FLUOR beam combiner at IOTA (Perrin et al. 1998). These data have been used to extend the temperature scale of oxygen-rich giant stars down to $T_{\rm eff} = 2800$ K. Diameters of carbon stars have been measured at wavelengths between 700 nm and 800 nm with the Mk III (Quirrenbach et al. 1994c) and at 2.2 μm with IOTA (Dyck et al. 1996b). Temperatures derived from optical and

IR interferometry, lunar occultations, and the IRFM disagree by up to ~300 K for some stars, and there is no clear trend of effective temperature with spectral classification. These difficulties reflect the uncertainties of the atmosphere models needed for the IRFM and for the limb-darkening correction; interferometric measurements of the surface brightness profiles at multiple wavelengths are clearly needed. There seems to be a general tendency, however, for the effective temperature to decrease from the coolest oxygen-rich stars to S-type to carbon stars (van Belle et al. 1997).

5.1.2. WAVELENGTH DEPENDENCE OF STELLAR DIAMETERS In addition to the variation of the uniform disk diameter with wavelength due to limb darkening (see above), one may expect a true wavelength dependence of the photospheric diameter in cool stars due to the wavelength dependence of the opacity. In essence, interferometry measures the diameter of the $\tau = 1$ surface, and the atmospheres of cool stars are so tenuous and extended that they become opaque at a substantially larger radius in absorption bands of molecules such as TiO than in the continuum. In fact, the very concept of a "photospheric diameter" becomes problematic under these circumstances (Baschek et al. 1991). It is frequently identified with the distance from the stellar center at which the Rosseland optical depth equals unity, but even that definition is not without difficulties for the coolest Mira stars for which monochromatic diameters at the same pulsational phase may differ from one another by a factor ~2 (Hofmann et al. 1998). Strong variations of diameter with TiO absorption depth have indeed been observed in the Mira variables *o* Ceti (Labeyrie et al. 1977, Bonneau et al. 1982), R Leo (Di Giacomo et al. 1991), R Cas (Haniff et al. 1995), and R Dor (Jacob et al. 2000a) in qualitative agreement with model predictions (Jacob et al. 2000b). Time series of measurements in well-defined narrow filters covering several pulsational cycles will be required for a more detailed comparison between observations and theory.

A sample of 42 stars, mostly "normal" K and M giants and supergiants, has been observed with the Mk III through 10 nm wide interference filters centered inside the strong TiO band at 712 nm and in the pseudo-continuum at 754 nm (Quirrenbach et al. 1993c, 2001). K stars, M0 stars, and carbon stars have identical diameters at 712 nm and 754 nm. For stars with spectral types later than M0, the diameters are systematically larger at 712 nm than at 754 nm. The diameter ratio increases with decreasing effective temperature, and it is larger for luminosity class I than luminosity class II and III stars. Although these trends are qualitatively expected, the observed 712 nm/754 nm diameter ratios are larger than predicted by current model atmospheres (B Plez, personal communication). This indicates that the available models do not adequately describe the TiO opacity in the tenuous outer layers of the atmosphere or at the base of the wind; the interferometric data lend support to the existence of an extended "molecular sphere," which has been postulated on the basis of IR spectroscopy of water bands in μ Cep (Tsuji 2000). Wavelength-dependent measurements of stellar diameters with spectral resolution across TiO and other molecular bands can thus provide a powerful new tool for the study of the extended atmospheres of cool stars.

5.1.3. PULSATION OF MIRA VARIABLES Interferometry plays an important role in the controversy about the pulsation mode of Mira variables. The pulsation constant

$$Q = P(M/M_\odot)^{1/2}(R/R_\odot)^{-3/2} \tag{18}$$

can be determined by combining the period P, angular diameter, and parallax with plausible values for the mass range of Mira stars, and then it can be compared with theoretical values of Q for fundamental-mode or first-overtone pulsations. The two major current limitations of this method are the large uncertainties of the HIPPARCOS parallaxes of most Miras and the difficulties in converting interferometric data to Rosseland mean diameters, owing to the strong limb darkening and wavelength dependence of the diameter. Based on measurements of 10 Miras at wavelengths between 700 nm and 900 nm, Haniff et al. (1995) concluded that all Miras have linear radii $\gtrsim 350\ R_\odot$, which implies that they are first-overtone pulsators. Furthermore, their data are consistent with a small mass range $1\ M_\odot \lesssim M \lesssim 1.5\ M_\odot$. In contrast, van Belle et al. (1996) derive a much larger spread of the linear diameters from observations of 18 Miras at 2.2 μm, implying a larger mass range and a mix of fundamental-mode and first-overtone pulsators. Interestingly, recent measurements at 2.2 μm of R Leo (Perrin et al. 1999), R Aql (Hofmann et al. 2000b), and R Cas (Weigelt et al. 2000) give systematically much smaller Rosseland radii than observations of the same stars at wavelengths $\lesssim 1\ \mu$m (Burns et al. 1998, Hofmann et al. 2000a, Haniff et al. 1995), which leads to the paradoxical situation that observations at the shorter wavelengths imply first-overtone, and those at $\lambda \geq 2\ \mu$m fundamental-mode pulsations. It is possible that the atmosphere models are at fault, but the presence of surface structure and deviations from circular symmetry as observed in *o* Ceti (Quirrenbach et al. 1992, see below) may also play a major role. Systematic observations with sufficient spectral resolution and coverage, with a range of baseline lengths, and covering the full pulsation cycle will be needed for critical tests of the model atmospheres, and for the eventual determination of the pulsation mode of Mira variables.

5.2. Limb Darkening and Stellar Structure

5.2.1. INTERFEROMETRIC MEASUREMENTS OF LIMB DARKENING Because the correct treatment of limb darkening is crucial for precise measurements of effective temperatures, it is important to perform observational checks of theoretical limb-darkening curves. This is a fairly difficult task, because data are required around and beyond the first null of the visibility function (Equation 12), where the SNR is low. The first such measurement was carried out with the Narrabri Intensity Interferometer; data from a 203-hour(!) integration on Sirius showed that the height of the second maximum of the visibility function was consistent with the prediction from a model atmosphere (Hanbury Brown et al. 1974b). Similar observations were carried out for Arcturus with the Mk III Interferometer (Quirrenbach et al. 1996). This experiment used wavelength bootstrapping (i.e., fringe tracking at $\lambda \approx 750$ nm, data taking at 450 nm, 500 nm, and 550 nm) and phase-referenced visibility averaging to obtain high signal-to-noise on data with visibilities as low as $V^2 \sim 10^{-4}$. The effective temperature of Arcturus was determined to be 4303 ± 47 K, and both the wavelength dependence of the uniform disk diameter and the shape of the visibility function were found to be in excellent agreement with the predictions of the models with $T_{\rm eff} = 4000$ K or $T_{\rm eff} = 4500$ K, $\log g = 1.5$ tabulated by Manduca et al. (1977) and Manduca (1979).

With the NPOI it is possible to use baseline bootstrapping to reach spatial frequencies beyond the first visibility null. Using this technique, Pauls et al. (1998) showed that limb-darkened diameters of cool stars can be obtained directly by fitting visibility curves with appropriate limb-darkening laws to the data. The quality of the fit near the first null provides an internal check of this method: The fit obtained with color-dependent limb-darkening coefficients from Van Hamme (1993) is good, but attempts to fit a uniform disk or a grey linear limb-darkening law with the Eddington approximation fail miserably. The expected 180° jump of the closure phase at the position of the null on the longest baseline has also been observed (Hajian et al. 1998).

The brightness profiles of cool supergiants (e.g., Betelgeuse) (Burns et al. 1997) and Mira variables (e.g., R Leo) (Perrin et al. 1999) have sufficiently soft edges or extended wings that a description by a polynomial (Equation 13) is inadequate. It would be very interesting to obtain visibility measurements of similar objects with high signal-to-noise and dense spatial frequency sampling, so that a direct determination of the surface brightness profile by taking the Fourier transform of the data would be possible. For the more "normal" warmer stars, interferometric measurements can provide important cross-checks of limb-darkening parameters determined with other methods (e.g., light curves of eclipsing binaries and of gravitational microlensing events) and useful tests of stellar atmosphere models.

5.2.2. STELLAR SURFACE STRUCTURE Observations of Mira (*o* Ceti) were carried out at $\lambda = 800$ nm with the shortest baselines (3.0–6.6 m) of the Mk III in 1990 at photometric phases 0.96, 0.05, and 0.14 (Quirrenbach et al. 1992). These measurements revealed that the star was not spherically symmetric and that variations in the size and position angle of the asymmetric structure occurred on a time scale of a few weeks. Comparison with a small amount of data taken one year earlier at almost the same phases also showed pronounced changes from cycle to cycle. These data could not be fitted with a single elliptical Gaussian or uniform disk, indicating the presence of surface structure. Results from aperture masking observations of five long-period variables at the William Herschel Telescope (WHT) can similarly be explained with bright compact surface features contributing between 5% and 20% of the stellar flux (Tuthill et al. 1999a). Young et al. (2000) have combined near-simultaneous WHT and Cambridge Optical Aperture Synthesis Telescope data at 700, 905, and 1290 nm on Betelgeuse. They found a strong asymmetry at 700 nm, a small one at 905 nm, and none at 1290 nm, and they suggested a model in which bright features are seen in regions where the atmospheric opacity has been reduced as a result of activity, perhaps convection. The questions raised by these first direct glimpses of variable stellar surface structure will have to be answered by observations with better resolution, *uv* coverage, and spectral coverage of the TiO absorption bands.

5.3. Cepheids

Diameter measurements of the nearest Cepheids have received much attention because they can contribute to distance estimates and thus to the calibration of the Cepheid period—luminosity relation, the first rung of the cosmological distance ladder. One approach is to combine the mean angular diameter with an estimate of the linear diameter (Mourard et al. 1997, Armstrong et al. 2001). This method is rather indirect as linear diameters derived from multicolor photometry (Barnes-Evans method) or from a combination of photometric data with radial-velocity observations (Baade-Wesselink method) are strongly model-dependent. Individual implementations of these methods differ in the way observations for the specific star under consideration, mean relations between surface brightness, colors and period, and atmospheric models are combined with each other. Interferometric diameter measurements of nonpulsating stars actually help in this regard, as they can be used for an empirical calibration of surface brightness—color relations (Welch 1994). However, there are still substantial disagreements between the different approaches; for example, there are discrepancies of up to $\pm 20\%$ between published linear diameters for δ Cep (Turner 1988, Fernley et al. 1989, Ripepi et al. 1997). Care must also be taken to properly average the data over the pulsation cycle.

Based on their GI2T measurement of 1.60 ± 0.12 mas for the limb-darkened diameter of δ Cep, and adopting a mean linear radius of $42.7 \pm 1\, R_{\odot}$ (a very optimistic error estimate from Turner 1988), Mourard et al. (1997) derived a distance of 240 ± 24 pc, corresponding to a parallax $\pi = 4.2 \pm 0.4$ mas. Armstrong et al. (2001) measured $\theta_{\mathrm{LD}} = 1.520 \pm 0.014$ mas with the NPOI and adopted $R = 41.5 \pm 5.1\, R_{\odot}$, which gives $D = 254 \pm 30$ pc, or $\pi = 3.94 \pm 0.47$ mas. The distance of δ Cep from interferometric measurements is thus in reasonable agreement with the HIPPARCOS value $\pi = 3.32 \pm 0.58$ mas, and has comparable accuracy. Application of the same method to η Aql ($\theta_{\mathrm{LD}} = 1.69 \pm 0.04$ mas, $R = 51.6 \pm 5.6\, R_{\odot}$) gives $D = 284 \pm 31$ pc or $\pi = 3.52 \pm 0.37$ mas (Armstrong et al. 2001), again in agreement with HIPPARCOS, $\pi = 2.78 \pm 0.91$ mas.

It is also possible to turn the above argument around and use parallaxes and angular diameters to derive the linear diameters of Cepheids. The current error bars are too large to differentiate between different alternatives for the period–radius relation (Nordgren et al. 2001), but the expected improvements in parallax measurements from space (DIVA, FAME, SIM, GAIA) and in angular diameters (from long baselines provided by NPOI, CHARA Array, VLTI) will make this an attractive method for the determination of physical parameters of Cepheids.

A promising way to obtain accurate Cepheid distances, which avoids the calibration uncertainties of the Barnes-Evans and Baade-Wesselink methods, is the direct interferometric measurement of the pulsation amplitude. If both angular diameter and radial velocity are known as a function of phase, it is possible to perform a simultaneous fit of distance and mean linear radius to the data. The pulsations of δ Cep and η Aql have been detected at the $\sim 1.5\,\sigma$ to $\sim 2\,\sigma$ level in NPOI data taken in the red part of the visible spectrum with baseline lengths up to 37.5 m (Armstrong et al. 2001). The distances derived from the pulsation solutions are consistent with the values quoted above but have very large error bars. PTI observations of ζ Gem at 1.6 μm (H band) show pulsations at a higher level of significance; the best-fit model parameters are $D = 336 \pm 44$ pc (i.e., $\pi = 2.98 \pm 0.39$ mas, compared with $\pi = 2.79 \pm 0.81$ from HIPPARCOS) and $R = 62 \pm 11\, R_\odot$ (Lane et al. 2000b). This result demonstrates that it is indeed possible to use interferometry and spectroscopy directly to determine Cepheid distances, without taking recourse to intermediate steps based on photometric methods.

Systematic errors remain, however. The radial velocity curves obtained from different lines in the same star may show substantial differences in amplitude, and asymmetries and line splitting observed in high-resolution IR spectra indicate the presence of pulsationally driven shocks (e.g., Sasselov & Lester 1990). This complicates the task of linking the spectroscopic and interferometric data, which should ideally trace the same layer in the star. Furthermore, corrections for limb darkening have to be applied to the angular diameter as well as the radial velocity curves. (The observed radial velocity is an integral over the stellar disk of the properly flux-weighted line-of-sight component of the pulsational velocity.) The correction factors currently have to be taken from model atmospheres, but interferometric observations on long baselines should soon directly provide empirical limb-darkened diameters. A more challenging future task is the measurement of angular diameter curves and limb darkening in the spectral lines used for the radial velocity curves, which would all but eliminate the current systematic uncertainties.

5.4. Binaries

The first interferometric determination of a "visual" orbit of a double-lined spectroscopic binary (SB2), namely Capella, by Anderson (1920) and Merrill (1922), showed the potential of this method for measuring stellar masses and distances. Adding the inclination from the interferometric orbit to the spectroscopic elements allows computation of the component masses, and combining the angular diameter of the orbit with the physical scale set by the spectroscopy yields the distance, or "orbital parallax." Because of the fundamental importance of these data, extensive observations of SB2s have been carried out with the Mk III, NPOI, and PTI (see Table 1).

TABLE 1 Interferometric determinations of orbits and component masses for double-lined spectroscopic binaries

System	Types	a [mas]	M_1 [$M_\odot$]	M_2 [$M_\odot$]	Instr.	Ref.	Remarks
θ Aql	B9.5III + B9.5III	3.2	3.6 ± 0.8	2.9 ± 0.6	Mk III	H95	
β Aur	A2V + A2V	3.3	2.41 ± 0.03	2.32 ± 0.03	Mk III	H95	eclipsing
12 Boo	F9IV + F9IV	3.4	1.435 ± 0.023	1.408 ± 0.020	PTI	B00	
64 Psc	F8V + F8V	6.5	1.223 ± 0.021	1.170 ± 0.018	PTI	B99b	
93 Leo	G5III + A7V	7.5	2.25 ± 0.29	1.97 ± 0.15	Mk III	H95	
ζ^1 UMa	A2V + A2V	9.6	2.51 ± 0.08	2.55 ± 0.07	Mk III	H95	
		9.8	2.43 ± 0.07	2.50 ± 0.07	NPOI	H98	
ι Peg	F5V + G8V	10.3	1.326 ± 0.016	0.819 ± 0.009	PTI	B99a	
η And	G8III + G8III	10.4	2.59 ± 0.30	2.34 ± 0.22	Mk III	H93	
α Equ	G2III + A5V	12.0	2.13 ± 0.29	1.86 ± 0.21	Mk III	A92b	
ζ Aur	K4Ib + B5V	16.2	5.8 ± 0.2	4.8 ± 0.2	Mk III	B96	eclipsing
θ^2 Tau	A7III + A:	18.6	2.1 ± 0.3	1.6 ± 0.2	Mk III	T95	Hyades
ϕ Cyg	K0III + K0III	23.7	2.536 ± 0.086	2.437 ± 0.082	Mk III	A92a	
α And	B8IV + A:	25.2	5.5:	2.3:	Mk III	P92,T95	
β Ari	A5V + G0V:	36.1	2.34 ± 0.10	1.34 ± 0.07	Mk III	P90	
α Aur	G1III + G8III	55.7	2.56 ± 0.04	2.69 ± 0.06	Mk III	H94a	

References: A92a, Armstrong et al. 1992a; A92b, Armstrong et al. 1992b; B96, Bennett et al. 1996; B99a, Boden et al. 1999a; B99b, Boden et al. 1999b; B00, Boden et al. 2000; H93, Hummel et al. 1993; H94a, Hummel et al. 1994a; H95, Hummel et al. 1995; H98, Hummel et al. 1998; P90, Pan et al. 1990; P92, Pan et al. 1992; T95, Tomkin et al. 1995.

A number of additional binaries have been observed interferometrically: Algol, an eclipsing triple system (Pan et al. 1993), the single-lined spectroscopic systems 113 Her (Hummel et al. 1995) and η Peg (Hummel et al. 1998), the nearly face-on SB2 δ Tri (Hummel et al. 1995), and the SB2 β Tri, for which the spectroscopic elements are questionable (Hummel et al. 1995). Interferometric measurements have been used to infer the duplicity of ζ Ori A (Hummel et al. 2000) and possibly FU Ori (Berger et al. 2000b).

The analysis of the earlier Mk III data (Pan et al. 1990, 1992, Armstrong et al. 1992a) proceeded in two steps. First, the separation ρ and position angle θ were fitted to the visibility data for each night, along with stellar parameters (diameters, colors, brightness ratio); this was possible because the three wavelength channels provided sufficient data to constrain ρ and θ. Then the orbit was determined from a least-squares fit to the nightly values of ρ and θ (for details see Armstrong et al. 1992a). An improvement of this algorithm takes into account linear approximations $d\rho/dt$ and $d\theta/dt$ of the orbital motion in the course of each night (Armstrong et al. 1992b). A better method was developed by Hummel et al. (1993) and applied in most subsequent analyses. Here the seven orbital elements and relevant stellar parameters are fitted directly to the observed visibilities. This approach has the advantages that fast orbital motion of short-period binaries is taken into account properly, and that small data sets from nights that do not allow determination of ρ and θ can still be used. The latter property is crucial for the PTI, which unlike the Mk III normally provides data in only one wavelength band. Hummel et al. (1998) have generalized the “global” approach by fitting orbital parameters in a simultaneous solution to interferometric and radial velocity data.

The determination of precise stellar masses requires good spectroscopy and interferometry. Because the masses depend on $\sin^3 i$, nearly face-on systems are not suited for mass measurements, and most speckle orbits are not precise enough to give masses to better than 10% (for a recent compilation see Pourbaix 2000). The stellar masses determined from interferometric observations of SB2s are summarized in Table 1. The orbital solutions and error estimates are taken from the references cited. Although all error bars refer formally to 1 σ, some authors may be more conservative than others in assessing systematics in the data or dealing with discrepancies between different published spectroscopic data sets or between spectroscopy and interferometry. It should be noted that determining the scale of the orbit (in mas), and the subsequent computation of the orbital parallax, requires knowledge of the effective central wavelength of the interferometric observations, which depends on the stellar color (Hummel et al. 1994a). Systematic errors in this quantity may easily go unnoticed because they do not affect the χ^2 of the orbit fit.

It is instructive to compare Table 1 to the masses of eclipsing binaries listed by Andersen (1991). Only a handful of the interferometrically determined masses meet Andersen’s 2% accuracy criterion for being useful for critical tests of main-sequence stellar models, and the baselines used in the observations compiled in Table 1 are too short to give good stellar radii (with the exception of Capella). On the other hand, the agreement of the component masses of β Aur, the only system in common between the two samples, is encouraging. Furthermore, analyses of pairs with evolved components such as Capella, ϕ Cyg, and α Equ provide useful tests of post-main-sequence evolutionary models (e.g., Armstrong et al. 1992b). The availability of orbital parallaxes, which are often better than HIPPARCOS values, is a clear advantage in this respect. Further improvements can be expected soon because NPOI and the CHARA Array will provide long baselines ($\gtrsim$100 m) in the visible and increased sensitivity compared with the Mk III. This will make many more SB2s available for precise interferometric orbit determination. The key to noticeable progress will be observations of stars with well-determined spectroscopic elements and state-of-the-art determination of the metal abundance. Many of the eclipsing systems in Andersen (1991) are also accessible to NPOI and CHARA, which could provide improved distances and better luminosity ratios for partially eclipsing systems. The good instantaneous coverage of the *uv* plane afforded by the multiple baselines and wavelength channels of these two arrays will allow determination of orbits from snapshot observations, making the NPOI and the CHARA Array very efficient instruments for binary programs.

5.5. Mid-Infrared Observations of Dust Shells Around Late-Type Stars

The observing program of the UC Berkeley Infrared Spatial Interferometer (ISI) has been aimed mainly at determining the spatial structure and temporal evolution of dust shells around long-period variables. Data for 13 stars obtained on baselines ranging from 4 m to 13 m at 11.15 μm have been discussed in detail by Danchi et al. (1994). This survey showed that the radius of dust formation depends on the spectral type of the star. The oxygen-rich supergiants α Sco, α Ori, and α Her, with spectral classifications ranging from M1.5 to M5, have dust shells with inner radius r_0 far from the star, $20 \leq r_0/r_* \lesssim 50$. This is consistent with sporadic formation of dust in episodes occurring in intervals of $\sim$50 years. In contrast, the Mira stars *o* Ceti, R Leo, and IK Tau, as well as the carbon star IRC +10216, have dust close to the stellar photosphere, $2 \lesssim r_0/r_* \lesssim 5$, indicating dust formation during each pulsation cycle. Semiregular and irregular variables were found to be similar to the Miras, and S stars were intermediate between the extremes.

Subsequent observations with ISI have expanded the target list by adding stars with little visible flux (IRC +10011 and IRC +10420) (Lipman et al. 2000) and focused on more detailed studies of individual stars. Data on *o* Ceti from 1988–1995 show a strong dependence on the pulsation phase but are inconsistent with simple heating and cooling of the dust due to the changes in luminosity (Lopez et al. 1997). A good fit can be obtained with models that include inhomogeneities such as clumps or partial shells, but the *uv* coverage is insufficient to distinguish between these alternatives. Temporal variations in the visibilities of IK Tau can be explained with an expanding multiple-shell model (Hale et al. 1997). ISI and near-IR speckle observations of NML Cygni also indicate the presence of two concentric dust shells (Monnier et al. 1997), which appear to be moving outward

(Danchi et al. 1999). Sudol et al. (1999) have obtained complementary visibility data at lower spatial frequencies using the 2.3 m Wyoming Infrared Observatory telescope. They find good agreement with the models based on the ISI observations for VY CMa, IRC +10216, χ Cyg, and IK Tau, but suggest a model including both steady and episodic mass loss for α Ori and a bipolar outflow model for NML Cyg. Near-simultaneous observations with good *uv* coverage and closure phases are clearly needed to resolve these issues and to obtain better constraints on models for objects such as *o* Ceti, VY CMa, and the symbiotic star R Aqr, for which the ISI data indicate strong deviations from spherical symmetry (Tuthill et al. 2000a, Monnier et al. 2000a). The impending addition of a third telescope to ISI will be an important step in that direction. The rotating spiral structure of the dust around the Wolf-Rayet binary WR 104 (Tuthill et al. 1999b) should serve as a reminder that complicated morphologies may be the rule rather than the exception.

The recent addition of a filter bank system to the ISI (Monnier et al. 2000c) has enabled spectrally resolved observations of NH_3 and SiH_4 absorption lines towards IRC +10216 and VY CMa (Monnier et al. 2000b). The visibility ratios on and off the lines are consistent with unity in all cases, which implies that ammonia and silane molecules are formed at large radii, significantly beyond the dust formation zone.

5.6. Be Stars and Herbig Ae/Be Stars

5.6.1. CLASSICAL BE STARS Classical Be stars are rapidly rotating hot stars of luminosity class III to V that have shown emission lines at least at one epoch. The structure of Be star envelopes has long been the subject of intense study and debate, with the majority view that the geometry is a fairly thin disk (e.g., Poeckert & Marlborough 1978, Bjorkman & Cassinelli 1993), but a minority advocating a moderately flattened "onion-shell" structure (e.g., Doazan 1987). These models can be tested directly, and subsequently refined, with interferometric observations of the emission lines. The first successful observations of a Be star, γ Cas, in the light of the Hα line were carried out with the I2T; the diameter of the emission region (FWHM of a Gaussian model) was found to be 3.25 ± 0.4 mas (Thom et al. 1986). Taking advantage of the improved sensitivity afforded by the 1.5 m apertures of the GI2T, Mourard et al. (1989) observed γ Cas with a spectral resolution of 1.5 Å, which gives seven resolution elements across the Hα line. They found pronounced changes of the visibility modulus and phase across the line, and concluded that their data agreed well with the predictions from a rotating disk model with inclination $i \sim 45^\circ$. These results clearly demonstrated the potential of observations that combine spectral and spatial resolution, but also that extensive modeling is required to interpret measurements obtained with very limited sampling of the *uv* plane. Stee & de Araújo (1994) and Stee et al. (1995) have developed an axisymmetric model for Be star envelopes based on a radiatively driven wind, and shown that the free parameters of this model can be constrained by comparison of predicted line profiles and visibilities with GI2T data. Similar observations in the Hβ and HeI λ6678 lines indicate that the Hβ-emitting region is only about half the size of the Hα region, and the HeI region even smaller (Stee et al. 1998). This should come as no surprise, as the excitation and ionization decrease with distance from the star; the Hα region was also found to be much larger than the HeI region in the luminous blue variable P Cygni (Vakili et al. 1997).

The geometry of seven Be star envelopes was determined more directly with the Mk III (Quirrenbach et al. 1993b, 1994a, 1997). Observations in a 1 nm wide filter centered on Hα were carried out on a number of different baselines and over the largest possible range in hour angle to obtain a good coverage of the uv plane through Earth-rotation synthesis. Although the program stars were unresolved in the 550 nm continuum, as expected from estimates of the photospheric diameters, they were all resolved in the Hα filter. Simple models consisting of a single elliptical Gaussian were fitted to the visibilities (see Table 2), with the exception of β CMi, for which the data were insufficient to constrain the ellipticity and a circular Gaussian model was used. For each star the contribution of photospheric light within the 1 nm wide filter was estimated from photometric data, and a small uniform disk with appropriate brightness subtracted before the model fitting. Comparison of the major axes ($a = 2.6 \ldots 4.5$ mas) with estimates of the photospheric diameters from the Barnes-Evans relation ($a^* = 0.34 \ldots 0.74$ mas) shows that the Hα regions have sizes ranging from about 3.5 to 12 photospheric diameters. In the Mk III sample the normalized disk size a/a^* is correlated with indicators of the circumstellar material, such as the equivalent width of the Hα line and 12 μm excess, but not with the equatorial rotation velocity.

The axial ratios r span a wide range, with $r < 0.5$ for ϕ Per, ψ Per and ζ Tau, intermediate ellipticity ($r = 0.7$) for γ Cas, and $r \sim 1$ for η Tau and 48 Per. This can easily be understood as an inclination effect in disk models, which require $r \geq \cos i$, with near-equality for thin disks. The strong correlation of the minimum inclination derived in this way with polarimetric estimates supports the thin-disk hypothesis (Quirrenbach et al. 1997); a disk opening angle of at most 20° is also implied by $r = 0.3$ as found for ζ Tau. In the four cases in which the ellipticity is significant, the position angle ϕ of the major axis is perpendicular to the position angle χ of the intrinsic linear polarization, as expected if the polarization is due to electron scattering in an optically or geometrically thin disk. Combining the lower limit on the inclination $i_{\mathrm{min}} = \arccos r$ with the projected rotation velocity $v \sin i$ gives an upper limit v_{max} for the true equatorial rotation velocity v, which is thus restricted to the range $v \sin i \leq v \leq v_{\mathrm{max}}$. Comparison of v with theoretical estimates of the critical rotation velocity (i.e., the rotation velocity at which the centrifugal and gravitational forces would be equal at the equator) (Collins 1974) indicates that γ Cas and ζ Tau rotate at a clearly subcritical rate, whereas the rotation of ϕ Per and ψ Per is close to critical.

Whereas these analyses have all been carried out within the framework of axisymmetric models, deviations from such symmetry have been found in GI2T

TABLE 2 Measurements of Be star disks with the Mk III Interferometer. The V magnitudes are from the Bright Star Catalog, Hα excesses E_α from Coté & Waters (1987), $v \sin i$ from Slettebak (1982), the intrinsic polarization position angles χ and interferometric results from Quirrenbach et al. (1997). The central part of the table gives Gaussian model fits to the interferometer data, with formal uncertainties. The models have three free parameters: major axis a, axial ratio r, and position angle of the major axis ϕ. Point source models accounting for the photospheric emission were subtracted before model fitting. The two columns on the right list the difference between polarimetric and interferometric position angles, and minimum inclination derived from the Mk III data.

Star	V	E_α [Å]	$v \sin i$ [km s^{-1}]	χ [°]	a [mas]	r	ϕ [°]	$\lvert\phi - \chi\rvert$ [°]	i_{min} [°]
γ Cas	2.47	22.7	230	−69	3.47 ± 0.02	0.70 ± 0.02	19 ± 2	88	46
ϕ Per	4.07	57.6	400	25	2.67 ± 0.20	0.46 ± 0.04	−62 ± 5	87	63
ψ Per	4.23	43.2	280	45	3.26 ± 0.23	0.47 ± 0.11	−33 ± 11	78	62
η Tau	2.87	11.9	140	—	2.65 ± 0.14	0.95 ± 0.22	(19)	—	18
48 Per	4.04	24.2	200	—	2.77 ± 0.56	0.89 ± 0.13	(68)	—	27
ζ Tau	3.00	26.9	220	33	4.53 ± 0.52	0.28 ± 0.02	−58 ± 4	91	74
β CMi	2.85	11.5	245	—	2.65 ± 0.10	—	—	—	—

data on ζ Tau (Vakili et al. 1998) and γ Cas (Berio et al. 1999b); these data have been interpreted as indications of one-armed ($m = 1$) oscillations in the disk. Even more detailed questions regarding the formation, structure, and dynamics of Be star disks can be addressed with future monitoring of multiple emission lines, which will combine high spatial and spectral resolution with good *uv* and time coverage (Quirrenbach 1997).

5.6.2. HERBIG AE/BE STARS Herbig Ae/Be stars are intermediate-mass ($1.5\,M_\odot \lesssim M \lesssim 10\,M_\odot$) pre–main-sequence stars. Many of these objects have an IR excess indicating the presence of circumstellar disks or shells. A sample of 15 Herbig Ae/Be stars has been observed at the IOTA in the H and K' IR bands; 11 of the stars have been resolved with the 21 m and 38 m baselines of this instrument (Millan-Gabet et al. 2000). The visibilities of MWC 361-A show the clear signature of a binary with ~18 mas separation. None of the other resolved sources show any indication for a departure from circular symmetry on the sky. This suggests a spherical distribution of the material responsible for the IR excess, possibly in a thin spherical shell. The data for individual objects are also consistent with a flattened distribution seen nearly face-on (e.g., AB Aur) (Millan-Gabet et al. 1999), but this interpretation seems unlikely for the ensemble of observations. It is important to note that standard accretion disk models generally provide a poor fit to the visibility data. The linear sizes (for Gaussian models) range from $\leq$0.5 AU (unresolved) to ~6 AU, with a tendency for the largest sizes of the excess to occur for the hottest and most luminous stars. There is a considerable spread in this relation, however, indicative of a loose connection between the stellar parameters and the properties of the IR excess. The measured sizes, combined with the near-IR fluxes, require the emission to be optically thin, consistent with the requirement that the stars are visible, although viewed through a spherically symmetric envelope. Follow-up observations of Herbig Ae/Be stars with longer baselines and more complete *uv* coverage will certainly provide a wealth of information about the distribution of the circumstellar material, which is crucial for understanding the star formation process.

5.7. Nova Cygni 1992

Nova Cygni 1992 was observed with the Mk III Interferometer ~10 days after maximum light (Quirrenbach et al. 1993a). Combining the diameter of 5.1 mas measured in a 10 nm wide filter centered on the Hα line with an average observed expansion velocity of ~1100 km s^{-1} gave a distance to the nova of ~2.5 kpc, in good agreement with later Hubble Space Telescope (HST) results (Paresce et al. 1995). The limited time, baseline, and wavelength coverage of the Mk III data did not allow any more detailed modeling of the structure and evolution of the nova envelope. However, with the arrays currently under construction it will be possible to carry out spectrally resolved observations of emission lines for bright novae with $V_{\max} \approx 7$, which occur about once per year.

5.8. Astrometry

5.8.1. WIDE-ANGLE ASTROMETRY The Mk III Interferometer was designed specifically with the goal of performing wide-angle astrometry (Shao et al. 1988, 1990), i.e., measurements of stellar positions, proper motions, and parallaxes. The phase noise due to atmospheric fluctuations can be reduced by a factor ~3 to 10 with simultaneous observations in two widely separated wavelength bands (Colavita et al. 1987). The corrected "two-color" delay d is defined by

$$d \equiv d_{\mathrm{red}} - \frac{n_{\mathrm{red}} - 1}{n_{\mathrm{blue}} - n_{\mathrm{red}}} \left(d_{\mathrm{blue}} - d_{\mathrm{red}} \right), \tag{19}$$

where n_{red} and n_{blue} are the refractive indices of air in the two wavelength bands. The two-color method is limited by water vapor fluctuations because the wavelength dependence of n for water vapor is different from that for dry air. Observations at three wavelengths could in principle overcome this limitation, but the photon noise in a three-color delay estimator is prohibitively large (Hummel et al. 1994b).

The target list of the astrometric measurements with the Mk III included 11 FK5 stars, which were observed repeatedly over a period of 4 years (Hummel et al. 1994b). An accuracy of 13 mas in declination and 23 mas in right ascension was achieved, although the astrometric solution has a two-fold degeneracy. (The zero point of right ascension is indeterminate as a matter of principle, and the Mk III data have an additional near-degeneracy in declination because of the limited declination range spanned by the program stars.) The goal of the astrometric program of NPOI is the maintenance of the HIPPARCOS reference frame, which is constantly degrading owing to the uncertainties in the proper motions. First results have been reported by Hutter et al. (1998). In spite of recent progress towards the NPOI design goal of 2 mas accuracy, the future of wide-angle astrometry from the ground is somewhat doubtful in light of the planned capabilities of astrometric missions such as DIVA (Bastian et al. 1996, Röser 1999), FAME (Horner et al. 2001), and GAIA (Lindegren & Perryman 1996, Gilmore et al. 2001).

5.8.2. NARROW-ANGLE ASTROMETRY It was pointed out by Shao & Colavita (1992a) that very high accuracy can be achieved in ground-based differential astrometric measurements between stars separated by at most a few tens of arcseconds on the sky. In the narrow-angle regime $\theta \ll B/h$ the error σ for simultaneous differential observations due to atmospheric turbulence scales with θ, $B^{-2/3}$, and $t^{-1/2}$. (θ is the angular separation of the two stars, B the baseline length, h the effective height of the turbulence, and t the integration time.) For excellent sites such as Mauna Kea and Cerro Paranal, the atmospheric limit expected from Kolmogorov turbulence models is ~20 μas for $\theta = 10''$, $B = 100$ m, and $t = 1800$ s (Shao & Colavita 1992a, von der Lühe et al. 1995); an outer scale of the turbulence, $L_0 \lesssim B$, would give an even more favorable estimate. These calculations have been confirmed by observations of the long-period binary star Castor ($\theta = 3\overset{''}{.}3$) with the

Mk III Interferometer ($B = 12$ m), which had been modified for simultaneous observations of the fringe packets of both binary components (Colavita 1994). As a by-product of these measurements, the separation and position angle at epoch B1992.9589 were determined to be $3''\!.281 \pm 0''\!.01$ and $73.^\circ 23 \pm 0.^\circ 15$. In addition to the atmospheric phase fluctuations, the error budget for differential astrometry includes contributions from photon and sky noise, which limit the precision of phase measurements from instrumental errors and from differential atmospheric refraction (Gubler & Tytler 1998). A precision of 100 μas has been achieved for measurements of the two components in the 61 Cyg system ($\theta = 31''$) with the Palomar Testbed Interferometer ($B = 110$ m) over a 1-week period (Lane et al. 2000a); over 70 days the rms precision is 170 μas.

The most important scientific driver for the development of narrow-angle interferometric astrometry is the potential to detect extrasolar planets (see Marcy & Butler 1998 and Quirrenbach 2000b for recent reviews). Astrometry is complementary to the successful radial velocity technique in many respects. Interferometric observations of stars with planets known from the radial velocity surveys can yield the inclination i of the orbit and therefore the planet's mass m, whereas only $m \sin i$ can be determined spectroscopically. For stars with multiple planets it is possible to determine whether their orbits are coplanar, an important clue to the dynamical history of the system. The different detection biases with respect to the orbital semimajor axis a of the radial velocity method (signal $\propto a^{-1/2}$) and astrometry (signal $\propto a$) imply sensitivity to different architectures of the planetary system. The radial velocity surveys favor systems with masses increasing with orbital semimajor axis [e.g., υ And (Butler et al. 1999)], whereas astrometric searches should preferentially detect systems in which the masses decrease with orbital radius. The combined results of both techniques will therefore give us a better picture about the types of systems realized by nature. Interferometric astrometry will also enable a census of planets around stars of all spectral types, including pre–main-sequence objects, an important step towards the understanding of planetary system formation (Boss 1998).

An important consideration for narrow-angle astrometry is the sky coverage, i.e., the probability of finding suitable astrometric references near the target stars. Because the targets for planet searches are bright, they can be used to phase the interferometer, enabling long coherent integrations and therefore the use of relatively faint stars as astrometric references, provided that the separation between the two stars is smaller than the isoplanatic angle θ_0. The $\lambda^{6/5}$ scaling of θ_0, as well as the desire to avoid resolving the target star, favors observations in the near-IR. Reference stars with IR magnitudes $\lesssim 17$ are needed to keep the photon and sky noise from dominating the astrometric error budget. Because of the substantial variations in stellar density with galactic latitude and inhomogeneities due to clusters, etc., estimates of the sky coverage based on catalog cross-correlation are preferred over statistical models. For example, cross-correlating the HIPPARCOS catalog with USNO-A1.0 yields 734 stars with declination $\delta \leq +20^\circ$ and parallax $\pi \geq 20$ mas that have 3 potential references within a radius of $30''$ (Quirrenbach

2000a); targets for a planet search with the VLT Interferometer could be drawn from this sample. An alternative strategy is searching for planets in double stars. The Washington Double Star Catalog contains 745 F, G, and K main-sequence stars with $\delta \leq +20^\circ$ and $V \leq 10$ in pairs with separation $5'' \leq \theta \leq 20''$; 23 of these are G main-sequence stars with $V \leq 7.5$ (Quirrenbach 2000a). The technical advantage of a bright reference star at a small separation, which minimizes the instrumental and atmospheric errors, may outweigh the difficulty of determining to which one of the two stars any detected planet belongs.

Astrometric programs with targets that are not bright enough for fringe tracking (i.e., fainter than $K \approx 13$) will normally have to be carried out from space because the probability of finding a reference for phasing is exceedingly small. Therefore, observations of gravitational microlensing events (Boden et al. 1998, Paczyński 1998) will have to be carried out with the Space Interferometry Mission, unless fringe tracking at $K \approx 16$ can be achieved with the large apertures of the Keck Interferometer or VLTI. Another fundamental limitation of narrow-angle astrometry is the inability to measure parallaxes. This is because "background stars" with negligible parallax do not exist at the relevant level of accuracy and because the parallax ellipses have identical axial ratio and orientation in a small patch of sky. (There may be a few exceptional cases such as observations near quasars or stars projected towards the Magellanic Clouds.) Precise parallax measurements therefore require wide-angle measurements and going to space. There is considerable overlap in the scientific capabilities of scanning astrometric missions (DIVA, FAME, GAIA) and the pointed satellite Space Interferometry Mission (SIM) in areas such as Cepheid parallaxes, Galactic structure, dynamics of open and globular clusters, and cluster distances. In general, the scanning satellites will have a clear advantage where large numbers of targets are required, whereas SIM will get the most precise results at the faint end of the magnitude range ($V \approx 20$). SIM will also excel for astrometric planet searches, because of its capability to perform many visits of the same star with optimized temporal sampling and its superior performance in the narrow-angle regime. It is quite reassuring, however, that planet searches, probing the Galactic potential with halo streamers (e.g., Zhao et al. 1999), and other important projects can be carried out with two different techniques.

6. CONCLUSION

Optical interferometry is an extremely active field, both technically and scientifically. The results that could be reviewed in the last section of this article have mostly been obtained with single-baseline instruments operating with small apertures in the single-r_0 regime. Still, important contributions have been made to a wide range of topics in stellar astrophysics. The technical developments described in the earlier sections will undoubtedly enable observing programs with an even more profound impact. To bring high-resolution optical astronomy to full fruition it will be necessary to make a wide community aware of the unique opportunities and challenges of this field and to train a generation of young astronomers in the judicious and competent use of interferometric methods. Instruments with large apertures and astrometric capabilities, multielement arrays, and space-borne facilities will provide many rewarding opportunities for ground-breaking observations and new discoveries.

ACKNOWLEDGMENTS

I thank my former Mk III colleagues at the US Naval Observatory, Naval Research Laboratory, and Jet Propulsion Laboratory for teaching me most of what I know about optical interferometry. I have also benefited tremendously from many discussions with the staff and users community of the European Southern Observatory. Thanks are due to Peter Lawson and Theo ten Brummelaar for their careful reading of the manuscript. The interferometry literature list at http://huey.jpl.nasa.gov/olbin maintained by Peter Lawson has been very helpful.

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