3 Observations and Instruments

Up to the end of the Middle Ages, the most important means of observation in astronomy was the human eye. It was aided by various mechanical devices to measure the positions of celestial bodies in the sky. The telescope was invented in Holland at the beginning of the 17th century, and in 1609 Galileo Galilei made his first astronomical observations with this new instrument. Astronomical photography was introduced at the end of the 19th century, and during the last few decades many kinds of electronic detectors have been adopted for the study of electromagnetic radiation from space. The electromagnetic spectrum from the shortest gamma rays to long radio waves can now be used for astronomical observations.

3.1 Observing Through the Atmosphere

With satellites and spacecraft, observations can be made outside the atmosphere. Yet, the great majority of astronomical observations are carried out from the surface of the Earth. In the preceding chapter, we discussed refraction, which changes the apparent altitudes of objects. The atmosphere affects observations in many other, more serious ways as well. Real problems are the weather, restlessness of the atmosphere, and, worst of all, opacity of the atmosphere in many wavelength regions.

The air is never quite steady, and there are layers with different temperatures and densities; this causes convection and turbulence. When the light from a star passes through the unsteady air, rapid changes in refraction in different directions result. Thus, the amount of light reaching a detector, e.g. the human eye, constantly varies; the star is said to *scintillate* (Fig. 3.1). Planets shine more steadily, since they are not point sources like the stars, and variations in the light coming from different parts partly cancel each others.

When a wavefront passes an aperture its amplitude and phase vary with time. The amplitude variations cause scintillation and phase changes blur the image. A telescope collects light over a larger area, which evens out rapid changes and diminishes scintillation. Instead, differences in the phase do not cancel out the same way; they smear the image and point sources are seen in telescopes as vibrating speckles. This phenomenon is called *seeing*. When the exposure time is at least a few seconds the light oscillates in all directions, and the stellar image becomes a round disk (Fig. 3.3). The diameter of the seeing disk can vary from less than an arc second to even tens of arc seconds. If the size of the disk small, the seeing is good. It is only in good conditions on high mountains that the seeing is less than one arc second.

Seeing limits how small details can be seen with a telescope. Sometimes the atmosphere can calm down for a short period of time, and then it is possible to glimpse visually details that are not seen in pictures exposed for a longer time.

Some wavelength regions in the electromagnetic spectrum are strongly absorbed by the atmosphere (Fig. 3.2). The most important transparent

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interval is the *optical window* from about 300 to 800 nm. This interval coincides with the region of sensitivity of the human eye (about 400–700 nm).

At wavelengths under 300 nm absorption by atmospheric ozone prevents radiation from reaching the ground. The ozone is concentrated in a thin layer at a height of about 20–30 km, and this layer protects the Earth from harmful ultraviolet radiation. At still shorter wavelengths, the main absorbers are O_2, N_2 and free atoms. Nearly all of the radiation under 300 nm is absorbed by the upper parts of the atmosphere.

At wavelengths longer than visible light, in the near-infrared region, the atmosphere is fairly transparent up to 1.3 µm. There are some absorption belts caused by water and molecular oxygen, but the atmosphere gets more opaque only at wavelengths of longer than 1.3 µm. At these

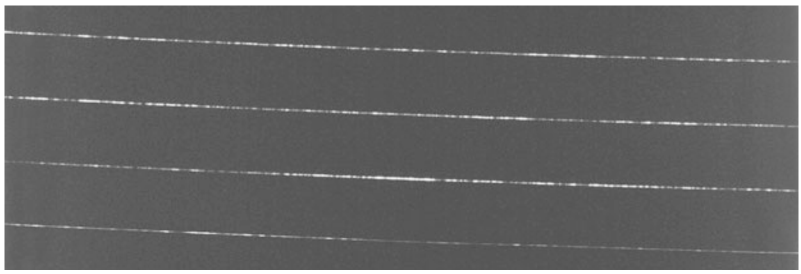


Fig. 3.1 Scintillation of Sirius during four passes across the field of view. The star was very low on the horizon. (Photo by Pekka Parviainen)

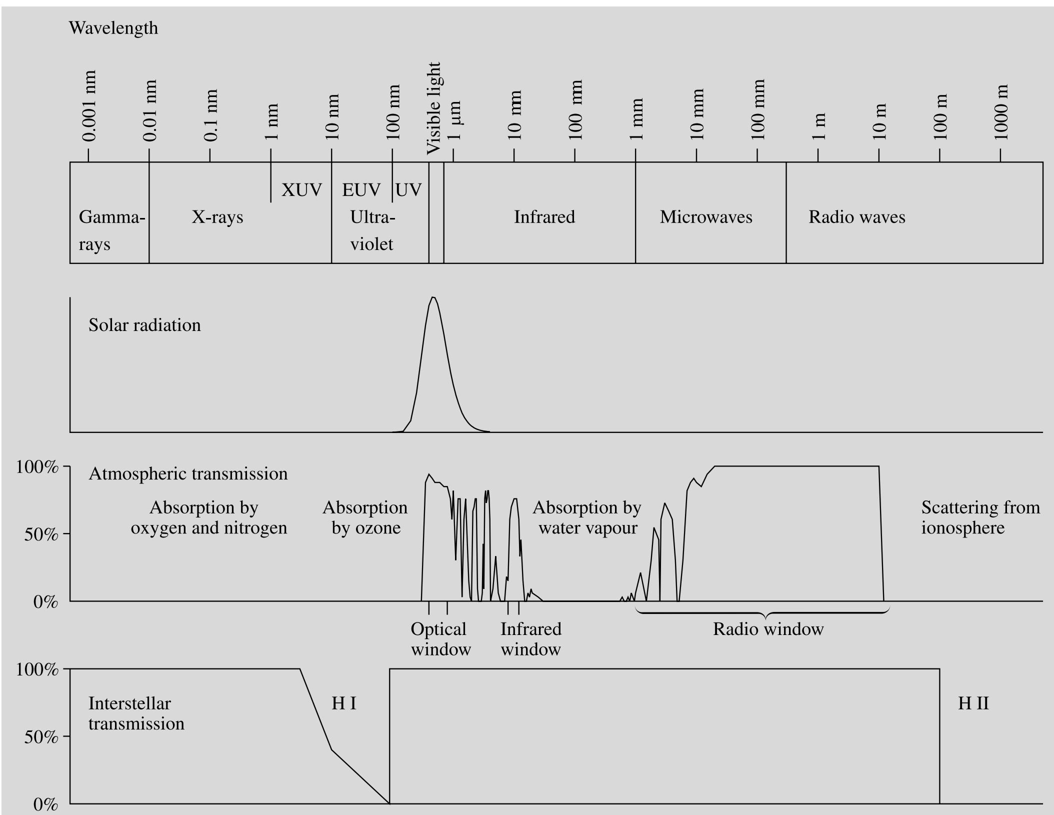


Fig. 3.2 The transparency of the atmosphere at different wavelengths. 100 % transmission means that all radiation reaches the surface of the Earth. The radiation is also absorbed by interstellar gas, as shown in the lowermost very schematic figure. The interstellar absorption also varies very much depending on the direction (Chap. 15)

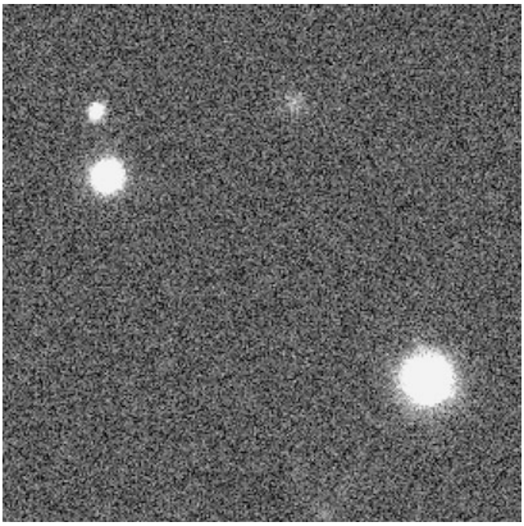


Fig. 3.3 Stellar images vibrate about equally often in all directions, on the average. If the picture is exposed for more than about one second the image of the star spreads into a circular seeing disc. The better the observing conditions are the smaller is the diameter of the disc

Fig. 3.4 Night views from the top of Mount Wilson. *The upper photo* was taken in 1908, *the lower one* in 1988. The lights of Los Angeles, Pasadena, Hollywood and more than 40 other towns are reflected in the sky, causing considerable disturbance to astronomical observations. (Photos by Ferdinand Ellerman and International Dark-Sky Association)

wavelengths, radiation reaches the lower parts of the atmosphere only in a few narrow windows. All wavelengths between 20 μm and 1 mm are totally absorbed.

At wavelengths longer than 1 mm, there is the *radio window* extending up to about 20 m. At still longer wavelengths, the ionosphere in the upper parts of the atmosphere reflects all radiation (Fig. 3.2). The exact upper limit of the radio window depends on the strength of the ionosphere, which varies during the day. (The structure of the atmosphere is described in Chap. 7.)

At optical wavelengths (300–800 nm), light is scattered by the molecules and dust in the atmosphere, and the radiation is attenuated. Scattering and absorption together are called *extinction*. Extinction must be taken into account when one measures the brightness of celestial bodies (Chap. 4).

In the 19th century Lord Rayleigh succeeded in explaining why the sky is blue. Scattering caused by the molecules in the atmosphere is inversely proportional to the fourth power of the wavelength. Thus, blue light is scattered more than red light. The blue light we see all over the sky is scattered sunlight. The same phenomenon colours the setting sun red, because owing to the long, oblique path through the atmosphere, all the blue light has been scattered away.

In astronomy one often has to observe very faint objects. Thus, it is important that the background sky be as dark as possible, and the atmosphere as transparent as possible. That is why the large observatories have been built on mountain tops far from the cities. The air above an observatory site must be very dry, the number of cloudy nights few, and the seeing good.

Astronomers have looked all over the Earth for optimal conditions and have found some exceptional sites. In the 1970's, several new major observatories were founded at these sites. Among the best sites in the world are: the extinct volcano Mauna Kea on Hawaii, rising more than 4000 m above the sea; the dry mountains in northern Chile; the Sonoran desert in the U.S., near the border of Mexico; and the mountains on La Palma, in the Canary Islands. Many older observatories are severely plagued by the lights of nearby cities (Fig. 3.4).

In radio astronomy atmospheric conditions are not very critical except when observing at the shortest wavelengths, damped by atmospheric moisture. Constructors of radio telescopes have much greater freedom in choosing their sites than optical astronomers. Still, radio telescopes are also often constructed in uninhabited places to isolate them from disturbing radio and television broadcasts.

3.2 Optical Telescopes

The telescope fulfils three major tasks in astronomical observations:

1. It collects light from a large area, making it possible to study very faint sources.
2. It improves resolution and increases the apparent angular diameter of the object.
3. It is used to measure the positions of objects.

The light-collecting surface in a telescope is either a lens or a mirror. Thus, optical telescopes are divided into two types, lens telescopes or *refractors* and mirror telescopes or *reflectors* (Fig. 3.5).

Geometrical Optics Refractors have two lenses, the *objective* which collects the incoming light and forms an image in the focal plane, and the *eyepiece* which is a small magnifying glass for looking at the image (Fig. 3.5). The lenses are at the opposite ends of a tube which can be directed towards any desired point. The distance between the eyepiece and the focal plane can be adjusted to get the image into focus. The image formed by the objective lens can also be registered, e.g. on a photographic film, as in an ordinary camera.

The diameter of the objective, D, is called the *aperture* of the telescope. The ratio of the aperture D to the focal length f, $F = D/f$, is called the *aperture ratio*. This quantity is used to characterise the light-gathering power of the telescope. If the aperture ratio is large, near unity, one has a powerful, "fast" telescope; this means that one can take photographs using short exposures, since the image is bright. A small aperture ratio (the focal length much greater than the aperture) means a "slow" telescope.

In astronomy, as in photography, the aperture ratio is often denoted by f/n (e.g. $f/8$), where n is the focal length divided by the aperture. For fast telescopes this ratio can be $f/1 \ldots f/3$, but usually it is smaller, $f/8 \ldots f/15$.

The *scale* of the image formed in the focal plane of a refractor can be geometrically determined from Fig. 3.6. When the object is seen at the angle u, it forms an image of height s,

$$s = f \tan u \approx f u, \tag{3.1}$$

since u is a very small angle. If the telescope has a focal length of, for instance, 343 cm, one arc minute corresponds to

$$\begin{aligned} s &= 343\ \text{cm} \times 1' \\ &= 343\ \text{cm} \times (1/60) \times (\pi/180) \\ &= 1\ \text{mm}. \end{aligned}$$

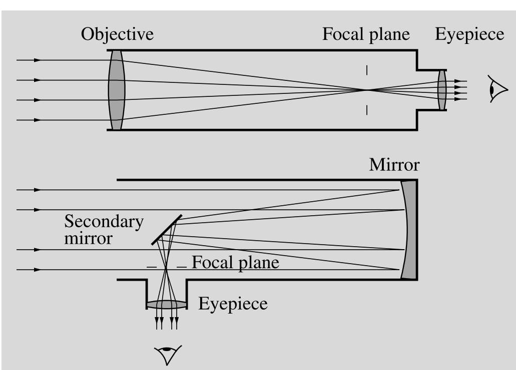


Fig. 3.5 A lens telescope or refractor and a mirror telescope or reflector

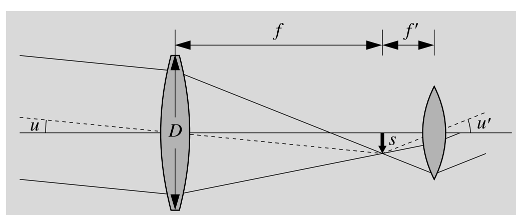


Fig. 3.6 The scale and magnification of a refractor. The object subtends an angle u. The objective forms an image of the object in the focal plane. When the image is viewed through the eyepiece, it is seen at an angle u'

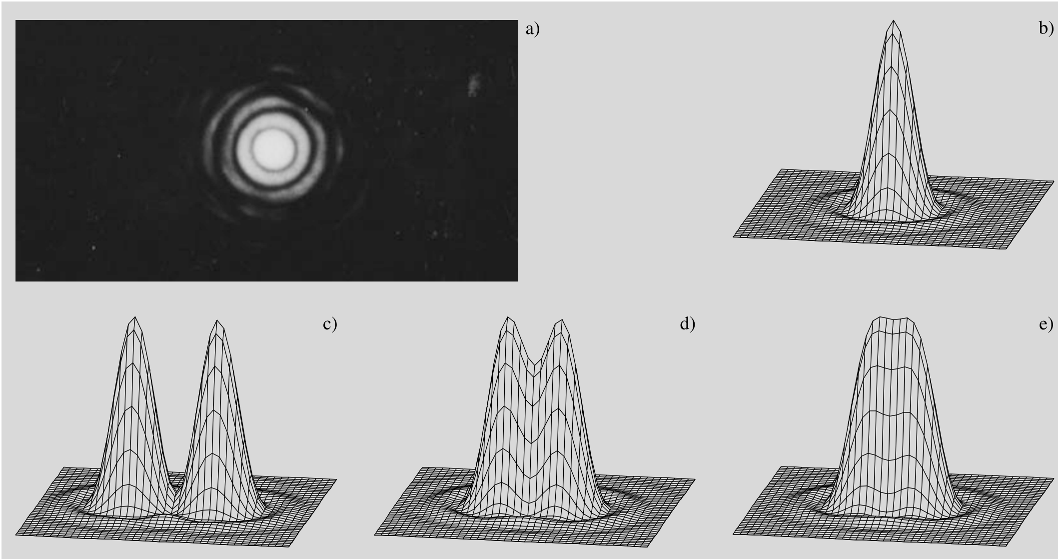


Fig. 3.7 Diffraction and resolving power. The image of a single star (**a**) consists of concentric diffraction rings, which can be displayed as a mountain diagram (**b**). Wide pairs of stars can be easily resolved (**c**). For resolving close binaries, different criteria can be used. One is the Rayleigh limit $1.22\,\lambda/D$ (**d**). In practice, the resolution can be written λ/D, which is near the Dawes limit (**e**). (Photo (**a**) Sky and Telescope)

The *magnification* ω is (from Fig. 3.6)

$$\omega = u'/u \approx f/f', \tag{3.2}$$

where we have used the equation $s = fu$. Here, f is the focal length of the objective and f' that of the eyepiece. For example, if $f = 100$ cm and we use an eyepiece with $f' = 2$ cm, the magnification is 50-fold. The magnification is not an essential feature of a telescope, since it can be changed simply by changing the eyepiece.

A more important characteristic, which depends on the aperture of the telescope, is the *resolving power*, which determines, for example, the minimum angular separation of the components of a binary star that can be seen as two separate stars. The theoretical limit for the resolution is set by the diffraction of light: The telescope does not form a point image of a star, but rather a small disk, since light "bends around the corner" like all radiation (Fig. 3.7).

The theoretical resolution of a telescope is often given in the form introduced by Rayleigh (see Box 3.1)

$$\sin\theta \approx \theta = 1.22\,\lambda/D, \quad [\theta] = \text{rad}. \tag{3.3}$$

As a practical rule, we can say that two objects are seen as separate if the angular distance between them is

$$\theta \gtrsim \lambda/D, \quad [\theta] = \text{rad}. \tag{3.4}$$

This formula can be applied to optical as well as radio telescopes. For example, if one makes observations at a typical yellow wavelength ($\lambda = 550$ nm), the resolving power of a reflector with an aperture of 1 m is about $0.1''$. However, seeing spreads out the image to a diameter of typically one arc second. Thus, the theoretical diffraction limit cannot usually be reached on the surface of the Earth.

In photography the resolution is further restricted by the properties of the photographic plate or the pixel size of the CCD camera, decreasing the resolution as compared with visual observations. The grain size of photographic emulsions is about 0.01–0.03 mm, which is also the minimum size of the image. For a focal length of 1 m, the scale is 1 mm = $206''$, and thus 0.01 mm corresponds to about 2 arc seconds. This is similar to the theoretical resolution of a telescope with an aperture of 7 cm in visual observations. The pixel sizes of CCD cameras have decreased, and currently the resolution can be as good as or even better than the resolution of photographic films.

In practice, the resolution of visual observations is determined by the ability of the eye to see details. In night vision (when the eye is perfectly adapted to darkness) the resolving capability of the human eye is about $2'$.

The *maximum magnification* $\omega_{\max}$ is the largest magnification that is worth using in telescopic observations. Its value is obtained from the ratio of the resolving capability of the eye, $e \approx 2' = 5.8 \times 10^{-4}$ rad, to the resolving power of the telescope, θ,

$$\omega_{\max} = e/\theta \approx eD/\lambda = \frac{5.8 \times 10^{-4}\, D}{5.5 \times 10^{-7}\ \mathrm{m}} \approx D/1\ \mathrm{mm}. \tag{3.5}$$

If we use, for example, an objective with a diameter of 100 mm, the maximum magnification is about 100. If larger magnifications are used the object is seen bigger but no more details are resolved. Especially small and cheap amateur telescopes are advertised having huge magnifications, which are useless; rather they make observing very difficult. However, experienced observers may sometimes use very big magnifications in some special occasions.

The *minimum magnification* $\omega_{\min}$ is the smallest magnification that is useful in visual observations. Its value is obtained from the condition that the diameter of the *exit pupil* L of the telescope must be smaller than or equal to the pupil of the eye.

The exit pupil is the image of the objective lens, formed by the eyepiece, through which the light from the objective goes behind the eyepiece. From Fig. 3.8 we obtain

$$L = \frac{f'}{f} D = \frac{D}{\omega}. \tag{3.6}$$

Thus the condition $L \le d$ means that

$$\omega \ge D/d. \tag{3.7}$$

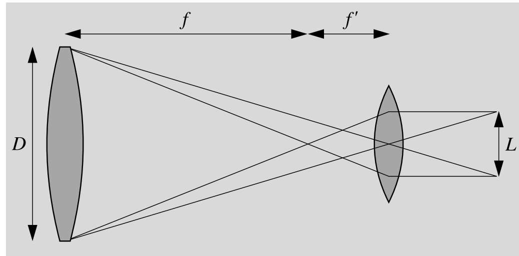


Fig. 3.8 The exit pupil L is the image of the objective lens formed by the eyepiece

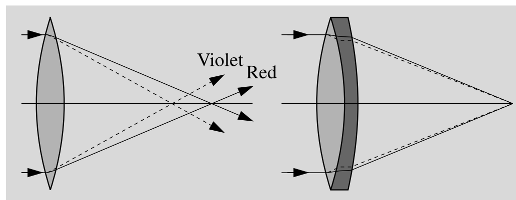


Fig. 3.9 Chromatic aberration. Light rays of different colours are refracted to different focal points (*left*). The aberration can be corrected with an achromatic lens consisting of two parts (*right*)

In the night, the diameter of the pupil of the human eye is about 6 mm, and thus the minimum magnification of a 100 mm telescope is about 17.

Refractors In the first refractors, which had a simple objective lens, the observations were hampered by the *chromatic aberration*. Since glass refracts different colours by different amounts, all colours do not meet at the same focal point (Fig. 3.9), but the focal length increases with increasing wavelength. To remove this aberration, *achromatic lenses* consisting of two parts were developed in the 18th century. The colour dependence of the focal length is much smaller than in single lenses, and at some wavelength, λ_0, the focal length has an extremum (usually a minimum). Near this point the change of focal length with wavelength is very small (Fig. 3.10). If the telescope is intended for visual observations, we choose $\lambda_0 = 550$ nm, corresponding to the maximum sensitivity of the eye. Objectives for photographic refractors were usually constructed with $\lambda_0 \approx 425$ nm, since normal photographic plates were most sensitive to the blue part of the spectrum.

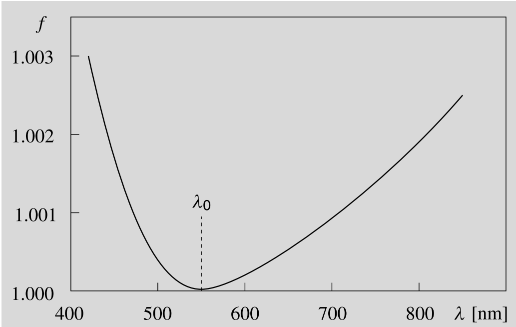


Fig. 3.10 The wavelength dependence of the focal length of a typical achromatic objective for visual observations. The focal length has a minimum near $\lambda = 550$ nm, where the eye is most sensitive. In bluer light ($\lambda = 450$ nm) or in redder light ($\lambda = 800$ nm), the focal length increases by a factor of about 1.002

By combining three or even more lenses of different glasses in the objective, the chromatic aberration can be corrected still better (as in apochromatic objectives). Also, special glasses have been developed where the wavelength dependences of the refractive index cancel out so well that two lenses already give a very good correction of the chromatic aberration. They have, however, hardly been used in astronomy so far.

The largest refractors in the world have an aperture of about one metre (102 cm in the Yerkes Observatory telescope (Fig. 3.11), finished in 1897, and 91 cm in the Lick Observatory telescope (1888)). The aperture ratio is typically $f/10 \ldots f/20$.

The use of refractors is limited by their small field of view and awkwardly long structure. Refractors are still used, e.g. in solar telescopes and various meridian telescopes for measuring the positions of stars, and also for visual observations of binary stars.

A wider field of view is obtained by using more complex lens systems, and telescopes of this kind are called *astrographs*. Astrographs have an objective made up of typically 3–5 lenses and an aperture of less than 60 cm. The aperture ratio is $f/5 \ldots f/7$ and the field of view about 5°. Astrographs are used to photograph large areas of the sky, e.g. for proper motion studies and for statistical brightness studies of the stars.

Reflectors By far the most common telescope type in astrophysical research nowadays is the mirror telescope or reflector. As a light-collecting surface, it employs a mirror coated with a thin layer of aluminium. The form of the mirror is usually parabolic. A parabolic mirror reflects all light rays entering the telescope parallel to the main axis into the same focal point. The image formed at this point can be observed through an eyepiece or registered with a detector. One of the advantages of reflectors is the absence of chromatic aberration, since all wavelengths are reflected to the same point.

In the very largest telescopes, the instruments can be installed at the *primary focus* (Fig. 3.12) without eclipsing too much of the incoming light. Sometimes even the observer could sit in a cage at the primary focus. In smaller telescopes, this is not possible, and the image must be inspected from outside the telescope. In modern telescopes instruments are remotely controlled, and the observer must stay away from the telescope to reduce thermal turbulence.

The idea of a reflector was described in 1663 by James Gregory. The first practical reflector, however, was built by Isaac Newton. He guided the light perpendicularly out from the telescope with a small flat mirror. Therefore the focus of the image in such a system is called the *Newton focus*. A typical aperture ratio of a Newtonian telescope is $f/3 \ldots f/10$. Another possibility is to bore a hole at the centre of the primary mirror and reflect the rays through it with a small hyperbolic secondary mirror in the front end of the telescope. In such a design, the rays meet in the *Cassegrain focus*. Cassegrain systems have aperture ratios of $f/8 \ldots f/15$.

The effective focal length (f_e) of a Cassegrain telescope is determined by the position and convexity of the secondary mirror. Using the notations of Fig. 3.13, we get

$$f_e = \frac{b}{a} f_p . \tag{3.8}$$

If we choose $a \ll b$, we have $f_e \gg f_p$. In this way one can construct short telescopes with long focal lengths. Cassegrain systems are especially well suited for spectrographic, photometric and other instruments, which can be mounted in the secondary focus, easily accessible to the observer.

More complicated arrangements use several mirrors to guide the light through the declination axis of the telescope to a fixed *coudé focus* (from the French word *couder*, to bend), which can even be situated in a separate room near the telescope (Fig. 3.14). The focal length is thus very long and the aperture ratio $f/30\ldots f/40$. The coudé focus is used mainly for accurate spectroscopy,

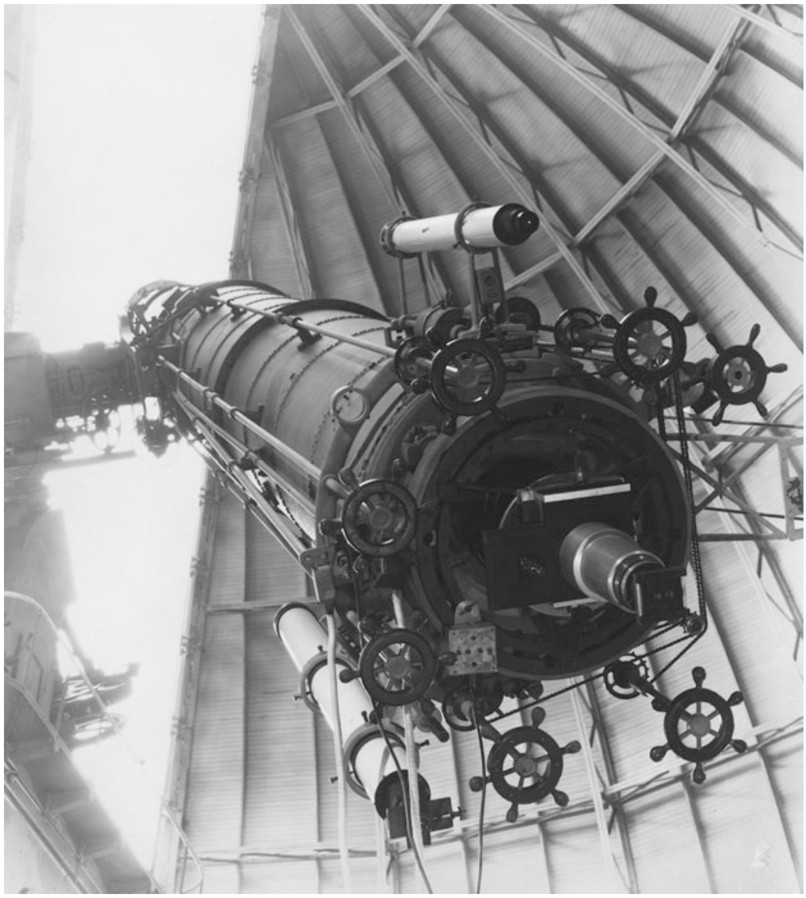


Fig. 3.11 The largest refractor in the world is at the Yerkes Observatory, University of Chicago. It has an objective lens with a diameter of 102 cm. (Photo by Yerkes Observatory)

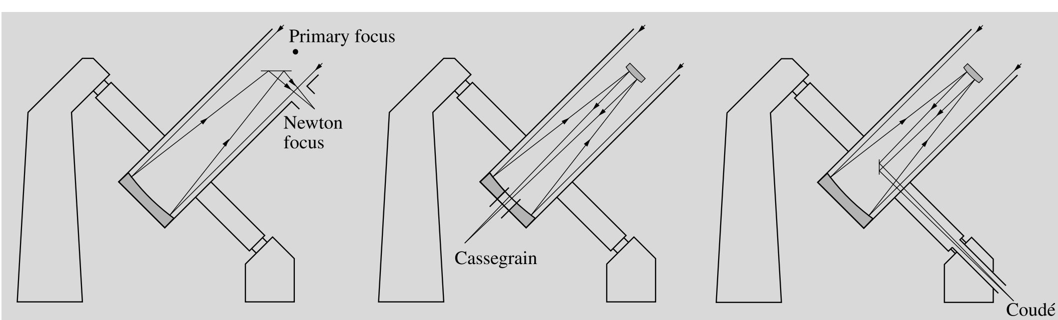


Fig. 3.12 Different locations of the focus in reflectors: primary focus, Newton focus, Cassegrain focus and coudé focus. The coudé system in this figure cannot be used for observations near the celestial pole. More complex coudé systems usually have three flat mirrors after the primary and secondary mirrors

since the large spectrographs can be stationary and their temperature can be held accurately constant. Nowadays this can also be achieved by using optical fibers. A drawback is that much light is lost in the reflections in the several mirrors of the coudé system. An aluminised mirror reflects about 80 % of the light falling on it, and thus in a coudé system of, e.g. five mirrors (including the primary and secondary mirrors), only $0.8^5 \approx 30$ % of the light reaches the detector. In new big telescopes the large instruments are usually mounted at the *Nasmyth focus* at the end of the horizontal axis of the telescope.

The reflector has its own aberration, *coma*. It affects images displaced from the optical axis. Light rays do not converge at one point, but form a figure like a comet. Due to the coma, the classical reflector with a paraboloid mirror has a very small correct field of view. The coma limits the diameter of the useful field to 2–20 minutes of arc, depending on the aperture ratio of the telescope. The 5 m Palomar telescope, for instance, has a useful field of view of about 4′, corresponding to about one-eighth of the diameter of the Moon. In practice, the small field of view can be enlarged by various correcting lenses.

If the primary mirror were spherical, there would be no coma. However, this kind of mirror has its own error, *spherical aberration*: light rays from the centre and edges converge at different points. To remove the spherical aberration, the Estonian astronomer *Bernhard Schmidt* developed a thin correcting lens that is placed in

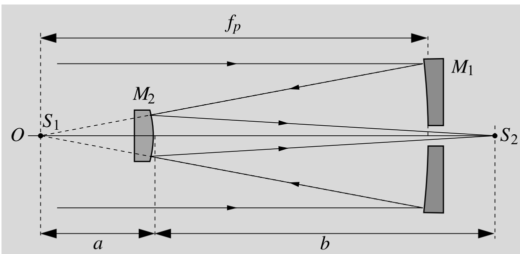


Fig. 3.13 The principle of a Cassegrain reflector. A concave (paraboloid) primary mirror M_1 reflects the light rays parallel to the main axis towards the primary focus S_1. A convex secondary mirror M_2 (hyperboloid) reflects the rays back through a small hole at the centre of the main mirror to the secondary focus S_2 outside the telescope

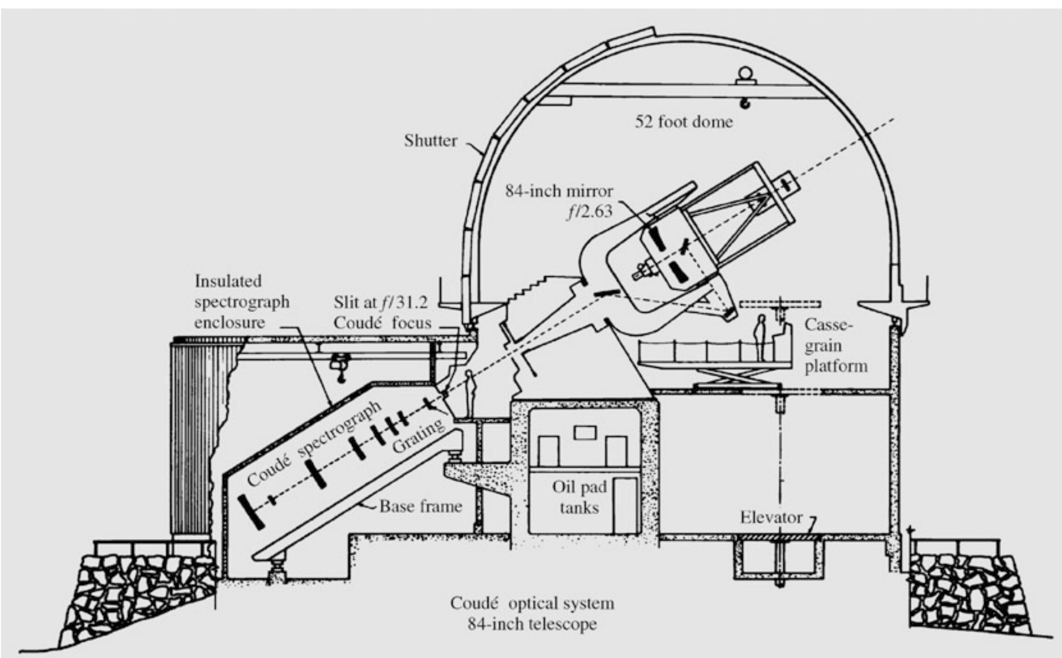


Fig. 3.14 The coudé system of the Kitt Peak 2.1 m reflector. (Drawing National Optical Astronomy Observatories, Kitt Peak National Observatory)

the way of the incoming light. Schmidt cameras (Figs. 3.15 and 3.16) have a very wide (about 7°), nearly faultless field of view, and the correcting lens is so thin that it absorbs very little light. The images of the stars are very sharp. In Schmidt telescopes the diaphragm with the correcting lens is positioned at the centre of the radius of curvature of the mirror (this radius equals twice the focal length). To collect all the light from the edges of the field of view, the diameter of the mirror must be larger than that of the correcting glass. The Palomar Schmidt camera, for example, has an aperture of 122 cm (correcting lens)/183 cm (mirror) and a focal length of 300 cm. The largest Schmidt telescope in the world is in Tautenburg, Germany, and its corresponding values are 134/203/400 cm.

A disadvantage of the Schmidt telescope is the curved focal plane, consisting of a part of a sphere. When the telescope is used for photography, the plate must be bent along the curved focal plane. Another possibility of correcting the curvature of the field of view is to use an extra correcting lens near the focal plane. Such a solution was developed by the Finnish astronomer Yrjö Väisälä in the 1930's, independently of Schmidt. Such a system is often called a *Schmidt–Väisälä telescope*. The chips of CCD cameras are usually much smaller than photographic plates, and then the curved focal surface is not a problem.

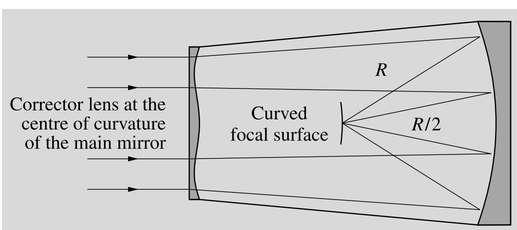


Fig. 3.15 The principle of the Schmidt camera. A correcting glass at the centre of curvature of a concave spherical mirror deviates parallel rays of light and compensates for the spherical aberration of the spherical mirror. (In the figure, the form of the correcting glass and the change of direction of the light rays have been greatly exaggerated.) Since the correcting glass lies at the centre of curvature, the image is practically independent of the incoming angle of the light rays. Thus there is no coma or astigmatism, and the images of stars are points on a spherical surface at a distance of $R/2$, where R is the radius of curvature of the spherical mirror. In photography, the plate must be bent into the form of the focal surface, or the field rectified with a corrector lens

Schmidt cameras have proved to be very effective in mapping the sky. They have been used to photograph the Palomar Sky Atlas mentioned in the previous chapter and its continuation, the *ESO/SRC Southern Sky Atlas*.

The Schmidt camera is an example of a *catadioptric telescope*, which has both lenses and mirrors. *Schmidt–Cassegrain telescopes* used by many amateurs are modifications of the Schmidt camera. They have a secondary mirror mounted at the centre of the correcting lens; the mirror reflects the image through a hole in the primary mirror. Thus the effective focal length can be rather long, although the telescope itself is very short. Another common catadioptric telescope is the *Maksutov* telescope. Both surfaces of the correcting lens as well as the primary mirror of a Maksutov telescope are concentric spheres.

Another way of removing the coma of the classical reflectors is to use more complicated mirror surfaces. The *Ritchey–Chrétien* system has hyperboloidal primary and secondary mirrors, providing a fairly wide useful field of view. Ritchey–Chrétien optics are used in many large telescopes.

Mountings of Telescopes A telescope has to be mounted on a steady support to prevent its shaking, and it must be smoothly rotated during observations. There are two principal types of mounting, *equatorial* and *azimuthal* (Fig. 3.17).

In the equatorial mounting, one of the axes is directed towards the celestial pole. It is called the *polar axis* or *hour axis*. The other one, the *declination axis*, is perpendicular to it. Since the hour axis is parallel to the axis of the Earth, the apparent rotation of the sky can be compensated for by turning the telescope around this axis at a constant rate.

The declination axis is the main technical problem of the equatorial mounting. When the telescope is pointing to the south its weight causes a force perpendicular to the axis. When the telescope is tracking an object and turns west-

Fig. 3.16 The large Schmidt telescope of the European Southern Observatory. The diameter of the mirror is 1.62 m and of the free aperture 1 m. (Photo ESO)

ward, the bearings must take an increasing load parallel with the declination axis.

In the azimuthal mounting, one of the axes is vertical, the other one horizontal. This mounting is easier to construct than the equatorial mounting and is more stable for very large telescopes. In order to follow the rotation of the sky, the telescope must be turned around both of the axes with changing velocities. The field of view will also rotate; this rotation must be compensated for when the telescope is used for photography.

If an object goes close to the zenith, its azimuth will change 180° in a very short time. Therefore, around the zenith there is a small region where observations with an azimuthal telescope are not possible.

The largest telescopes in the world were equatorially mounted until the development of computers made possible the more complicated guidance needed for azimuthal mountings. Most of the recently built large telescopes are already azimuthally mounted. Azimuthally mounted telescopes have two additional obvious places for foci, the *Nasmyth foci* at both ends of the horizontal axis.

The *Dobson mounting*, used in many amateur telescopes, is azimuthal. The magnification of the Newtonian telescope is usually small, and the telescope rests on pieces of teflon, which make it very easy to move. Thus the object can easily be tracked manually.

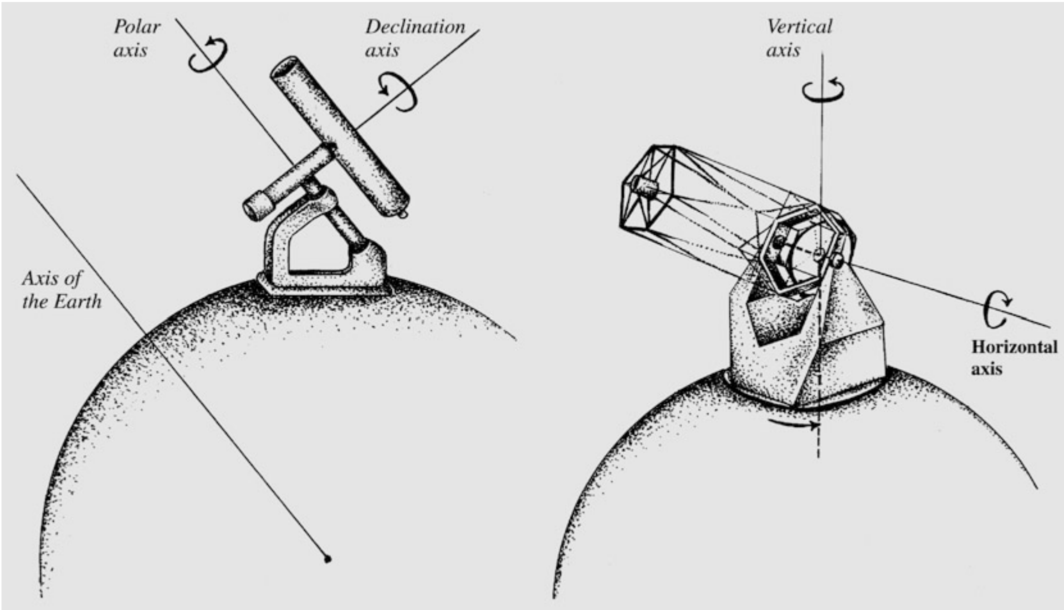


Fig. 3.17 The equatorial mounting (*left*) and the azimuthal mounting (*right*)

Another type of mounting is the *coelostat*, where rotating mirrors guide the light into a stationary telescope. This system is used especially in solar telescopes.

To measure absolute positions of stars and accurate time, telescopes aligned with the north–south direction are used. They can be rotated around one axis only, the east–west horizontal axis. *Meridian circles* or *transit instruments* with this kind of mounting were widely constructed for different observatories during the 19th century. A few are still used for astrometry, but they are now highly automatic. The most important of the still operational instruments is that of the U.S. Naval Observatory in Flagstaff. The meridian circle on La Palma funded by the Carlsberg foundation was closed in 2013.

New Techniques Detectors are already approaching the theoretical limit of efficiency, where all incident photons are registered. Ultimately, to detect even fainter objects the only solution is to increase the light gathering area (Fig. 3.18), but also the mirrors are getting close to the practical maximum size. Thus, new technical solutions are needed.

One new feature is *active optics*, used in most new big telescopes. The mirror is very thin, but its shape is kept exactly correct by a computer controlled support mechanism. The weight and production cost of such a mirror are much smaller compared with a conventional thick mirror. Because of the smaller weight also the supporting structure can be made lighter.

Developing the control mechanism further leads to *adaptive optics*. A reference star (or an artificial beam) is monitored constantly in order to obtain the shape of the seeing disk. The attitude or the shape of a small auxiliary mirror is adjusted up to hundreds of times a second to keep the image as concentrated as possible. Adaptive optics has been taken into use in the largest telescopes of the world from about the year 2000 on.

The mirrors of large telescopes need not be monolithic, but can be made of smaller pieces that are, e.g. hexagonal. These *mosaic mirrors* are very light and can be used to build up mirrors with diameters of several tens of metres (Fig. 3.20). Using active optics, the hexagons can be accurately focused. The California Association for Research in Astronomy has constructed the William M. Keck telescope with a 10 m mo-

Fig. 3.18 The largest telescopes in the world in 1947–2000. (**a**) For nearly 30 years, the 5.1 m Hale telescope on Mount Palomar, California, USA, was the largest telescope in the world. (**b**) The BTA, Big Azimuthal Telescope, is situated in the Caucasus in the southern Soviet Union. Its mirror has a diameter of 6 m. It was set in operation at the end of 1975. (**c**) The William M. Keck Telescope on the summit of Mauna Kea, Hawaii, was completed in 1992. The 10 m mirror consists of 36 hexagonal segments. Another similar telescope next to it was completed in 1996. (Photos Palomar Observatory, Spetsialnaya Astrofizitsheskaya Observatorya, and Roger Ressmeyer–Starlight for the California Association for Research in Astronomy)

saic mirror. It is located on Mauna Kea, and the last segment was installed in 1992. Also the *E-ELT* telescope (Fig. 3.21) planned by ESO will have a mosaic mirror.

The reflecting surface does not have to be continuous, but can even consist of several separate telescopes. Next to the Keck another similar telescope, the Keck II, was completed in 1996. The pair can be used as an interferometer that has the same resolving power as a single 75 metre mirror.

The European Southern Observatory has constructed its own multi-mirror telescope. ESO's Very Large Telescope (VLT) has four closely located telescopes (Fig. 3.19). The diameter of each mirror is eight metres, and the total area corresponds to one telescope with a 16 m mirror. The resolution is even better, since the "aperture", i.e. the maximum distance between the mirrors, is several tens of meters.

An important astronomical instruments of the 20th century is the *Hubble Space Telescope*, launched in 1990 (Fig. 3.22). It has a mirror with a diameter of 2.4 m. The resolution of the telescope (after the faulty optics was corrected) is near the theoretical diffraction limit, since there is no disturbing atmosphere. A second generation Space Telescope, now called the *James Webb Space Telescope*, will have a mosaic mirror of 6.5 m. The launch has been postponed mainly for financial reasons. The current estimates for the launch are around 2018.

Fig. 3.18 (Continued)

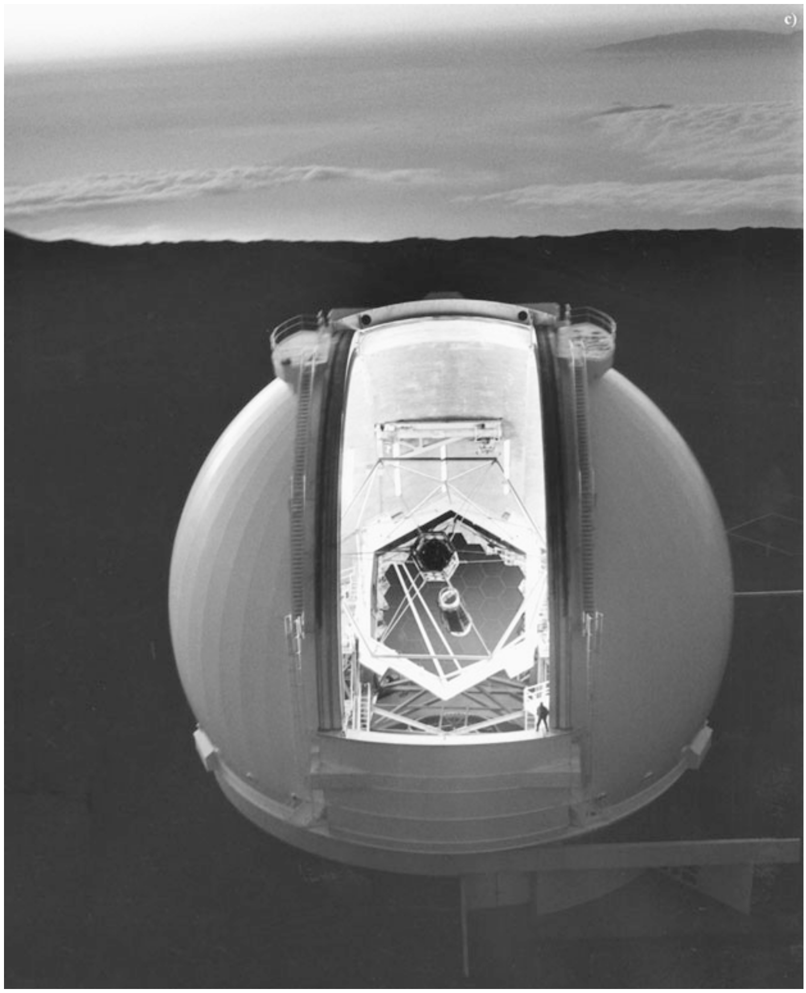
Space telescopes avoid the atmospheric perturbations. However, comparable results are nowadays obtained using the adaptive optics. Due to budgetary reasons, the majority of astronomical observations will still be carried out on the Earth, and great attention will be given to improving ground-based observatories and detectors.

In the future, satellites will continue to be mainly used for those wavelength regions where the radiation is absorbed by the atmosphere. (Fig. 3.2 and Sect. 3.5).

3.3 Detectors and Instruments

Only a limited amount of information can be obtained by looking through a telescope with the unaided eye. Yet until the end of the 19th century this was the only way to make observations. The invention of photography in the middle of the 19th century brought a revolution in astronomy. The next important step forward in optical astronomy was the development of photoelectric photometry in the 1940's and 1950's. A new revolution, comparable to that caused by the invention of photography, took place in the middle of the 1970's with the introduction of different semiconductor detectors. The sensitivity of detectors has grown so much that today, a 60 cm telescope can be used for observations similar to those made with the Palomar 5 m telescope when it was set in operation in the 1940's.

Fig. 3.18 (Continued)



The Photographic Plate *Photography* has long been one of the most common methods of observation in astronomy. In astronomical photography glass plates were used, rather than film, since they keep their shape better, but nowadays they are no more manufactured, and CCD-cameras have largely replaced photography. The sensitive layer on the surface of the film or plate is made up of a silver halide, usually silver bromide, AgBr. A photon absorbed by the halide excites an electron that can move from one atom to another. A silver ion, Ag^+, can catch the electron, becoming a neutral atom. When the necessary amount of silver atoms have been accumulated at one place, they form a latent image. The latent image can be made into a permanent negative by treating the plate after exposure with various chemicals, which transform the silver bromide crystals enclosing the latent image into silver ("development"), and remove the unexposed crystals ("fixing").

The photographic plate has many advantages over the human eye. The plate can register up to millions of stars (picture elements) at one time, while the eye can observe at most one or two objects at a time. The image on a plate is practically permanent—the picture can be studied at any time. In addition, the photographic plate is cheap and easy to use, as compared to many other detectors. The most important feature of a plate is its capability to collect light over an extended time: the longer exposures are used, the more sil-

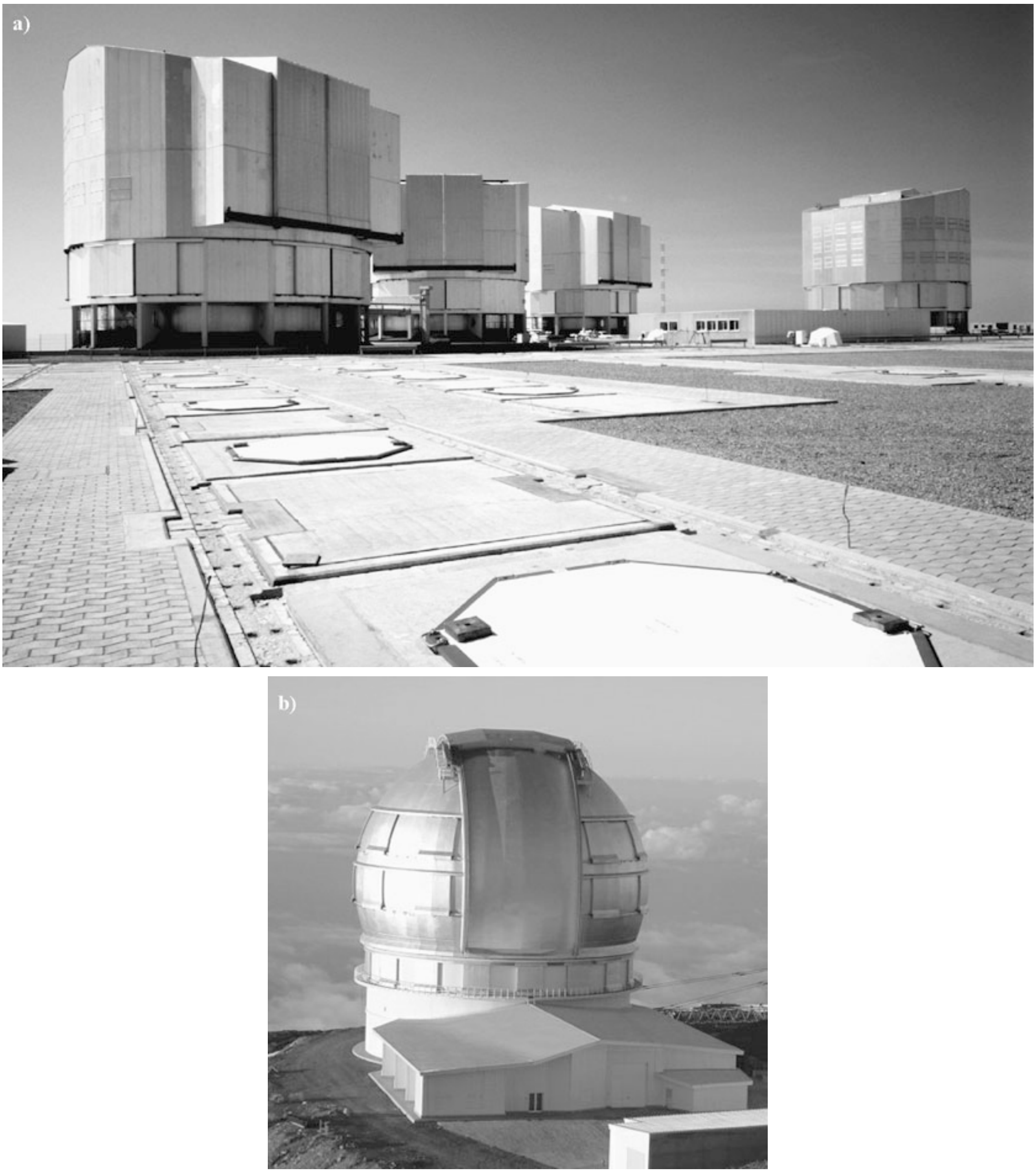


Fig. 3.19 Some new large telescopes. (**a**) The European Southern Observatory (ESO) was founded by Belgium, France, the Netherlands, Sweden and West Germany in 1962. Other European countries have joined them later. The VLT (Very Large Telescope) on Cerro Paranal in Northern Chile, was inaugurated in 1998–2000. (**b**) Currently the largest telescope is the GTC (Gran Telescopio Canarias or GranTeCan) on La Palma. The observations with the 10.4 m mirror were made in 2009. (**c**) The largest binocular is the LBT (Large Binocular Telescope) on Mount Graham in Arizona, completed in 2008. It has two 8.4 metre mirrors. (Photos ESO, IAC, Max Planck Institute)

ver atoms are formed on the plate (the plate darkens). By increasing the exposure times, fainter objects can be photographed. The eye has no such capacity: if a faint object does not show through a telescope, it cannot been seen, no matter how long one stares.

Fig. 3.19 (Continued)

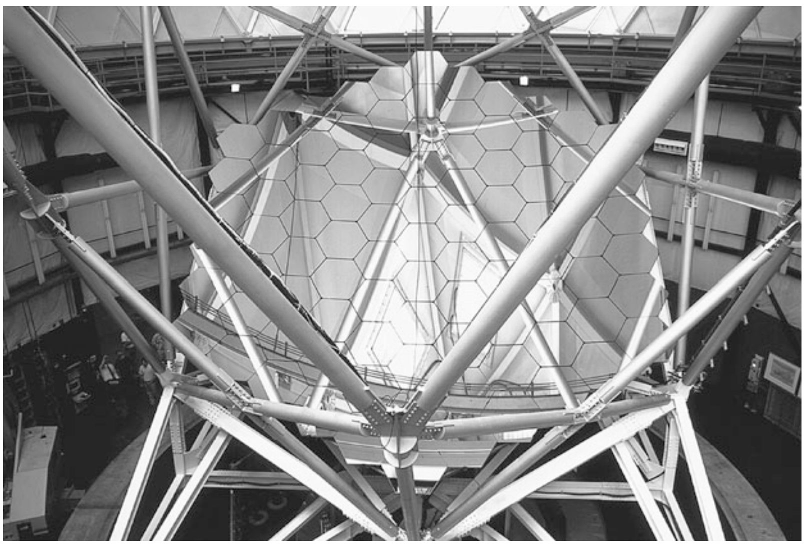


Fig. 3.20 The mirror of a telescope can be made up of several smaller segments, which are much easier to manufacture, as in the Hobby–Eberle Telescope on Mount Fowlkes, Texas. The effective diameter of the mirror is 9.1 m. A similar telescope is being built in South Africa. (Photo MacDonald Observatory)

One disadvantage of the photographic plate is its low sensitivity. Only one photon in a thousand causes a reaction leading to the formation of a silver grain. Thus the *quantum efficiency* of the plate is only 0.1 %. Several chemical treatments can be used to sensitise the plate before exposure. This brings the quantum efficiency up to a few percent. Another disadvantage is the fact that a silver bromide crystal that has been exposed once does not register anything more, i.e. a saturation point is reached. On the other hand, a certain number of photons are needed to produce an image. Doubling the number of photons does not necessarily double the density (the 'blackness' of the image): the density of the plate depends nonlinearly on the amount of incoming light (Fig. 3.23). The sensitivity of the plate is also strongly dependent on the wavelength of the light. For the reasons mentioned above the accuracy with which brightness can be measured

Fig. 3.21 In order to observe objects essentially fainter than nowadays the size of the objective must be increased considerably. The European Southern Observatory ESO started to make plans for a 100 metre OWL telescope (Overwhelmingly Large telescope), but when the price turned out to be "astronomical" the size has been reduced several times. When the E-ELT telescope (European Extremely Large Telescope) will be completed the main mirror will have a diameter of 39 metres. The mirror will consist of 798 hexagonal pieces, each having a diameter of 1.45 metres. (ESO)

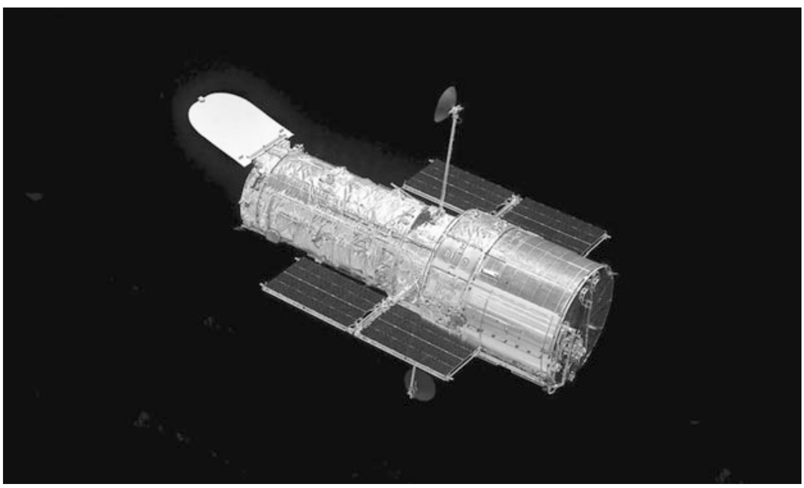


Fig. 3.22 The Hubble Space Telescope after the latest service flight in 2009. The telescope got new solar panels and several other upgrades. (Photo NASA)

on a photographic plate is usually worse than about 5 %. Thus the photographic plate makes a poor photometer, but it can be excellently used, e.g. for measuring the positions of stars (positional astronomy) and for mapping the sky. Old photography archives are still valuable for studying proper motions and brightness variations that are not too minuscule.

Photocathodes, Photomultipliers A *photocathode* is a more effective detector than the photographic plate. It is based on the photoelectric effect. A light quantum, or photon, hits the photocathode and loosens an electron. The electron moves to the positive electrode, or anode, and gives rise to an electric current that can be measured. The quantum efficiency of a photocathode

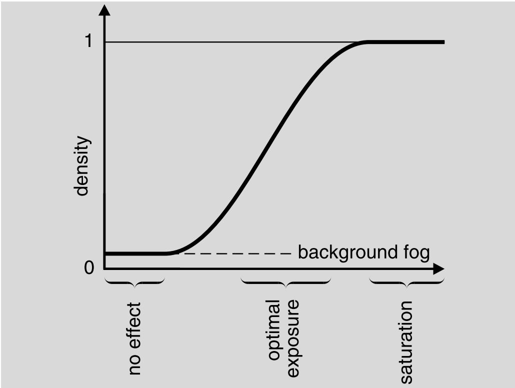


Fig. 3.23 The density of a photographic emulsion depends on the incident light. If there is very little light it has no effect. Even then the film is not completely transparent since it has a light background fog. When the amount of light increases the density will begin to grow first nonlinearly. When the exposure is optimal, which depends of the type of the film, the dependence of the density on exposure is nearly linear. When there is too much light, all grains of the film become dark and the image is saturated

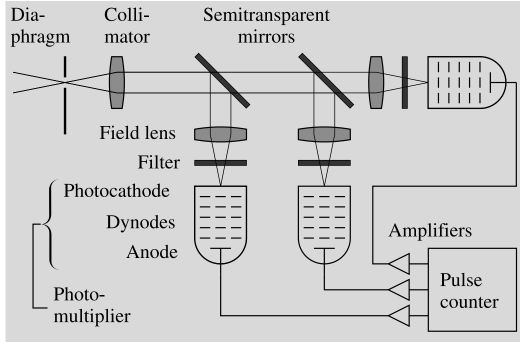


Fig. 3.24 The principle of a photoelectric multicolour photometer. Light collected by the telescope arrives from the left. The light enters the photometer through a small hole in the focal plane, the diaphragm. A lens collimates the light into a parallel beam. Semitransparent mirrors divide the beam to several photomultipliers. A field lens guides the light through a filter onto the photocathode of the photomultiplier. The quanta of light, photons, release electrons from the cathodes. The electrons are accelerated towards the dynodes with a voltage of about 1500 V. The electrons hitting the dynodes release still more electrons, and the current is greatly enhanced. Every electron emitted from the cathode gives rise to a pulse of up to 10^8 electrons at the anode; the pulse is amplified and registered by a pulse counter. In this way, the photons from the star are counted

is about 10–20 times better than that of a photographic plate; optimally, an efficiency of 30 % can be reached. A photocathode is also a linear detector: if the number of electrons is doubled, the outcoming current is also doubled.

The *photomultiplier* is one of the most important applications of the photocathode. In this device, the electrons leaving the photocathode hit a dynode. For each electron hitting the dynode, several others are released. When there are several dynodes in a row, the original weak current can be intensified a millionfold. The photomultiplier measures all the light entering it, but does not form an image. Photomultipliers are mostly used in photometry, and an accuracy of 0.1–1 % can be attained.

Photometers, Polarimeters A detector measuring brightness, *a photometer*, is usually located behind the telescope in the Cassegrain focus. In the focal plane there is a small hole, the *diaphragm*, which lets through light from the object under observation. In this way, light from other stars in the field of view can be prevented from entering the photometer. A *field lens* behind the diaphragm refracts the light rays onto a photocathode. The outcoming current is intensified further in a preamplifier. The photomultiplier needs a voltage of 1000–1500 volts.

Observations are often made in a certain wavelength interval, instead of measuring all the radiation entering the detector. In this case a *filter* is used to prevent other wavelengths from reaching the photomultiplier. A photometer can also consist of several photomultipliers (Fig. 3.23), which measure simultaneously different wavelength bands. In such an instrument beam splitters or semitransparent mirrors split the light beam through fixed filters to the photomultipliers.

In a device called the *photopolarimeter*, a *polarising filter* is used, either alone or in combination with other filters. The degree and direction of polarisation can be found by measuring the intensity of the radiation with different orientations of the polarisers.

In practice, the diaphragm of a photometer will always also let through part of the background sky around the observed object. The measured brightness is in reality the combined brightness of the object and the sky. In order to find the brightness of the object, the background brightness must be measured separately and subtracted from the combined brightness. The accuracy of the measurements is decreased if long observation times are used and the background brightness undergoes fast changes. The problem can be solved by observing the brightness of the background sky and the object simultaneously.

Photometric observations are often relative. If one is observing, e.g. a variable star, a *reference star* close to the actual target is observed at regular intervals. Using the observations of this reference star it is possible to derive a model for the slow changes in the atmospheric extinction (see Chap. 4) and remove their effect. The instrument can be calibrated by observing some *standard stars*, whose brightness is known very accurately.

Image Intensifiers Different *image intensifiers* based on the photocathode have been used since the 1960's. In the intensifier the information about the starting point of the electron on the photocathode is preserved and the intensified image is formed on a fluorescent screen. The image can then be registered, e.g. with a CCD camera. One of the advantages of the image intensifier is that even faint objects can be imaged using relatively short exposures, and observations can be made at wavelengths where the detector is insensitive.

Another common type of detector is based on the TV camera (*Vidicon camera*). The electrons released from the photocathode are accelerated with a voltage of a few kilovolts before they hit the electrode where they form an image in the form of an electric charge distribution. After exposure, the charge at different points of the electrode is read by scanning its surface with an electron beam row by row. This produces a video signal, which can be transformed into a visible image on a TV tube. The information can also be saved in digital form. In the most advanced systems, the scintillations caused by single electrons on the fluorescent screen of the image intensifier can be registered and stored in the memory of a computer. For each point in the image there is a memory location, called a picture element or *pixel*.

Since the middle of the 1970's, detectors using semiconductor techniques began to be used in increasing numbers. With semiconductor detectors a quantum efficiency of about 70–80 % can be attained; thus, sensitivity cannot be improved much more. The wavelength regions suitable for these new detectors are much wider than in the case of the photographic plate. The detectors are also linear: if the number of photons is doubled, also the signal will be doubled. Computers are used for collecting, saving and analysing the output data available in digital form.

CCD Camera Currently the most important detector is the *CCD camera* (Charge Coupled Device). The detector consists of a surface made up of light sensitive silicon diodes, arranged in a rectangular array of image elements or pixels. The largest cameras can have as many as 4096×4096 pixels, although most are considerably smaller. Recently, CCD's have become widely used in ordinary digital cameras.

A photon hitting the detector can release an electron, which will remain trapped inside a pixel. After the exposure varying potential differences are used to move the accumulated charges row by row to a readout buffer. In the buffer the charges are moved pixel by pixel to an analogy/digital converter, which transmits the digital value to a computer. Reading an image also clears the detector (Fig. 3.24). If the exposures are very short the readout times may take a substantial part of the observing time.

The CCD camera is nearly linear: the number of electrons is directly proportional to the number of photons. Calibration of the data is therefore much easier than with photographic plates (Fig. 3.25).

The quantum efficiency, i.e. the number of electrons per incident photon, is high, and the CCD camera is much more sensitive than a photographic plate. The sensitivity is highest in the red wavelength range, about 600–800 nm, where the quantum efficiency can be 80–90 % or even higher.

The range of the camera extends far to the infrared. In the ultraviolet the sensitivity drops due to the absorption of the silicon very rapidly below about 500 nm. Two methods have been used

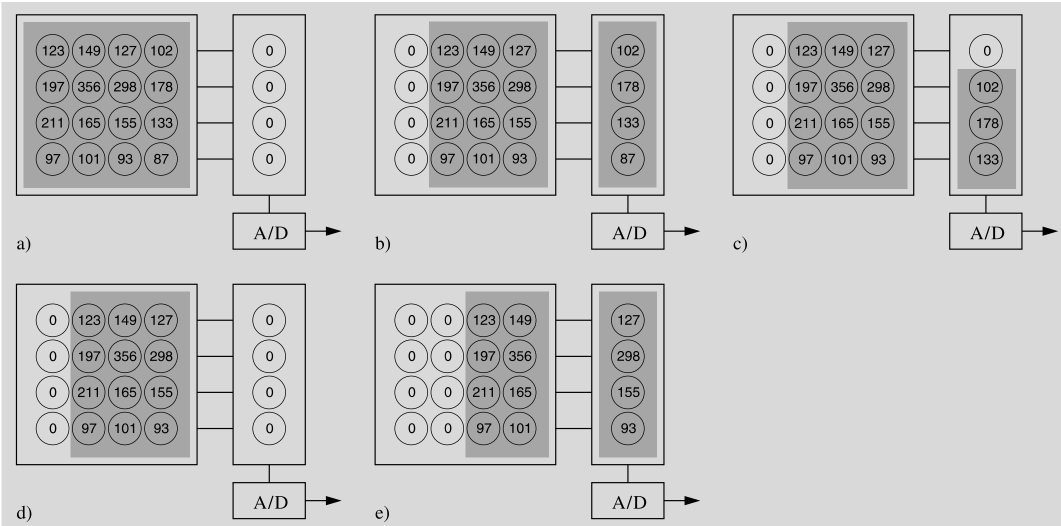


Fig. 3.25 The principle of reading a CCD camera. (**a**) During an exposure electrons are trapped in potential wells corresponding to pixels of the camera. The number at each pixel shows the number of electrons. (**b**) After the exposure each horizontal line is moved one pixel to the right; the rightmost row moves to the readout buffer. (**c**) The contents of the buffer is moved down by one pixel. The lowermost charge moves to the A/D converter, which sends the number of electrons to the computer. (**d**) After moving the buffer down several times one vertical row has been read. (**e**) The image is again shifted right by one pixel. This procedure is repeated till the whole image is read

to avoid this problem. One is to use a coating that absorbs the ultraviolet photons and emits light of longer wavelength. Another possibility is to turn the chip upside down and make it very thin to reduce the absorption.

The thermal noise of the camera generates *dark current* even if the camera is in total darkness. To reduce the noise the camera must be cooled. Astronomical CCD cameras are usually cooled with liquid nitrogen, which efficiently removes most of the dark current. However, the sensitivity is also reduced when the camera is cooled; so too cold is not good either. The temperature must be kept constant in order to obtain consistent data. For amateurs there are already moderately priced CCD cameras, which are electrically cooled. Many of them are good enough also for scientific work, if very high sensitivity is not required.

The dark current can easily be measured by taking exposures with the shutter closed. Subtracting this from the observed image gives the real number of electrons due to incident light (Fig. 3.26).

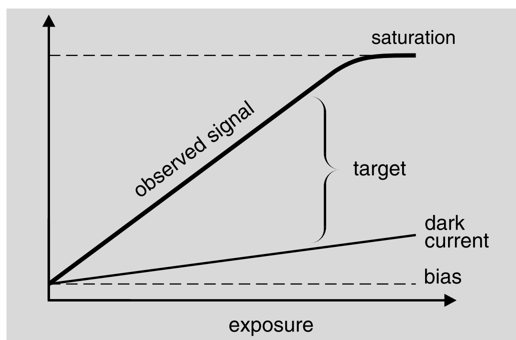


Fig. 3.26 The signal measured by a CCD-camera depends almost linearly on the number of photons. During the exposure also the dark current will increase and has to be subtracted from the observed values. Bias is the dark current registered even if the exposure time were zero

The sensitivity of individual pixels may be slightly different. This can be corrected for by taking an image of an evenly illuminated field, like a twilight sky, where stars are not visible. This image is called a *flat-field*. When observations are divided by the flat-field, the error caused by different pixels is removed (Fig. 3.27).

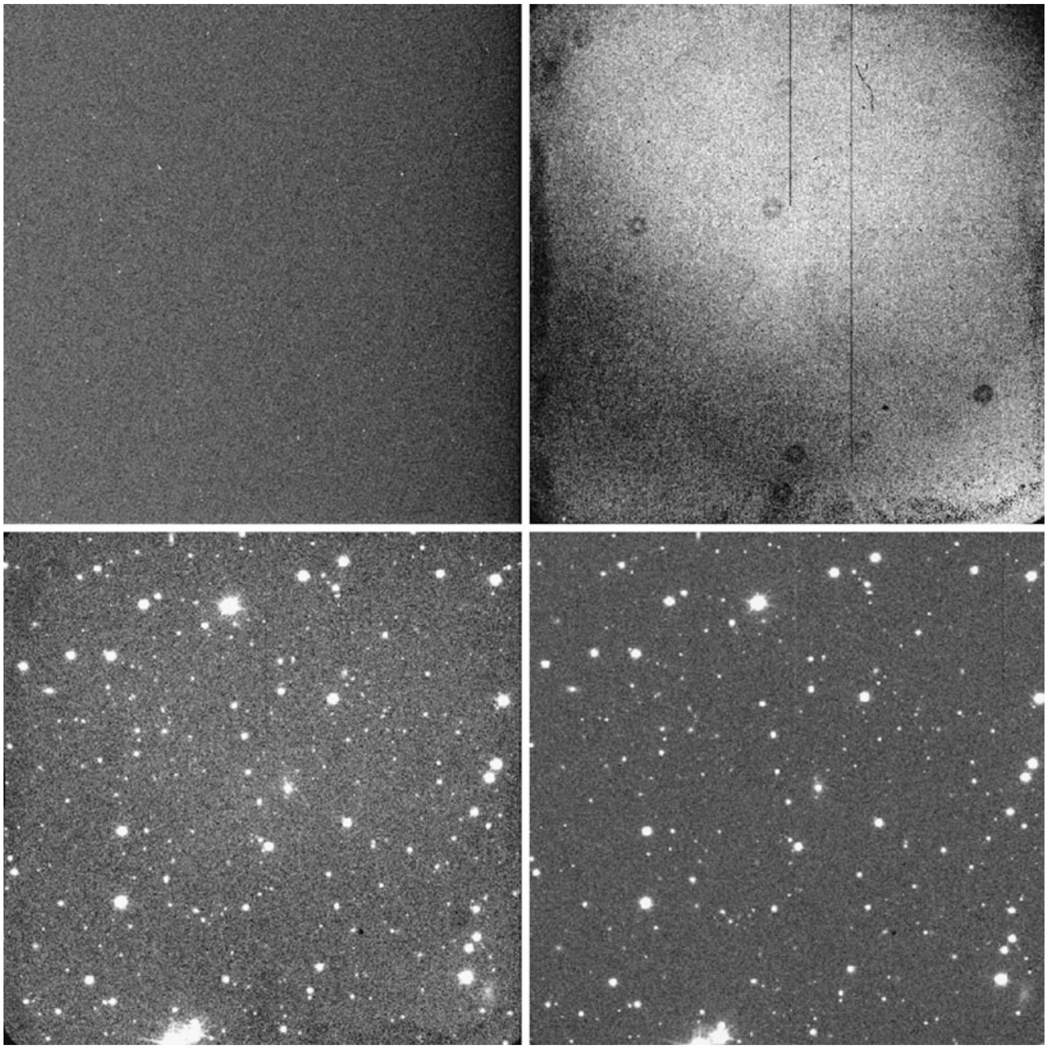


Fig. 3.27 Typical images related to CCD observations. *Top left* dark current, *top right* flat field, *bottom left* original raw image and *bottom right* reduced image

The CCD camera is very stable. Therefore it is not necessary to repeat the dark current and flat-field observations very frequently. Typically these calibration exposures are taken during evening and morning twilights, just before and after actual observations.

Cosmic rays are charged particles that can produce extraneous bright dots in CCD images. They are usually limited to one or two pixels, and are easily identified. Typically a short exposure of a few minutes contains a few traces of cosmic rays. Instead of a single long exposure it is usually better to take several short ones, clean the images from cosmic rays, and finally add the images on a computer.

A more serious problem is the *readout noise* of the electronics. In the first cameras it could be hundreds of electrons per pixel. In modern cameras it is a few electrons. This gives a limit to the faintest detectable signal: if the signal is weaker than the readout noise, it is indistinguishable from the noise. Nowadays there are so called

L3-cameras (low light level CCD), where electrons are multiplied during reading and the readout noise is the only source of noise.

Although the CCD camera is a very sensitive detector, even bright light cannot damage it. A photomultiplier, on the other hand, can be easily destroyed by letting in too much light. However, one pixel can only store a certain number of electrons, after which it becomes *saturated*. Excessive saturation can make the charge to overflow also to the neighbouring pixels. If the camera becomes badly saturated it may have to be read several times to completely remove the charges.

The largest CCD cameras are quite expensive, and even they are still rather small compared with the largest photographic plates and films. Because they are very sensitive and processing of the data is easy, they have, however, replaced photographic materials almost completely.

Spectrographs The simplest spectrograph is a prism that is placed in front of a telescope. This kind of device is called the *objective prism spectrograph*. The prism spreads out the different wavelengths of light into a spectrum which can be registered. During the exposure, the telescope is usually slightly moved perpendicularly to the spectrum, in order to increase the width of the spectrum. With an objective prism spectrograph, large numbers of spectra can be photographed, e.g. for spectral classification (Chap. 9).

For more accurate information the *slit spectrograph* must be used (Fig. 3.28). It has a narrow slit in the focal plane of the telescope. The light is guided through the slit to a collimator that reflects or refracts all the light rays into a parallel beam. After this, the light is dispersed into a spectrum by a prism and focused with a camera onto a detector, which nowadays is usually a CCD camera. A comparison spectrum is exposed next to the stellar spectrum to determine the precise wavelengths. In modern spectrographs using CCD cameras, the comparison spectrum is usually exposed as a separate image. A big slit spectrograph is often placed at the coudé or Nasmyth focus of the telescope.

The most important property of a spectrograph is the scale or *dispersion* of the spectrum. The dispersion tells how long a wavelength range corresponds to one distance unit of the detector. For the objective prism the dispersion is typically a few tens of nanometres per millimetre. Slit spectrographs may have a dispersion as high as 1–0.01 nm/mm, which makes it possible to study even the shapes of individual spectral lines. Often the dispersion is given as a dimensionless quantity. For instance, the dispersion 1 nm/mm means that the wavelength scale has increased millionfold; thus it can be expressed as 10^6.

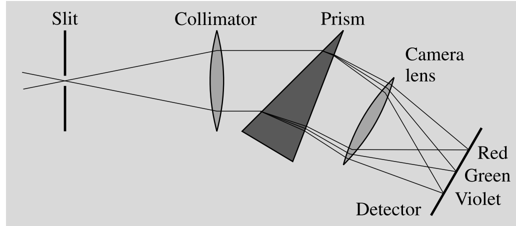


Fig. 3.28 The principle of the slit spectrograph. Light rays entering through a slit are collimated (made parallel to each other), dispersed into a spectrum by a prism and projected onto a photographic plate or a CCD. There can also be several prisms. However, grids are used more commonly since they give a better resolution

Instead of the prism a *diffraction grating* can be used to form the spectrum. A grating has narrow grooves, side by side, typically several hundred per millimetre. When light is reflected by the walls of the grooves, the adjoining rays interfere with each other and give rise to spectra of different orders. There are two kinds of gratings: *reflection* and *transmission gratings*. In a reflection grating no light is absorbed by the glass as in the prism or transmission grating. A grating usually has higher dispersion, or ability to spread the spectrum, than a prism. The dispersion can be increased by increasing the density of the grooves of the grating. In slit spectrographs the reflection grating is most commonly used. In some instruments a *grism* is used; it is a prism with a transmission grating on one of its surfaces.

Interferometers The resolution of a big telescope is in practice limited by seeing, and thus increasing the aperture does not necessarily improve the resolution. To get nearer to the theoretical resolution limit set by diffraction (Eq. (3.3), Fig. 3.6), different *interferometers* can be used.

There are two types of optical interferometers. One kind uses an existing large telescope; the other a system of two or more separate telescopes. In both cases the light rays are allowed to interfere. By analysing the outcoming interference pattern, the structures of close binaries can be studied, apparent angular diameters of the stars can be measured, etc.

One of the earliest interferometers was the *Michelson interferometer* that was built shortly before 1920 for the largest telescope of that time. In front of the telescope, at the ends of a six metre long beam, there were flat mirrors reflecting the light into the telescope. The form of the interference pattern changed when the separation of the mirrors was changed. In practice, the interference pattern was disturbed by seeing, and only a few positive results were obtained with this instrument.

The diameters of over 30 of the brightest stars have been measured using *intensity interferometers*. Such a device consists of two separate telescopes that can be moved in relation to each other. This method is suitable for the brightest objects only.

In 1970 the Frenchman *Antoine Labeyrie* introduced the principle of *speckle interferometry*. In traditional imaging the pictures from long exposures consist of a large number of instantaneous images, "speckles", that together form the seeing disk. In speckle interferometry very short exposures and large magnifications are used and hundreds of pictures are taken. When these pictures are combined and analysed (usually in digital form), the actual resolution of the telescope can nearly be reached.

The accuracy of interferometric techniques was improved at the beginning of 00's. The first experiments to use the two 10 m Keck telescopes as one interferometer, were made in 2001. Similarly, the ESO VLT can be used as an interferometer.

3.4 Radio Telescopes

Radio astronomy, started in the 1930's, extended the observable range of the electromagnetic spectrum by several orders of magnitude. Radio frequencies range from a few megahertz (100 m) up to about 300 GHz (1 mm). The low-frequency limit of the radio band is determined by the opacity of the ionosphere, while the high-frequency limit is due to the strong absorption from oxygen and water bands in the lower atmosphere. Neither of these limits is very strict, and under favourable conditions radio astronomers can work into the submillimetre region or through ionospheric holes during sunspot minima.

At the beginning of the 20th century attempts were made to observe radio emission from the Sun. These experiments, however, failed because of the low sensitivity of the antenna–receiver systems, and because of the opaqueness of the ionosphere at the low frequencies at which most of the experiments were carried out. The first observations of cosmic radio emission were later made by the American engineer Karl G. Jansky in 1932, while studying thunderstorm radio disturbances at a frequency of 20.5 MHz (14.6 m). He discovered radio emission of unknown origin, which varied within a 24 hour period. Somewhat later he identified the source of this radiation to be in the direction of the centre of our Galaxy.

The real birth of radio astronomy may perhaps be dated to the late 1930's, when Grote Reber started systematic observations with his homemade 9.5 m paraboloid antenna. Thereafter radio astronomy developed quite rapidly and has greatly improved our knowledge of the Universe.

Observations are made both in the continuum (broad band) and in spectral lines (radio spectroscopy). Much of our knowledge about the structure of our Milky Way comes from radio observations of the 21 cm line of neutral hydrogen and, more recently, from the 2.6 mm line of the carbon monoxide molecule. Radio astronomy has resulted in many important discoveries; e.g. both pulsars and quasars were first found by radio astronomical observations. The importance of the field can also be seen from the fact that the Nobel prize in physics has recently been awarded three times to radio astronomers.

A radio telescope collects radiation in an aperture or antenna, from which it is transformed to an electric signal by a receiver, called a radiometer. This signal is then amplified, detected and integrated, and the output is registered on some recording device, nowadays usually by a computer. Because the received signal is very weak, one has to use sensitive receivers. These are often cooled to minimise the noise, which could otherwise mask the signal from the source. Because radio waves are electromagnetic radiation, they are reflected and refracted like ordinary light waves. In radio astronomy, however, only reflecting telescopes are used.

At low frequencies the antennas are usually dipoles (similar to those used for radio or TV), but in order to increase the collecting area and improve the resolution, one uses dipole arrays, where all dipole elements are connected to each other.

The most common antenna type, however, is a parabolic reflector, which works exactly as an optical mirror telescope. At long wavelengths the reflecting surface does not need to be solid, because the long wavelength photons cannot see the holes in the reflector, and the antenna is therefore usually made in the form of a metal mesh. At high frequencies the surface has to be smooth, and in the millimetre-submillimetre range, radio astronomers even use large optical telescopes, which they equip with their own radiometers. To ensure a coherent amplification of the signal, the surface irregularities should be less than one-tenth of the wavelength used.

The main difference between a radio telescope and an optical telescope is in the recording of the signal. Radio telescopes are not imaging telescopes (except for synthesis telescopes, which will be described later); instead, a feed horn, which is located at the antenna focus, transfers the signal to a receiver. The wavelength and phase information is, however, preserved.

The resolving power of a radio telescope, θ, can be deduced from the same formula (3.4) as for optical telescopes, i.e. λ/D, where λ is the wavelength used and D is the diameter of the aperture. Since the wavelength ratio between radio and visible light is of the order of 10,000, radio antennas with diameters of several kilometres are needed in order to achieve the same resolution as for optical telescopes. In the early days of radio astronomy poor resolution was the biggest drawback for the development and recognition of radio astronomy. For example, the antenna used by Jansky had a fan beam with a resolution of about 30° in the narrower direction. Therefore radio observations could not be compared with optical observations. Neither was it possible to identify the radio sources with optical counterparts.

The world's biggest radio telescope is the Arecibo antenna in Puerto Rico, whose main reflector has a diameter of 305 m. It is a fixed metal mesh built over a natural round valley (Fig. 3.29). In the late 1970's the antenna surface and receivers were upgraded, enabling the antenna to be used down to wavelengths of 5 cm. The mirror of the Arecibo telescope is not parabolic but spherical, and the antenna is equipped with a movable feed system, which makes observations possible within a 20° radius around the zenith.

A similar and only partly steerable radio telescope is being built in China. It will be the largest in the world with an antenna diameter of 500 m.

The biggest completely steerable radio telescope is the Green Bank telescope in Virginia, U.S.A., dedicated at the end of 2000. It is slightly asymmetric with a diameter of 100×110 m (Fig. 3.30). Before the Green Bank telescope, for over two decades the largest telescope was the Effelsberg telescope in Germany. This antenna has a parabolic main reflector with a diameter of 100 m. The inner 80 m of the dish is made of solid aluminium panels, while the outmost portion of the disk is a metal mesh structure. By using only the inner portion of the telescope, it has been possible to observe down to wavelengths of 4 mm. The oldest and perhaps best-known big radio telescope is the 76 m antenna at Jodrell Bank in Britain, which was completed in the end of the 1950's.

The biggest telescopes are usually incapable of operating below wavelengths of 1 cm, because the surface cannot be made accurate enough. However, the millimetre range has become more and more important. In this wavelength range there are many transitions of interstellar molecules, and one can achieve quite high angular resolution even with a single dish telescope. At

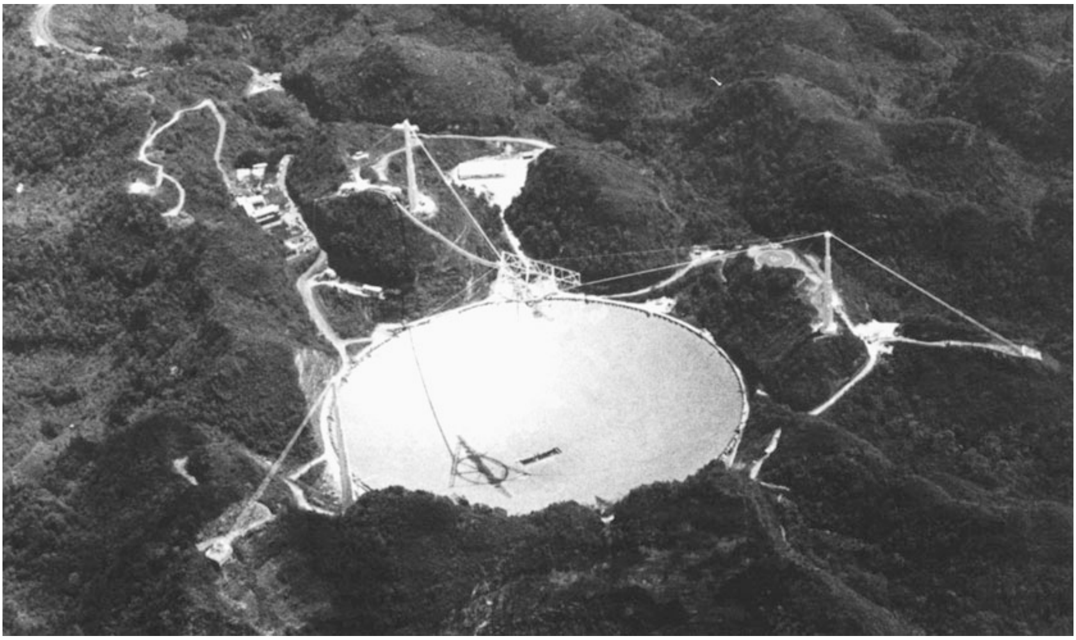


Fig. 3.29 The largest radio telescope in the world is the Arecibo dish in Puerto Rico. It has been constructed over a natural bowl and is 300 m in diameter. (Photo Arecibo Observatory)

Fig. 3.30 The largest fully steerable radio telescope is in Green Bank, Virginia. Its diameter is 100×110 m. (Photo NRAO)

present, the typical size of a mirror of a millimetre telescope is about 15 m. The development of this field is rapid, and at present several big millimetre telescopes are in operation (Table C.24). Among them are the 40 m Nobeyama telescope in Japan, which can be used down to

Fig. 3.31 The Atacama Large Millimetre Array (ALMA) built in cooperation by Europe, U.S.A. and Japan. The telescope has 64 antennas. (Photo ESO/NOAJ)

3 mm, the 30 m IRAM telescope at Pico Veleta in Spain, which is usable down to 1 mm, and the 15 m. The largest project in the first decade of the 21st century is ALMA (Atacama Large Millimetre Array), which consists of 64 telescopes with a diameter of 12 m (Fig. 3.31). It was built as an international project by the United States, Europe and Japan in Chile on the Chajnantor plateau.

As already mentioned, the resolving power of a radio telescope is far poorer than that of an optical telescope. The biggest radio telescopes can at present reach a resolution of 5 arc seconds, and that only at the very highest frequencies. To improve the resolution by increasing the size is difficult, because the present telescopes are already close to the practical upper limit. However, by combining radio telescopes and interferometers, it is possible to achieve even better resolution than with optical telescopes.

As early as 1891 Michelson used an interferometer for astronomical purposes. While the use of interferometers has proved to be quite difficult in the optical wavelength regime, interferometers

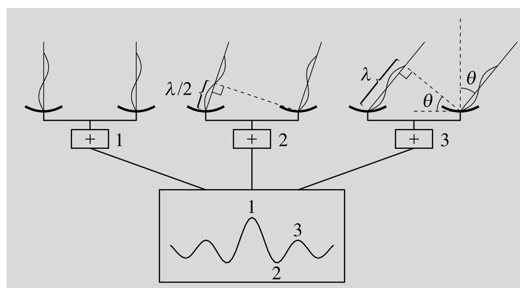


Fig. 3.32 The principle of an interferometer. If the radiation reaches the radio telescopes in the same phase, the waves amplify each other and a maximum is obtained in the combined radiation (cases *1* and *3*). If the incoming waves are in opposite phase, they cancel each other (case *2*)

are extremely useful in the radio region. To form an interferometer, one needs at least two antennas coupled together. The spacing between the antennas, D, is called the baseline. Let us first assume that the baseline is perpendicular to the line of sight (Fig. 3.32). Then the radiation arrives at both antennas with the same phase, and the summed signal shows a maximum. However, due to the rotation of the Earth, the direction of the baseline changes, producing a phase difference between the two signals. The result is a sinusoidal interference pattern, in which minima occur when the phase difference is 180 degrees. The distance between the peaks is given by

$$\theta D = \lambda,$$

where θ is the angle the baseline has turned and λ is the wavelength of the received signal. The resolution of the interferometer is thus equal to that of an antenna with a linear size equal to D.

If the source is not a point source, the radiation emitted from different parts of the source will have phase differences when it enters the antennas. In this case the minima of the interference pattern will not be zero, but will have some positive value $P_{\min}$. If we denote the maximum value of the interference pattern by $P_{\max}$, the ratio

$$\frac{P_{\max} - P_{\min}}{P_{\max} + P_{\min}}$$

gives a measure of the source size (fringe visibility).

More detailed information about the source structure can be obtained by changing the baseline by changing the relative positions of the antennas. In principle it is possible to observe the same amount of details as with a single antenna with the same size as the area covered by the movable antennas in their different positions. This way interferometry is transformed into a technique called *aperture synthesis*.

The theory and techniques of aperture synthesis were developed particularly by the British astronomer Sir Martin Ryle. In Fig. 3.33 the principle of aperture synthesis is illustrated. If the telescopes are located on an east–west track, the spacing between them, projected onto the sky, will in 12 hours describe a circle or an ellipse, depending on the position of the source as the Earth rotates around its axis. If one varies the distance between the telescopes, one will get a series of circles or ellipses on the sky. As we can see from Fig. 3.29, one does not have to cover all the spacings between the telescopes, because any antenna combination which has the same relative distance will describe the same path on the sky. In this way one can synthesise an antenna, a filled aperture, with a size equal to the maximum spacing between the telescopes. Interferometers working according to this principle are called *aperture syn-*

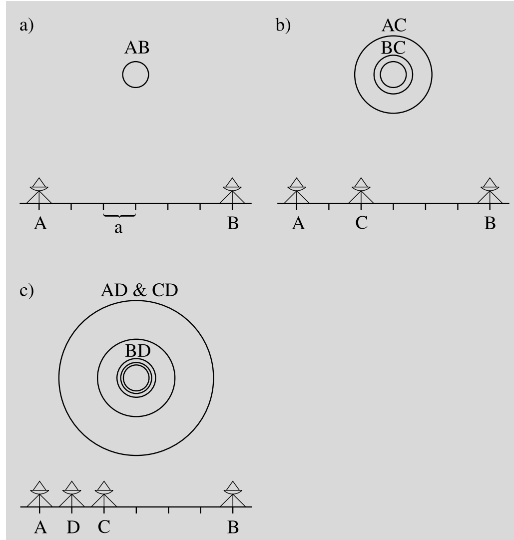


Fig. 3.33 To illustrate the principle of aperture synthesis, let us consider an east–west oriented interferometer pointed towards the celestial north. Each antenna is identical, has a diameter D and operates at a wavelength λ. The minimum spacing between each antenna element is a, and the maximum spacing is $6a$. In (**a**) there are only two antennas, A and B, displaced by the maximum spacing $6a$. When the earth rotates, antennas A and B will, in the course of 12 hours, track a circle on the plane of the sky with a diameter $\lambda/(6a)$, the maximum resolution that can be achieved with this interferometer. In (**b**) the antenna C is added to the interferometer, thus providing two more baselines, which track the circles AC and BC with radii of $\lambda/(2a)$ and $\lambda/(4a)$, respectively. In (**c**) there is still another antenna D added to the interferometer. In this case two of the baselines are equal, AD and CD, and therefore only two new circles are covered on the plane of the sky. By adding more interferometer elements, one can fill in the missing parts within the primary beam, i.e. the beam of one single dish, and thus obtain a full coverage of the beam. The resolution of the system depends on the maximum separation of the antennas. It is evident from (**c**), that not all of the antenna positions are needed to provide all the different spacings; some antenna spacings will in such a case be equal and therefore provide no additional information. It is essential to position the antennas in such a way that as many different baselines as possible are formed. The observed values and the image of the target are related by a Fourier transform

Fig. 3.34 The VLA at Socorro, New Mexico, is a synthesis telescope consisting of 27 movable antennas

thesis telescopes. If one covers all the spacings up to the maximum baseline, the result will be an accurate map of the source over the primary beam of an individual antenna element. Aperture synthesis telescopes therefore produce an image of the sky, i.e. a "*radio photograph*".

A typical aperture synthesis telescope consists of one fixed telescope and a number of movable telescopes, usually located on an east–west track, although T or Y configurations are also quite common. The number of telescopes used determines how fast one can synthesise a larger disk, because the number of possible antenna combinations increases as $n(n-1)$, where n is the number of telescopes. It is also possible to synthesise a large telescope with only one fixed and one movable telescope by changing the spacing between the telescopes every 12 hours, but then a full aperture synthesis can require several months of observing time. In order for this technique to work, the source must be constant, i.e. the signal cannot be time variable during the observing session.

The most efficient aperture synthesis telescopes at present are the VLA (Very Large Array) in New Mexico, USA (Fig. 3.34) and ALMA in Chile.

VLA consists of 27 paraboloid antennas, each with a diameter of 25 m, which are located on a Y-shaped track. The Y-formation was chosen because it provides a full aperture synthesis in 8 hours. Each antenna can be moved by a specially built carrier, and the locations of the telescopes are chosen to give optimal spacings for each configuration. In the largest configuration each arm is about 21 km long, thereby resulting in an antenna with an effective diameter of 35 km. If the VLA is used in its largest configuration and at its highest frequency, 23 GHz (1.3 cm), the resolution achieved is 0.1 arc second, clearly superior to any optical telescope. Similar resolution can also be obtained with the British MERLIN telescope, where already existing telescopes have been coupled together by radio links. Other well-known synthesis telescopes are the Cambridge 5 km array in Britain and the Westerbork array in the Netherlands, both located on east–west tracks.

Even higher resolution can be obtained with an extension of the aperture synthesis technique, called *VLBI* (Very Long Baseline Interferometry). With the VLBI technique the spacing between the antennas is restricted only by the size of the Earth. Some of the antennas can even be on satellites, in which case the baseline can be arbitrarily long. The antennas are all pointed towards the same source. The signal is recorded together with accurate timing signals from atomic clocks. The data files are correlated against each other, resulting in maps similar to those obtained with a normal aperture synthesis telescope. With VLBI techniques it is possible to achieve resolutions of $0.00001''$ if one of the telescopes is orbiting the Earth. Because interferometry is very sensitive to the distance between the telescopes, the VLBI technique also provides one of the most accurate methods to measure distances. Currently one can measure distances with an accuracy of a few millimetres on intercontinental baselines. This is utilised in geodetic VLBI experiments, which study continental drift and polar motion as a function of time.

In radio astronomy the maximum size of single antennas has also been reached. The trend is to build synthesis antennas. In the 1990's The United States built a chain of antennas extending across the whole continent, and the Australians have constructed a similar, but north–south antenna chain across their country.

More and more observations are being made in the submillimetre region. The disturbing effect of atmospheric water vapour becomes more serious at shorter wavelengths; thus, submillimetre telescopes must be located on mountain tops, like optical telescopes. All parts of the mirror are actively controlled in order to accurately maintain the proper form like in the new optical telescopes. Several new submillimetre telescopes are under construction.

3.5 Other Wavelength Regions

The Earth receives electromagnetic radiation from the sky at all wavelengths However, as mentioned in Sect. 3.1, not all radiation reaches the ground. The wavelength regions absorbed by the atmosphere have only been studied more extensively since the 1970's, using Earth-orbiting satellites (Fig. 3.37). Besides the optical and radio regions, there are only some narrow wavelength ranges in the infrared that can be observed from high mountain tops.

The first observations in each new wavelength region were usually carried out from balloons, but not until rockets came into use could observations be made from outside the atmosphere. The first actual observations of an X-ray source, for instance, were made on a rocket flight in June 1962, when the detector rose above the atmosphere for about 6 minutes. Finally, Earth orbiting satellites have made it possible to map the whole sky in the wavelength regions invisible from the ground.

Gamma Radiation Gamma ray astronomy studies radiation quanta with energies of 10^5–10^{14} eV. The boundary between gamma and X-ray astronomy, 10^5 eV, corresponds to a wavelength of 10^{-11} m. The boundary is not fixed; the regions of hard (= high-energy) X-rays and soft gamma rays partly overlap.

While ultraviolet, visible and infrared radiation are all produced by changes in the energy states of the electron envelopes of atoms, gamma and hard X-rays are produced by transitions in atomic nuclei or in mutual interactions of elementary particles. Thus observations of the shortest wavelengths give information on processes different from those giving rise to longer wavelengths.

The first observations of gamma sources were obtained at the end of the 1960's, when a device in the OSO 3 satellite (Orbiting Solar Observatory) detected gamma rays from the Milky Way. Later on, many satellites have been especially designed for gamma astronomy. The most recent satellites include the Compton Gamma Ray Observatory, operating in 1991–2000, the European Integral, launched in 2002, and the American Fermi, originally called GLAST.

The quanta of gamma radiation have energies a million times greater than those of visible light, and therefore they cannot be observed with similar detectors. Observations are made with various *scintillation detectors*, usually composed of several layers of detector plates, where gamma

Fig. 3.35 The MAGIC telescope on La Palma in the Canary islands observes the Cherenkov radiation produced by gamma rays. (Photo Robert Wagner, Max Planck Institut für Physik)

radiation is transformed by the photoelectric effect into visible light, detectable by photomultipliers.

The energy of a gamma quantum can be determined from the depth to which it penetrates the detector. Analysing the trails left by the quanta gives information on their approximate direction. The field of view is limited by the grating. The directional accuracy is low, and in gamma astronomy the resolution is still far below that of other wavelength regions.

Gamma radiation cannot pass through the atmosphere, yet it has effects that can be observed with ground based instruments. When a high-energy photon enters the atmosphere it can create particle-antiparticle pairs. These particles can get such a high energy that they move faster than the local speed of light in the atmosphere. Such particles radiate Čherenkov radiation that can be detected as visible light (Fig. 3.35).

X-rays The observational domain of X-ray astronomy includes the energies between 10^2 and 10^5 eV, or the wavelengths 10–0.01 nm. The regions 10–0.1 nm and 0.1–0.01 nm are called *soft* and *hard X-rays*, respectively. X-rays were discovered in the late 19th century. Systematic studies of the sky at X-ray wavelengths only became possible in the 1970's with the advent of satellite technology.

The first all-sky mapping was made in the early 1970's by SAS 1 (Small Astronomical Satellite), also called Uhuru. At the end of the 1970's, two High-Energy Astronomy Observatories, HEAO 1 and 2 (the latter called Einstein), mapped the sky with much higher sensitivity than Uhuru.

The Einstein Observatory was able to detect sources about a thousand times fainter than earlier X-ray telescopes. In optical astronomy, this would correspond to a jump from a 15 cm reflector to a 5 m telescope. Thus X-ray astronomy has developed in 20 years as much as optical astronomy in 300 years.

The latest X-ray satellites have been the American Chandra and the European XMM-Newton, both launched in 1999.

Besides satellites mapping the whole sky, there have been several satellites observing the X-ray radiation of the Sun. The most recent ones are the Japanese Yohkoh and Hinode and the American RHESSI.

The first X-ray telescopes used detectors similar to those in gamma astronomy. Their directional accuracy was never better than a few arc minutes. The more precise X-ray telescopes utilise the principle of *grazing reflection* (Fig. 3.36). An X-ray hitting a surface perpendicularly is not reflected, but absorbed. If, however, X-rays meet the mirror nearly parallel to its surface, just grazing it, a high quality surface can reflect the ray.

The mirror of an X-ray reflector is on the inner surface of a slowly narrowing cone. The outer part of the surface is a paraboloid and the inner part a hyperboloid. The rays are reflected by both surfaces and meet at a focal plane. In practice, several tubes are installed one within another. For instance, the four cones of the Einstein Observatory had as much polished optical surface as a normal telescope with a diameter of 2.5 m. The resolution in X-ray telescopes is of the order of a few arc seconds and the field of view about 1 deg.

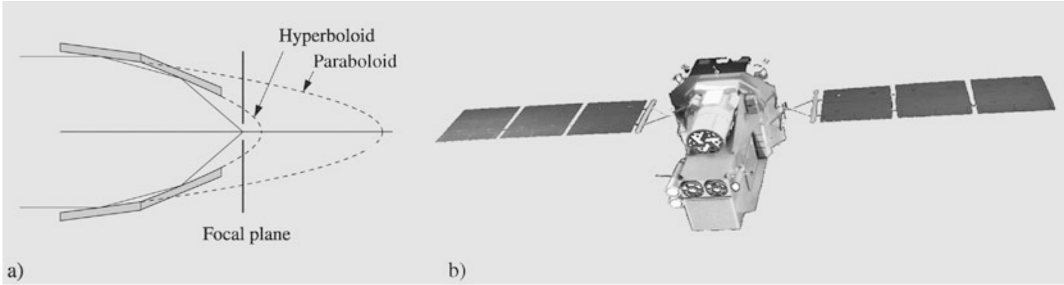


Fig. 3.36 X-rays are not reflected by an ordinary mirror, and the principle of grazing reflection must be used for collecting them. Radiation meets the paraboloid mirror at a very small angle, is reflected onto a hyperboloid mirror and further to a focal point. In practice, several mirrors are placed one inside another, collecting radiation in a common focus

The detectors in X-ray astronomy are usually *Geiger–Müller counters*, *proportional counters* or scintillation detectors. Geiger–Müller and proportional counters are boxes filled with gas. The walls form a cathode, and an anode wire runs through the middle of the box; in more accurate counters, there are several anode wires. An X-ray quantum entering the box ionises the gas, and the potential difference between the anode and cathode gives rise to a current of electrons and positive ions. In recent imaging instruments (like the ACIS in Chandra) also CCD cameras are used.

Ultraviolet Radiation Between X-rays and the optical region lies the domain of ultraviolet radiation, with wavelengths between 10 and 400 nm. Most ultraviolet observations have been carried out in the *soft UV* region, at wavelengths near those of optical light, since most of the UV radiation is absorbed by the atmosphere. The wavelengths below 300 nm are completely blocked out. The short wavelength region from 10 to 91.2 nm is called the *extreme ultraviolet* (EUV, XUV).

Extreme ultraviolet was one of the last regions of the electromagnetic radiation to be observed systematically. The reason for this is that the absorption of interstellar hydrogen makes the sky practically opaque at these wavelengths. The visibility in most directions is limited to some hundred light years in the vicinity of the Sun. In some directions, however, the density of the interstellar gas is so low that even extragalactic objects can be seen.

The first dedicated EUV satellite was the Extreme Ultraviolet Explorer (EUVE), operating in 1992–2000. It observed about a thousand EUV sources. In EUV grazing reflection telescopes similar to those used in X-ray astronomy are employed.

In nearly all branches of astronomy important information is obtained by observations of ultraviolet radiation. Many emission lines from stellar chromospheres or coronas, the Lyman lines of atomic hydrogen, and most of the radiation from hot stars are found in the UV domain. In the *near-ultraviolet*, telescopes can be made similar to optical telescopes and, equipped with a photometer or spectrometer, installed in a satellite orbiting the Earth.

The most effective satellites in the UV have been the European TD-1, the American Orbiting Astronomical Observatories OAO 2 and 3 (Copernicus), and the International Ultraviolet Explorer IUE. The instruments of the TD-1 satellite included both a photometer and a spectrometer. The satellite measured the magnitudes of over 30,000 stars in four different spectral regions between 135 and 274 nm, and registered UV spectra from over 1000 stars. The OAO satellites were also used to measure magnitudes and spectra, and OAO 3 worked for over eight years.

The IUE satellite, launched in 1978, was one of the most successful astronomical satellites. IUE had a 45 cm Ritchey-Chrétien telescope with an aperture ratio of $f/15$ and a field of view of

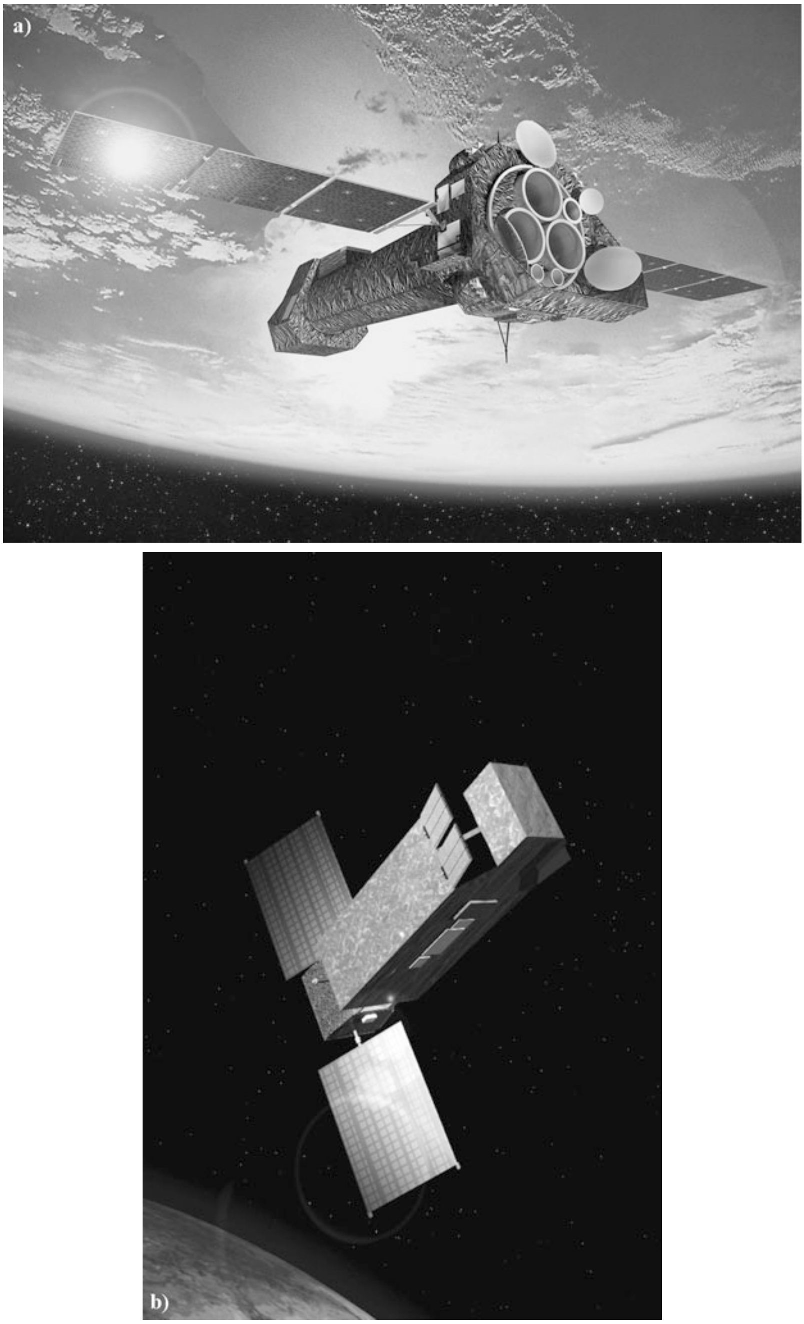


Fig. 3.37 (**a**) The European X-ray satellite XMM-Newton was launched in 1999. (Drawing D. Ducros, XMM Team, ESA.) (**b**) FUSE satellite has photographed far ultraviolet objects from Earth orbit since 1999. (Graphics NASA/JHU Applied Physics Laboratory)

16 arc minutes. The satellite had two spectrographs to measure spectra of higher or lower resolution in wavelength intervals of 115–200 nm or 190–320 nm. For registration of the spectra, a Vidicon camera was used. IUE was different from earlier satellites, since it could be used

almost like ground based telescopes. The observer could follow the observations and change them in real time. IUE worked on the orbit for 20 years.

The Hubble space telescope can also make observations in the ultraviolet. Some ultraviolet telescopes have also been on shuttle flight in the 1990's.

Infrared Radiation Radiation with longer wavelengths than visible light is called infrared radiation. This region extends from about 1 micrometre to 1 millimetre, where the radio region begins. Sometimes the *near-infrared*, at wavelengths below 5 μm, and the submillimetre domain, at wavelengths between 0.1 and 1 mm, are considered separate wavelength regions.

In infrared observations radiation is collected by a telescope, as in the optical region. The incoming radiation consists of radiation from the object, from the background and from the telescope itself. Both the source and the background must be continually measured, the difference giving the radiation from the object. The background measurements can be made with a Cassegrain secondary mirror oscillating between the source and the background at a rate of, say, 100 oscillations per second, and thus the changing background can be eliminated (Fig. 3.38). To register the measurements, semiconductor detectors are used. The detector must always be cooled to minimise its own thermal radiation. Sometimes the whole telescope is cooled.

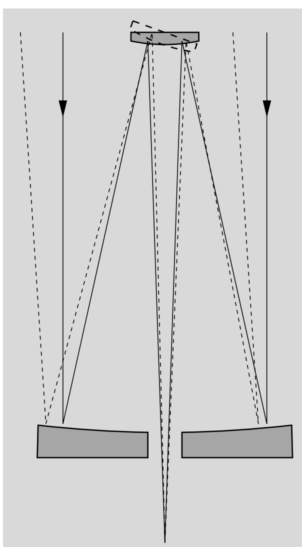


Fig. 3.38 Refractors are not suitable for infrared telescopes, because infrared radiation cannot penetrate glass. The Cassegrain reflectors intended especially for infrared observations have secondary mirrors nodding rapidly back and forth between the object and the background near the object. By subtracting the brightness of the background from the brightness of the object, the background can be eliminated

Infrared observatories have been built on high mountain tops, where most of the atmospheric water vapour remains below. Some favourable sites are, e.g. Mauna Kea on Hawaii, Mount Lemon in Arizona and Pico del Teide on Tenerife. For observations in the far-infrared these mountains are not high enough; these observations are carried out, e.g. on aeroplanes. One of the best-equipped planes is the Kuiper Airborne Observatory KAO, named after the well-known planetary scientist Gerard Kuiper.

Balloons and satellites are also used for infrared observations. The first powerful infrared observatory was the InfraRed Astronomy Satellite IRAS, built by the United States and Holland and launched in 1983. It made observations for eight months and mapped the sky at four wavelengths (12, 25, 60 and 100 μm), and found over 200,000 new infrared sources.

The European Infrared Space Observatory ISO, made more detailed observations of thousands of infrared objects in 1996–1998. Since 2003 Spitzer (originally SIRTF, Space InfraRed Telescope Facility). has been mapping the infrared sky. In 2009 ESA launched its own Herschel satellite (Fig. 3.39), which operated till 2013, when the liquid helium used for cooling was exhausted.

Infrared satellites have also been used to map the afterglow of the big bang of the universe. A very successful satellite was the 1989 launched COBE (Cosmic Background Explorer), which mapped the background radiation in submillimetre and infrared wavelengths. The Microwave Anisotropy Probe (MAP) has continued the work of COBE, starting in 2001. In 2009 the even more powerful European *Planck* satellite was launched.

Fig. 3.39 The European Herschel launched in 2009 was the most efficient recent infrared satellite. The diameter of its main mirror, 3.5 metres, was bigger than in any earlier astronomical satellite. The final shape to the highly curved mirror was given by the Opteon company at the Tuorla observatory in Finland. The observations ended in 2103 when the liquid helium used to cool down the camera ran out. (Photos ESA, Rami Rekola)

3.6 Other Forms of Energy

Besides electromagnetic radiation, energy arrives from space in other forms: particles (*cosmic rays*, *neutrinos*) and *gravitational radiation*.

Cosmic Rays Cosmic rays, consisting of electrons and totally ionised nuclei of atoms, are received in equal amounts from all directions. Their incoming directions do not reveal their origin, since cosmic rays are electrically charged; thus their paths are continually changed when they move through the magnetic fields of the Milky Way. The high energies of cosmic rays mean that they have to be produced by high-energy phenomena like supernova explosions. The majority of cosmic rays are protons (nearly 90 %) and helium nuclei (10 %), but some are heavier nuclei; their energies lie between 10^8 and 10^{20} eV.

Fig. 3.39 (Continued)

The most energetic cosmic rays give rise to *secondary radiation*, mainly muons, when they hit molecules of the atmosphere. This secondary radiation can be observed from the ground. Just like particles produces by gamma rays particles of the secondary radiation can move faster than light in the atmosphere emitting Čerenkov radiation. Thus the cosmic rays can also be detected by their Čerenkov radiation.

The Pierre Auger telescope in Argentina observes both secondary particles and Čerenkov radiation to detect cosmic rays with energies at least of the order of 10^{18} eV. Since only about one such particle hits an area of one square kilometre in a century, the detectors have been positioned in an area of 3000 square kilometres.

Primary cosmic rays can only be directly observed outside the atmosphere. The detectors used to observe cosmic rays are similar to those used in particle physics. Since Earth-based accelerators reach energies of only about 10^{12} eV, cosmic rays offer an excellent "natural" laboratory for particle physics. Many satellites and spacecraft have detectors for cosmic rays.

Neutrinos Neutrinos are elementary particles with no electric charge. Earlier it was thought that they are massless, but currently they seem to have a tiny rest mass. They interact with other matter only through the weak nuclear force, which is very much weaker than the electromagnetic or strong nuclear force.

Most neutrinos are produced in nuclear reactions within stars and in supernova explosions. Since they react very weakly with other matter, they escape directly from the stellar interior.

Due to their weak interaction neutrinos are very difficult to observe; trillions of them pass through this page every second without any noticeable consequences.

The first method of detection was the radiochemical method. As a reactive agent, e.g. tetrachloroethene (C_2Cl_4) can be used. When a neutrino hits a chlorine atom, the chlorine is transformed into argon, and an electron is freed:

$$^{37}\mathrm{Cl} + \nu \rightarrow {}^{37}\mathrm{Ar} + \mathrm{e}^- .$$

The argon atom is radioactive and can be observed. Instead of chlorine, lithium and gallium might be used to detect neutrinos. The first gallium detectors have been running in Italy and Russia from the end of the 1980's.

Another observation method is based on the Čerenkov radiation produced by neutrinos in ex-

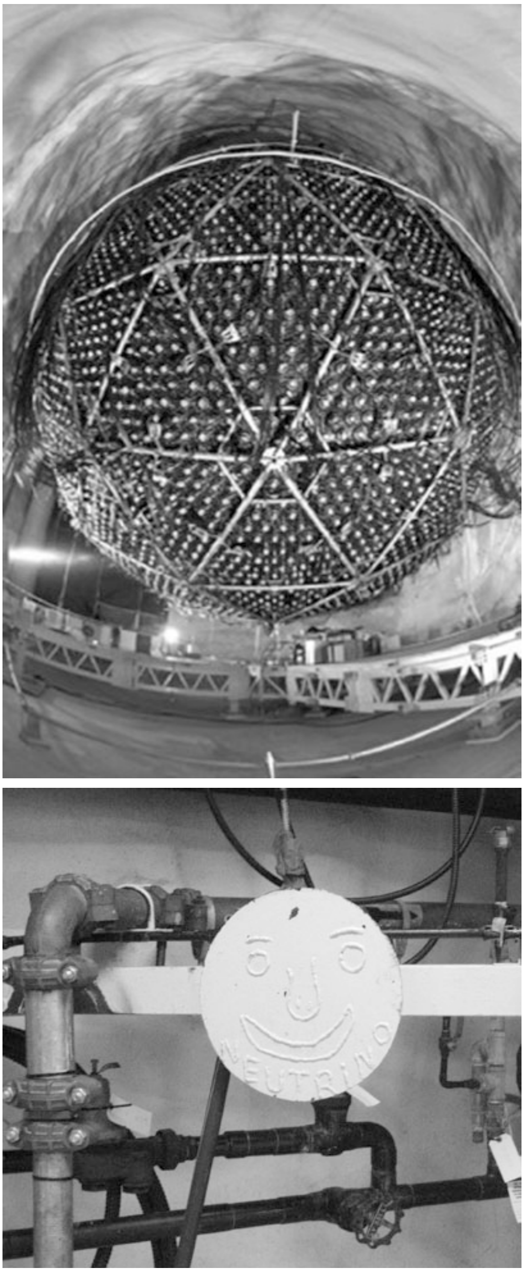


Fig. 3.40 The Sudbury neutrino observatory (SNO) is at the depth of two kilometres in a still operational mine. The spherical vessel contains liquid where neutrinos produce Cherenkov radiation. 9600 photomultipliers are used to register the flashes of the radiation. Earlier the liquid was heavy water but that was later replaced with alkyl benzene. The observations have already solved the solar neutrino problem (Sect. 13.1). The new SNO+ will be able to observe neutrinos of even lower energies. (Photos SNOlab, H. Karttunen)

tremely pure water. The flashes of light are registered with photomultipliers, and thus it is possible to find out the direction of the radiation. This method is used e.g. in the Japanese Kamiokande detector and in the Sudbury neutrino observatory in Canada (Fig. 3.40).

The largest neutrino detector is the IceCube in the Antarctic, finished in 2010. It consists of one cubic kilometre of ice.

Neutrino detectors must be located deep under the ground to protect them from the secondary radiation caused by cosmic rays. Also radioactive decay inside the Earth cases background noise.

The detectors have observed neutrinos from the Sun, and the Supernova 1987A in the Large Magellanic Cloud was also observed in 1987. In the future, also the dark matter and other neutrino producing phenomena will be studied.

Gravitational Radiation Gravitational astronomy is as young as neutrino astronomy. The first attempts to measure gravitational waves were made in the 1960's. Gravitational radiation is emitted by accelerating masses, just as electromagnetic radiation is emitted by electric charges in accelerated motion. Detection of gravitational waves is very difficult, and they have been observed only very recently.

The first type of gravitational wave antenna was the *Weber cylinder*. It is an aluminium cylinder which starts vibrating at its proper frequency when hit by a gravitational pulse. The distance between the ends of the cylinder changes by about 10^{-17} m, and the changes in the length are studied by strain sensors welded to the side of the cylinder.

Another type of modern gravity radiation detectors measures "spatial strain" induced by gravity waves and consists of two sets of mirrors in directions perpendicular to each other (Michelson interferometer), or one set of parallel mirrors (Fabry–Perot interferometer). The relative distances between the mirrors are monitored by laser interferometers. If a gravity pulse passes the detector, the distances change and the changes can be measured. The longest baseline between the mirrors is in the American LIGO (Laser Inter-

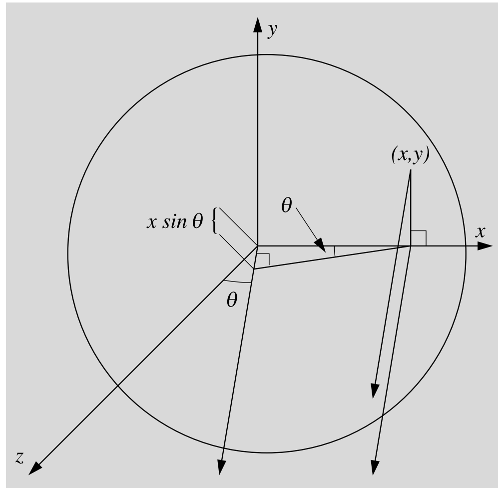
Fig. 3.41 The LIGO Livingston Observatory seen from the air. (Photo LIGO/Caltech)

ferometer Gravitational-wave Observatory) system (Fig. 3.41). LIGO made the first scientific observations in 2002. In early 2016 the first positive detection of gravitational waves was announced. The observation was made in September 14, 2015.

Although the gravitational detectors are isolated from their surroundings as carefully as possible, many perturbations can make them oscillate. A positive detection requires that the same phenomenon is observed by different instruments located far away from each others.

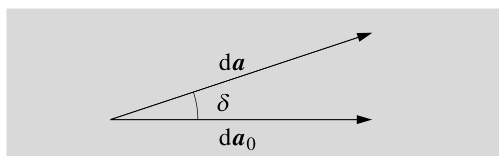
Box 3.1 (Diffraction by a Circular Aperture) Consider a circular hole of radius R in the xy plane. Coherent light enters the hole from the direction of the negative z axis (see figure). We consider light rays leaving the hole parallel to the xz plane forming an angle θ with the z axis. The light waves interfere on a screen far away. The phase difference between a wave through a point (x, y) and a wave going through the centre of the hole can be calculated from the different path lengths $s = x\sin\theta$:

$$\delta = \frac{s}{\lambda}2\pi = \frac{2\pi\sin\theta}{\lambda}x \equiv kx.$$



Thus, the phase difference δ depends on the x coordinate only. The sum of the amplitudes of the waves from a small surface element is proportional to the area of the element $\mathrm{d}x\,\mathrm{d}y$. Let the amplitude coming through the centre of the hole be $\mathrm{d}\boldsymbol{a}_0 = \mathrm{d}x\,\mathrm{d}y\hat{\boldsymbol{i}}$. The amplitude coming from the point (x, y) is then

$$\mathrm{d}\boldsymbol{a} = \mathrm{d}x\,\mathrm{d}y\left(\cos\delta\,\hat{\boldsymbol{i}} + \sin\delta\,\hat{\boldsymbol{j}}\right).$$



We sum up the amplitudes coming from different points of the hole:

$$\begin{aligned}\boldsymbol{a} &= \int_{\text{Aperture}} \mathrm{d}\boldsymbol{a} \\ &= \int_{x=-R}^{R}\int_{y=-\sqrt{R^2-x^2}}^{\sqrt{R^2-x^2}} \left(\cos kx\,\hat{\boldsymbol{i}} + \sin kx\,\hat{\boldsymbol{j}}\right)\mathrm{d}y\,\mathrm{d}x \\ &= 2\int_{-R}^{R}\sqrt{R^2-x^2}\left(\cos kx\,\hat{\boldsymbol{i}} + \sin kx\,\hat{\boldsymbol{j}}\right)\mathrm{d}x.\end{aligned}$$

Since sine is an odd function ($\sin(-kx) = -\sin(kx)$), we get zero when we integrate the second term. Cosine is an even function, and so

$$a \propto \int_0^R \sqrt{R^2-x^2}\cos kx\,\mathrm{d}x.$$

We substitute $x = Rt$ and define $p = kR = (2\pi r \sin\theta)/\lambda$, thus getting

$$a \propto \int_0^1 \sqrt{1-t^2} \cos pt \, \mathrm{d}t.$$

The zero points of the intensity observed on the screen are obtained from the zero points of the amplitude,

$$J(p) = \int_0^1 \sqrt{1-t^2} \cos pt \, \mathrm{d}t = 0.$$

Inspecting the function $J(p)$, we see that the first zero is at $p = 3.8317$, or

$$\frac{2\pi R \sin\theta}{\lambda} = 3.8317.$$

The radius of the diffraction disk in angular units can be estimated from the condition

$$\sin\theta = \frac{3.8317\lambda}{2\pi R} \approx 1.22 \frac{\lambda}{D},$$

where $D = 2R$ is the diameter of the hole.

In mirror telescopes diffraction is caused also by the support structure of the secondary mirror. If the aperture is more complex and only elementary mathematics is used calculations may become rather cumbersome. However, it can be shown that the diffraction pattern can be obtained as the Fourier transform of the aperture.

3.7 Examples

Example 3.1 The distance between the components of the binary star ζ Herculis is $1.38''$. What should the diameter of a telescope be to resolve the binary? If the focal length of the objective is 80 cm, what should the focal length of the eyepiece be to resolve the components, when the resolution of the eye is $2'$?

In the optical region, we can use the wavelength value of $\lambda \approx 550$ nm. The diameter of the objective is obtained from the equation for the resolution (3.4),

$$D \approx \frac{\lambda}{\theta} = \frac{550 \times 10^{-9}}{(1.38/3600) \times (\pi/180)} \text{ m} = 0.08 \text{ m} = 8 \text{ cm}.$$

The required magnification is

$$\omega = \frac{2'}{1.38''} = 87.$$

The magnification is given by

$$\omega = \frac{f}{f'},$$

and, thus, the focal length of the eyepiece should be

$$f' = \frac{f}{\omega} = \frac{80 \text{ cm}}{87} = 0.9 \text{ cm}.$$

Example 3.2 A telescope has an objective with a diameter of 90 mm and focal length of 1200 mm.

(a) What is the focal length of an eyepiece, the exit pupil of which is 6 mm (about the size of the pupil of the eye)?
(b) What is the magnification of such an eyepiece?
(c) What is the angular diameter of the Moon seen through this telescope and eyepiece?

(a) From Fig. 3.7 we get

$$L = \frac{f'}{f} D,$$

whence

$$f' = f\frac{L}{D} = 1200 \text{ mm} \frac{6 \text{ mm}}{90 \text{ mm}} = 80 \text{ mm}.$$

(b) The magnification is $\omega = f/f' = 1200 \text{ mm}/80 \text{ mm} = 15$.
(c) Assuming the angular diameter of the Moon is $\alpha = 31' = 0.52°$, its diameter through the telescope is $\omega\alpha = 7.8°$.

3.8 Exercises

Exercise 3.1 The Moon was photographed with a telescope, the objective of which had a diameter of 20 cm and focal length of 150 cm. The exposure time was 0.1 s.

(a) What should the exposure time be, if the diameter of the objective were 15 cm and focal length 200 cm?
(b) What is the size of the image of the Moon in both cases?
(c) Both telescopes are used to look at the Moon with an eyepiece the focal length of which is 25 mm. What are the magnifications?

Exercise 3.2 The radio telescopes at Amherst, Massachusetts, and Onsala, Sweden, are used as an interferometer, the baseline being 2900 km.

(a) What is the resolution at 22 GHz in the direction of the baseline?
(b) What should be the size of an optical telescope with the same resolution?