1 Lecture, 2 September 1999

1.1 Observational astronomy

Virtually all of our knowledge of astronomical objects was gained by observation of their light. We know how to make many kinds of detailed measurements on light, among them brightness, spectrum, polarization state, degree of coherence, and the shape of light-emitting regions. We also know a great deal about the mechanisms by which light can be produced in nature, and we can thus use any accurate measurements we can obtain on light to characterize the physical state of astronomical objects.

The design and construction of the equipment with which we make these measurements is usually a pretty complicated activity. It demands of its practitioners a good deal of expertise in many areas of physics and engineering besides that necessary in the analysis of astronomical results. Thus the development of observational-astronomical tools has become an important specialty area within astrophysics. Here is a list of some of the problems that provide challenges to astronomers who design instruments and telescopes.

- *The overwhelming majority of interesting astronomical objects are faint.* There are still interesting, unsolved problems involving the Sun and other bright, relatively nearby objects, but most astronomical objects are quite dim. Many examples are shown in the section of the Hubble Deep Field displayed in Figure 1.1: the faintest galaxies detected in this image have apparent visual magnitude 30, about a factor of 4×10^9 fainter than can be seen with an unaided eye. This represents the present state of the art in faintness of detected objects at visible wavelengths. We confidently expect this part of the sky – and all others, too – to contain galaxies and protogalaxies even fainter and more distant than these, the

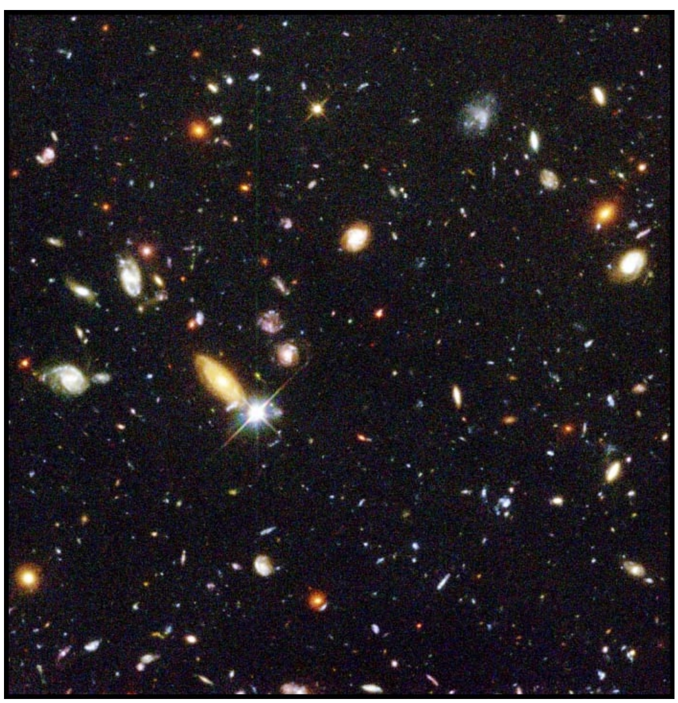

Figure 1.1: three-color image of part of the northern Hubble Deep Field (NASA/Space Telescope Science Institute)

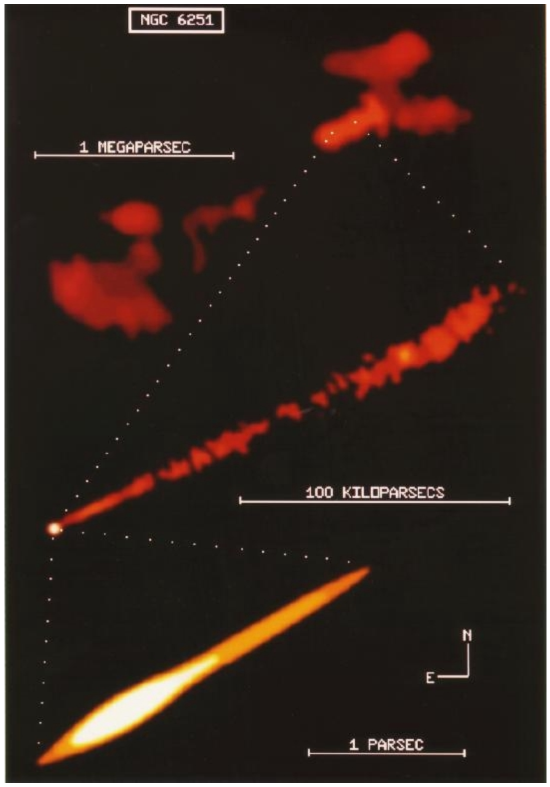

Figure 1.2: Three radio-frequency images of the radio glaxy NGC 6251, zooming in on the galactic nucleus with a wide range of magnification.. The distance to this galaxy is 79 Mpc assuming a Hubble constant of $65 \text{ km s}^{-1} \text{ Mpc}^{-1}$; 10^{-3} arcseconds corresponds to 0.38 parsecs at this distance. (A.C.S. Readhead and co-workers, Caltech)

detection of which is necessary for the understanding of galaxy formation in the early Universe. To achieve results like the Hubble Deep Field, and continually to improve the state of the art, astronomers contend with the fundamental limitations to sensitivity of detection: the various forms of *noise*, in the detected light itself and in the detection process.

- *Astronomical objects typically exhibit structure on a wide range of scales,* and many interesting astronomical objects have angular size smaller than the blur introduced by telescopes, instruments and Earth's atmosphere. Objects can subtend relatively large angles, yet have widespread sites with small sizes for which the understanding is crucial to the understanding of the whole. The smallest angular scales of importance are very frequently found to be smaller than the blur introduced in the images by the optics or the atmosphere. Consider for example the radio-frequency images of the active, elliptical galaxy NGC 6251, shown in Figure 1.2. Radio emission from NGC 6251 is distributed continuously over well over a degree of sky – far more than the visible extent of the galaxy – yet exhibits compact structures at scales as small as the limits of the measurements, in this case about 3×10^{-3} arcseconds. Also present are features compact in one dimension and long in another, like the jet shown in the central panel of Figure 1.2. Yet even these images are far from revealing the smallest interesting scales around the massive black hole thought to reside at the galaxy's center, which can be judged from the size of this black hole's event horizon to be 10^{-6} arcseconds or less.

- *The measurable spectra of astronomical objects often cover many orders of magnitude in wavelength, and exhibit structure on a wide range of bandwidth.* Continuous emission over a large fraction of the electro-

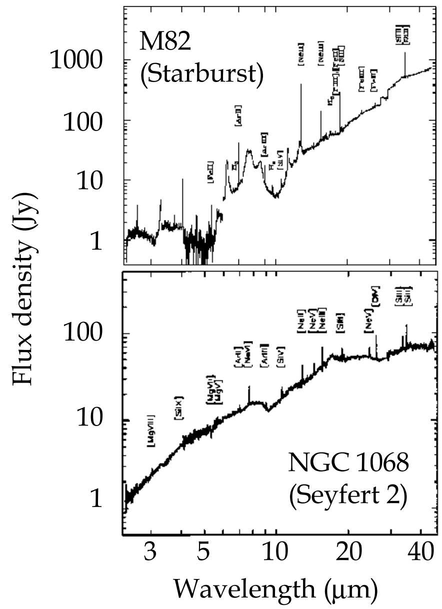

Figure 1.3: Infrared spectra of the nuclei of the starburst galaxy M82 and the active (Seyfert 2) galaxy NGC 1068. (R. Genzel and co-workers, and ESA).

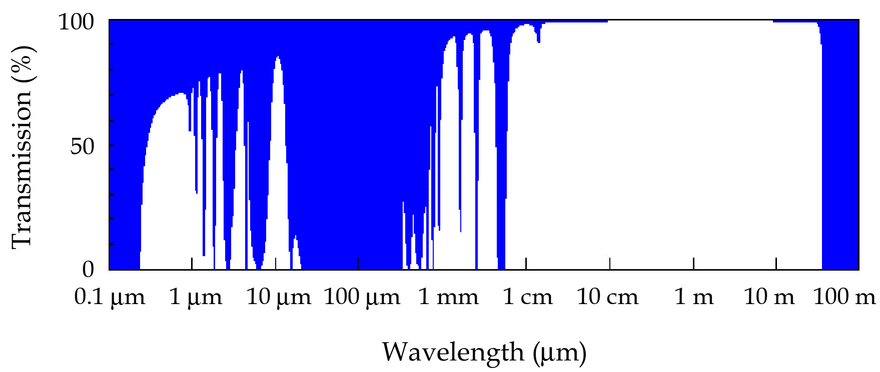

Figure 1.4: transmission of Earth's atmosphere (to approximately sea level), $\lambda = 0.1\ \mu\mathrm{m}$ - $100\ \mathrm{m}$. The transmission is zero at wavelengths shorter or longer than those plotted.

magnetic spectrum, narrow spectral lines and broad emission and absorption features are seen in most objects, as in Figure 1.3. Structure in the spectral lines themselves arises from the Doppler effect and line-of-sight differences in velocity of emitting material, and can tempt the observer to discriminate wavelengths as closely spaced as a part in a million or as wide as a part in a hundred, sometimes in the same object. Single celestial objects emit spectra with such variety, but no single device serves to reveal this variety sensitively to astronomers.

- *Earth's atmosphere is most uncooperative.* The parts of the electromagnetic spectrum for which the atmosphere is not opaque are indicated in Figure 1.4. Our atmosphere absorbs all incident radiation at wavelengths less than about 300 nm. At longer wavelengths there are "windows" in the spectrum, notably at visible wavelengths, several narrow ranges in the infrared, and a wide swath of the radio range, but all the rest is absorbed. Much of the absorption in the center of Figure 1.4 comes from wa-

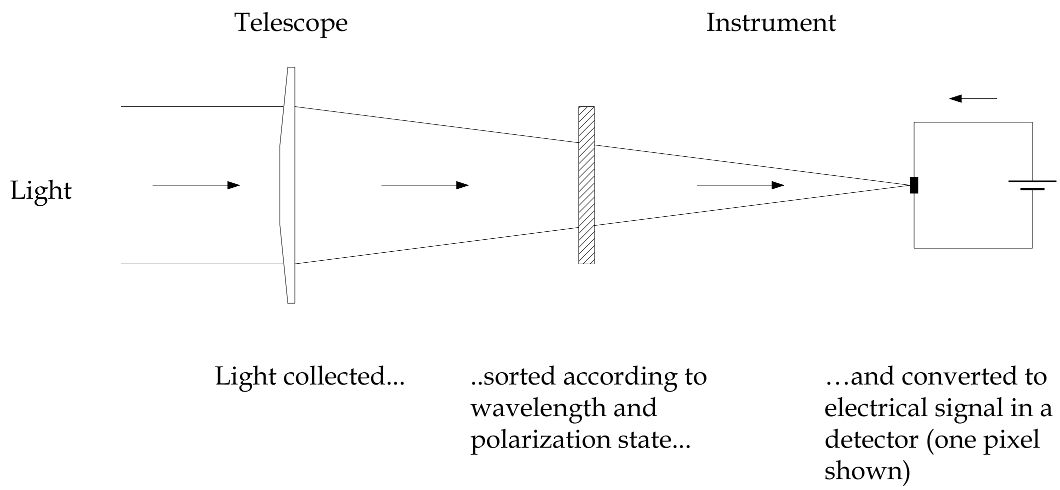

Figure 1.5: schematic diagram of a telescope-instrument combination.

ter vapor in the troposphere. The effect is less at altitudes accessible to aircraft, but the only way to avoid this absorption completely, or to observe at high energies or very long wavelengths at all, is to observe from space. The atmosphere is also turbulent; at visible and short infrared wavelengths this contributes a blur to images of celestial objects that cannot be corrected completely except by observing from above the atmosphere.

These challenges have been met remarkably successfully. The only remaining virgin range of the spectrum lies at very long radio wavelengths, beyond the ionospheric cut-off. Solutions have been found for sensitive imaging and spectroscopy at all other wavelengths of importance in astronomy.

1.2 Imaging, spectroscopy, resolution and sensitivity

In their measurements, astronomers use equipment, the components of which are usefully sorted into three categories according to task, as indicated in the diagram in Figure 1.5. First, as much light as possible is gathered, and brought to focus in an image; this of course is done with telescopes. On its way to the final focus, the physical state of the light is prepared, according to the intention of the observation: unwanted light rejected, the rest of the light sorted according properties like wavelength and polarization. This done, the remaining light is detected in a way amenable to accurate determination of radiant power as a function of wavelength, or polarization, or position in the sky. The equipment that does the selection and detection of light is generally called an *instrument*. A given telescope is usually outfitted with several instruments, individually optimized for particular wavelengths and types of observations, and matched optically to that telescope. Instruments themselves are grouped into two categories: *cameras*, which measure images of celestial objects from light within some given range of wavelengths; and *spectrographs*, which measure spectra of celestial objects lying within some range of angles on the sky.

Two-dimensional detector arrays have come into astronomical use at all but the very longest and shortest wavelengths. They generally consist of planar, rectangular grids of independent semiconducting or superconducting detectors (called *pixels*). Detector arrays are placed in the final focal plane of the telescope-instrument optics, and play the same role that photographic plates used to play on visible-wavelength telescopes. Their two dimensions are used in two different ways, as indicated in Figure 1.6. Most obvious is to use the dimensions as an ordinary camera does, to represent position within an image. The other way, used in instruments with *dispersive* optical elements like diffraction gratings, is to restrict the image to a one-dimensional *slit* and to use the other array dimension for wavelength. When one

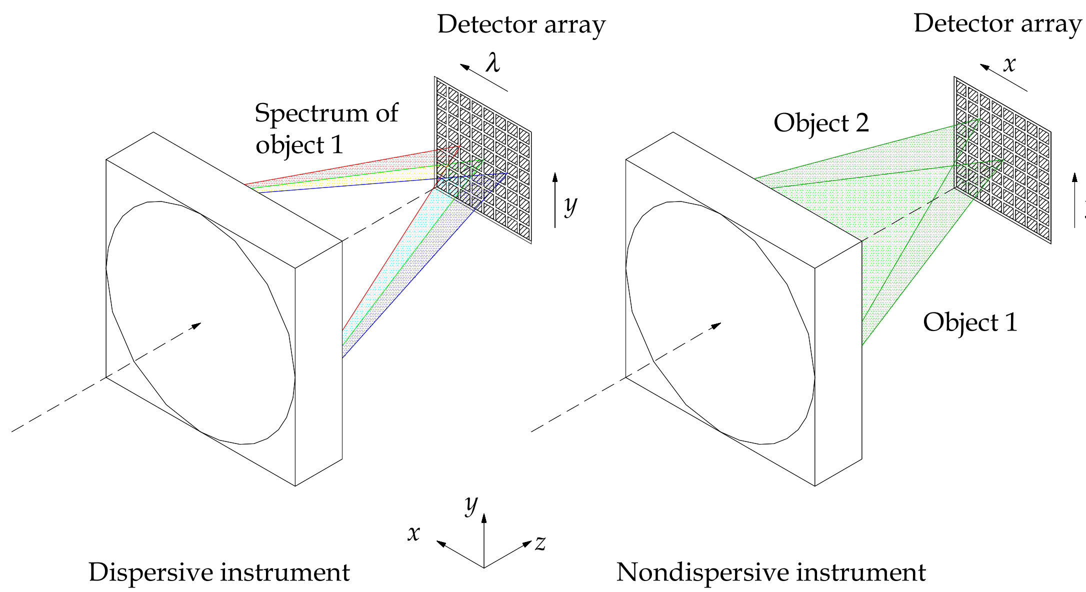

Figure 1.6: how two-dimensional detector arrays can be used for spectroscopy and imaging. On the left is a schematic diagram of a dispersive instrument that feeds the detector the spectrum of a one-dimensional image, or "slit," that contains a certain object. At right is depicted a nondispersive, imaging instrument that passes a certain band of wavelengths and forms a two-dimensional image of a field containing two objects.

matches optically a telescope and instrument one accounts for the detector array length and width, distance between pixels, field of view and spectral range to be covered, and conversion by the optics of angles to position in the final focal plane. For nondispersive instruments the relevant conversion is between angles on the sky to position on the detector, a factor called the *plate scale*. Wavelength-dependent angular deviations are introduced to light within dispersive instruments, and the match of this light to the final focal plane involves the conversion between angle and wavelength, called the *dispersion*.

A combination of telescope and instrument is designed to carry out certain kinds of observations on certain types of astronomical objects. Estimates of the properties of the objects are used to specify the properties of the system: the angular field of view and range of wavelengths to be covered, the smallest angular size and spectral interval (bandwidth) that is wished to be revealed, and the range of brightness that the celestial objects present within these fields and intervals. After the designed system is built, measurements are made of its properties to determine the limits to its performance. The lower bounds on the angular and spectral detail, and object brightness, that a system is capable of revealing are perhaps its most important properties. The first two bounds are called the *angular* (or *spatial*) *resolution* and *spectral resolution* of the system. Angular resolution is usually expressed as the angular size of the smallest discernable detail in an image. Here we define angular resolution to be the angle corresponding to the full width to half the maximum signal (or *FWHM*) for the focal-plane image of a point-like object. In most observing situations, stars are too small and distant for our telescopes and instruments to reveal their details, so for wavelengths at which they can be detected stars serve nicely as point-like objects, and the FWHM of a stellar images are used to determine the angular resolution $\Delta\theta$. Similarly, we define spectral resolution to be the FWHM in frequency or wavelength of the signal from a single frequency or wavelength – a spectral line of infinitesimal width. In practice, one could use system measurements on laboratory single frequency oscillators like lasers – which have finite, but very narrow, linewidths – to determine spectral resolution. We normally express spectral resolution as the dimensionless ratio of FWHM to the center frequency or wavelength of the spectral line used in the measurement, $\Delta\lambda/\lambda = \Delta\nu/\nu$. Small values of $\Delta\theta$, $\Delta\lambda/\lambda$ or $\Delta\nu/\nu$ correspond to small details visible in images or spectra; either case is (misleadingly) referred to as *high resolution* in the argot of astronomy. The fundamental limitations to angular and spectral resolution of telescope-instrument systems are determined respectively by the size of the light-gathering or dispersive optics relative to the wavelength of light, as we will find in lectures 12-15 and 24-28.

The lower bound on brightness that can be measured by a system is called the system's *sensitivity*: one system capable of detecting fainter objects than a second system is (again misleadingly) called the more sensitive of the two. Sensitivity is expressed as the smallest power, flux (i.e. power per unit area) or flux density (i.e. flux per unit spectral bandwidth) that a system is capable of detecting in a standard exposure time. It is generally measured by observing an object of known or calculable power, and measuring both the average and the *fluctuations* in signal for a large number of exposure-time intervals. By scaling the signal power (or flux, or flux density) by the measured ratio of signal and *noise*, the latter represented by the standard deviation of these fluctuations, one obtains the *noise equivalent power* or *NEP.* (Analogously, one can obtain a noise-equivalent flux or flux density.) *NEP* and its relatives are commonly used to represent sensitivity of systems. The noise that provides the fundamental limitation to sensitivity comes most often from *background* sources of light, as we will discuss in lectures 19-23.

