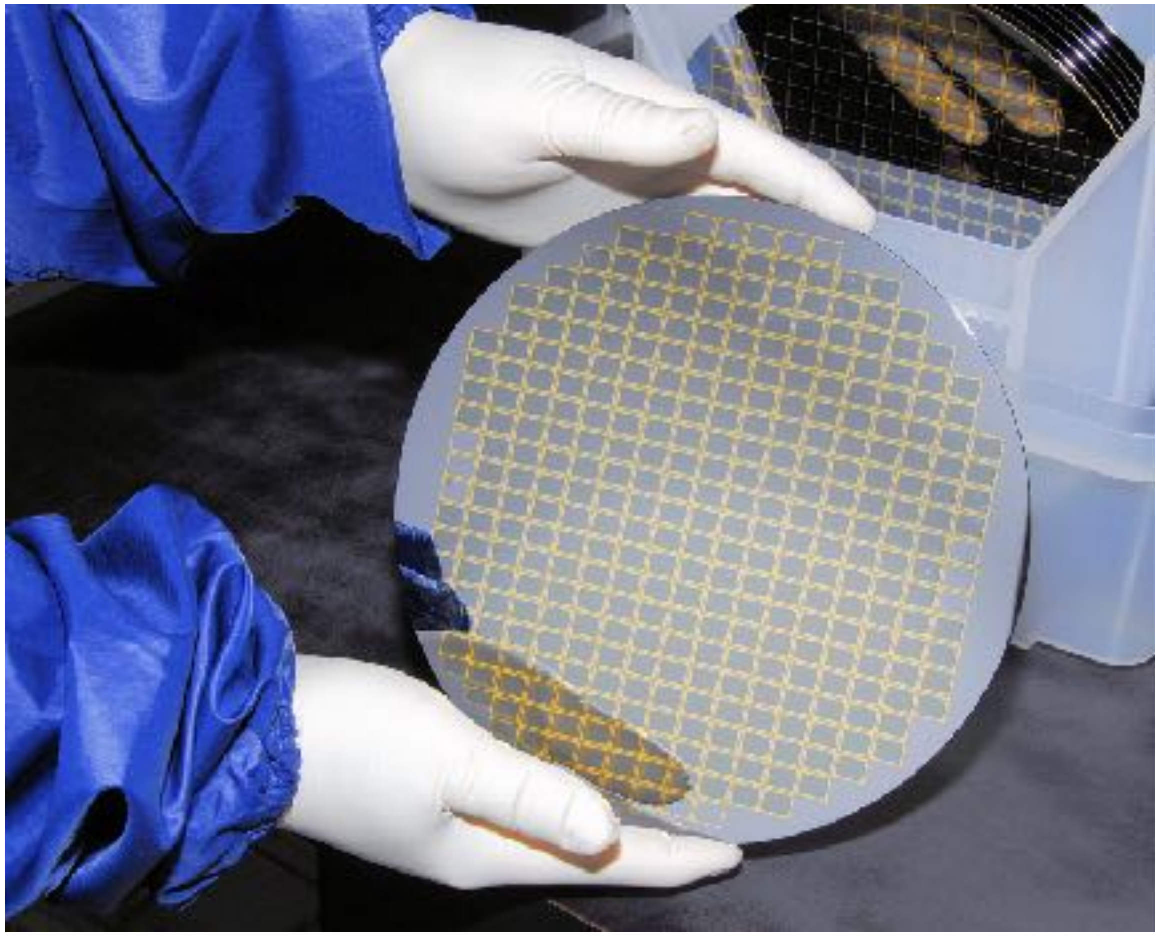
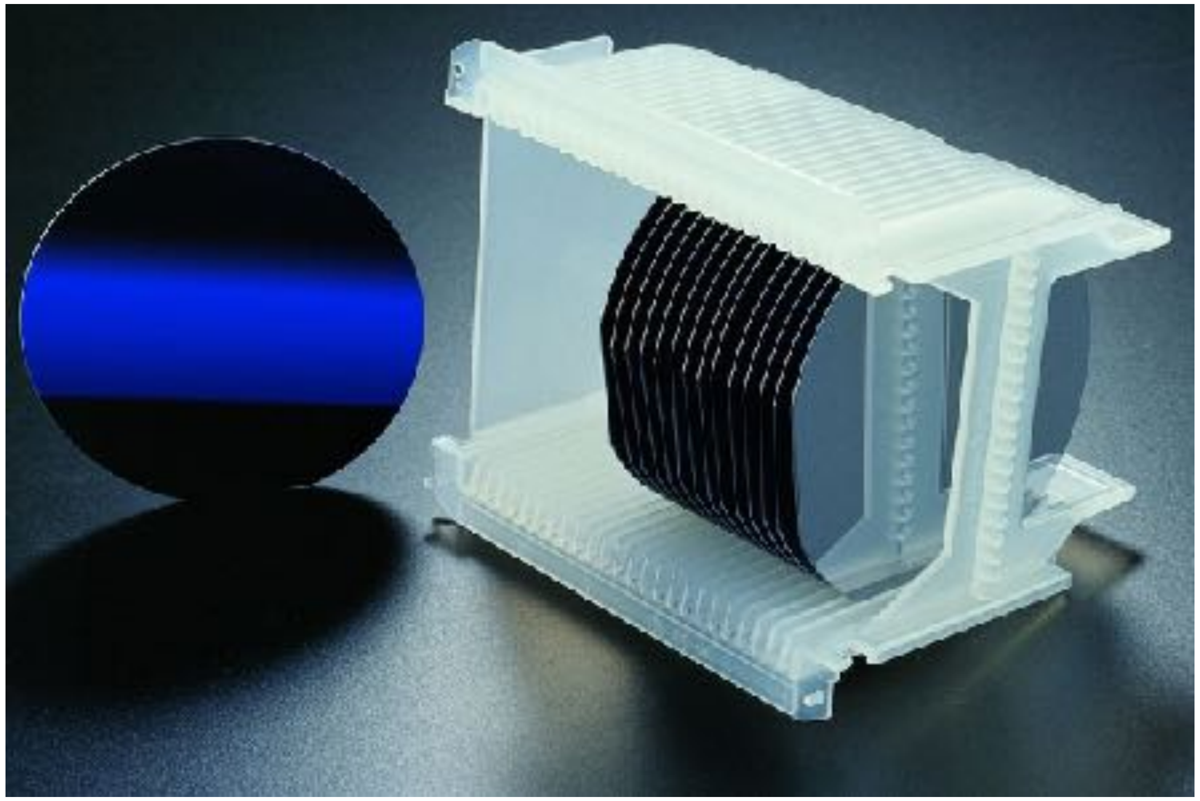
Astronomy 503
Observational Astronomy

Gautham Narayan

Lecture 09: CCDs, Imaging and Photometry

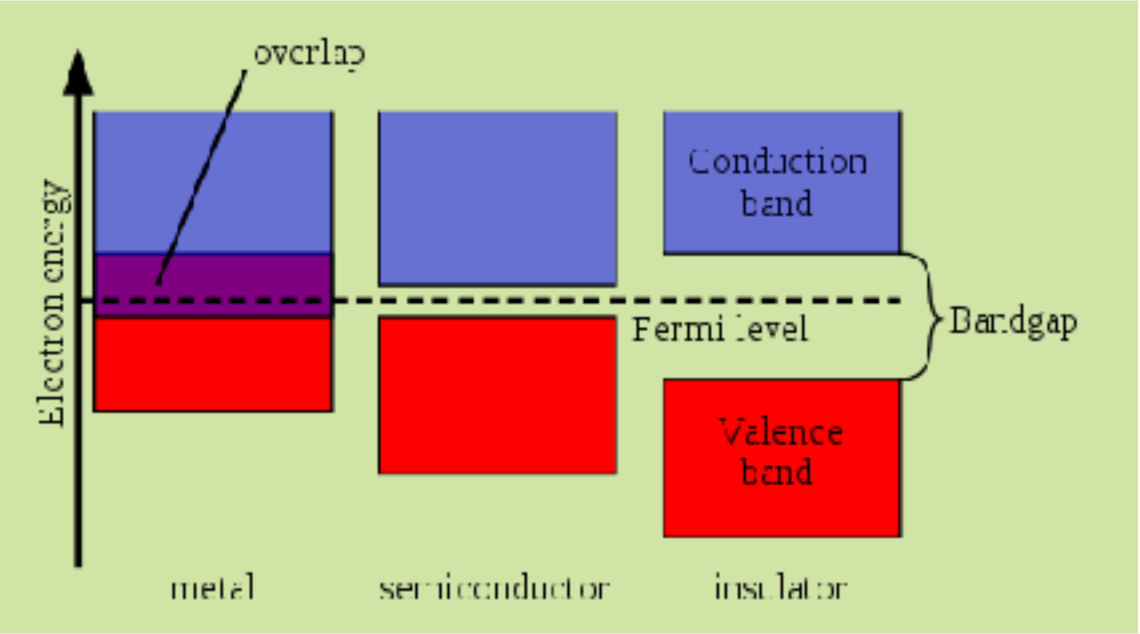
Silicon





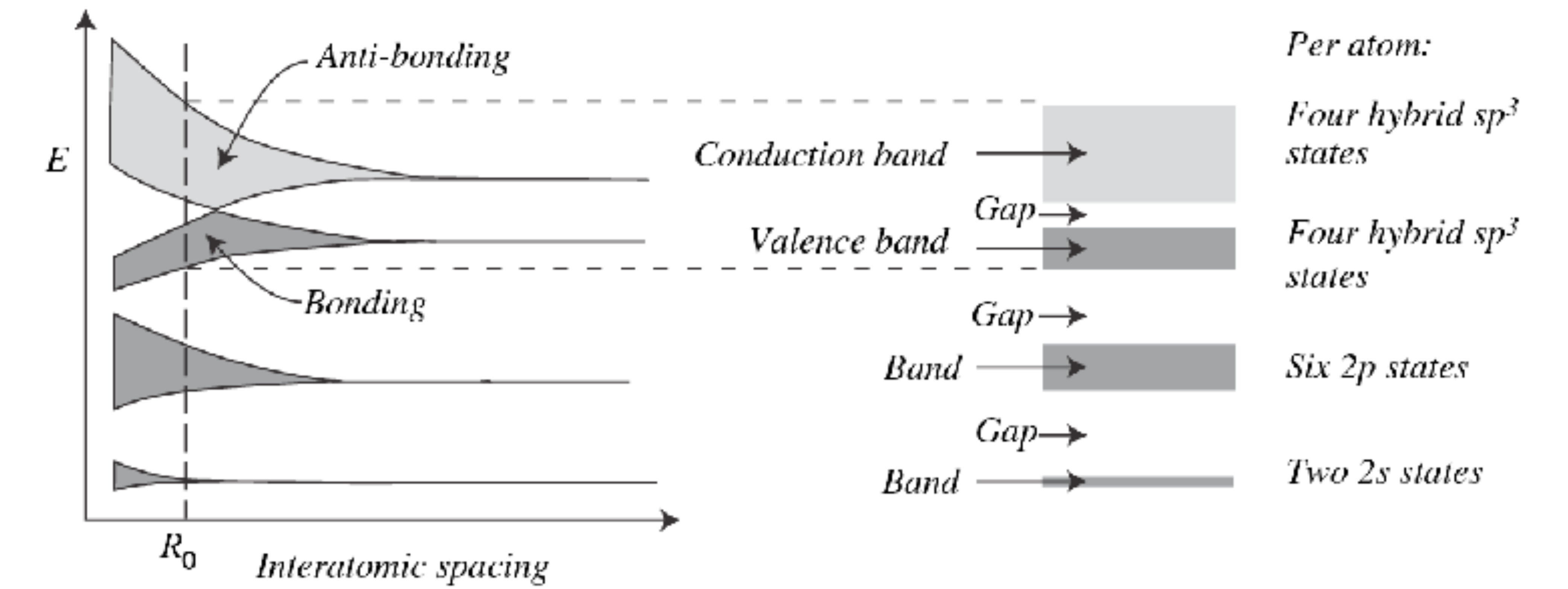
Semiconductor band structure

- Gap between valence and conduction band is the ***forbidden energy gap***:
  - For current to flow there must either be:
    - Electrons in the conduction band.
    - "Holes" in the valence band (previously occupied by electrons).
    - An applied electric field (a voltage across the semiconductor).
- ***Conductors:***
  - Overlapping valence and conduction bands.
- ***Insulators:***
  - Wide forbidden gap and no free electrons or holes.
- ***Semiconductors:***
  - Non-zero probability of electron-hole pairs forming and current conduction – properties lie between conductors and insulators.

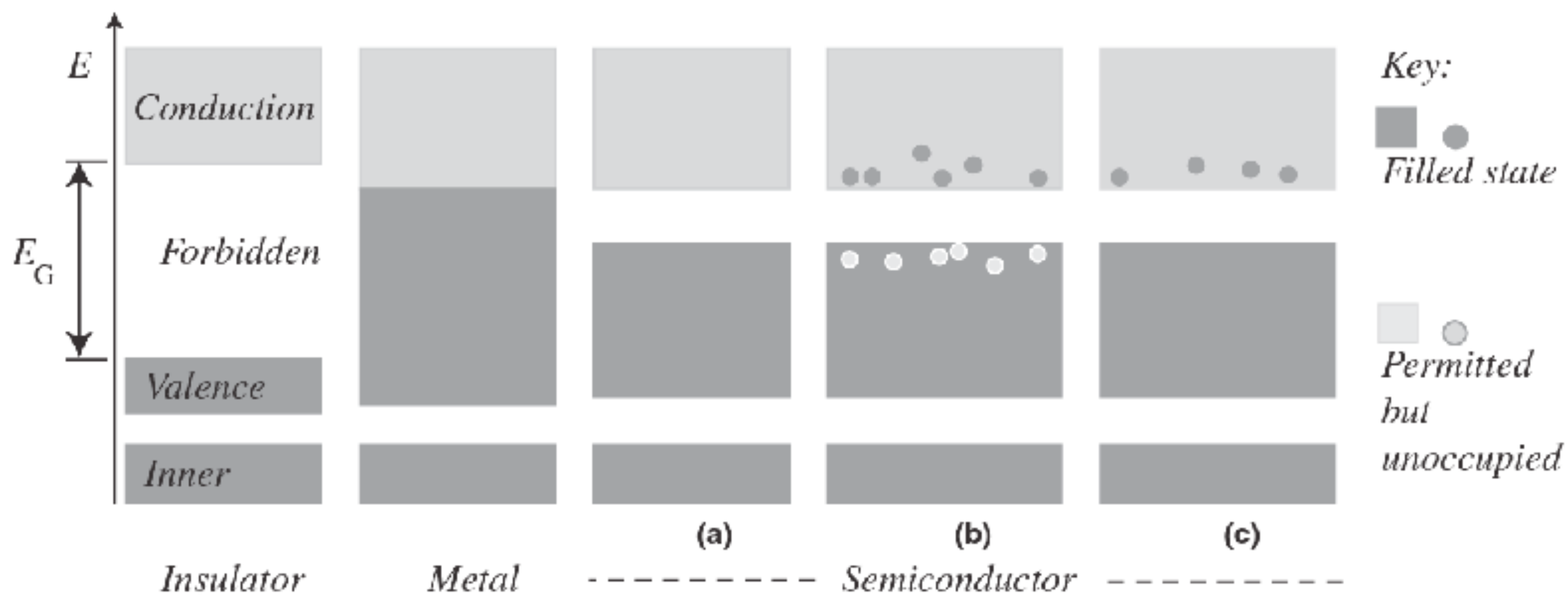


Fermi level: *energy at which there is a 50/50 probability of the electron energy state being occupied by an electron.*

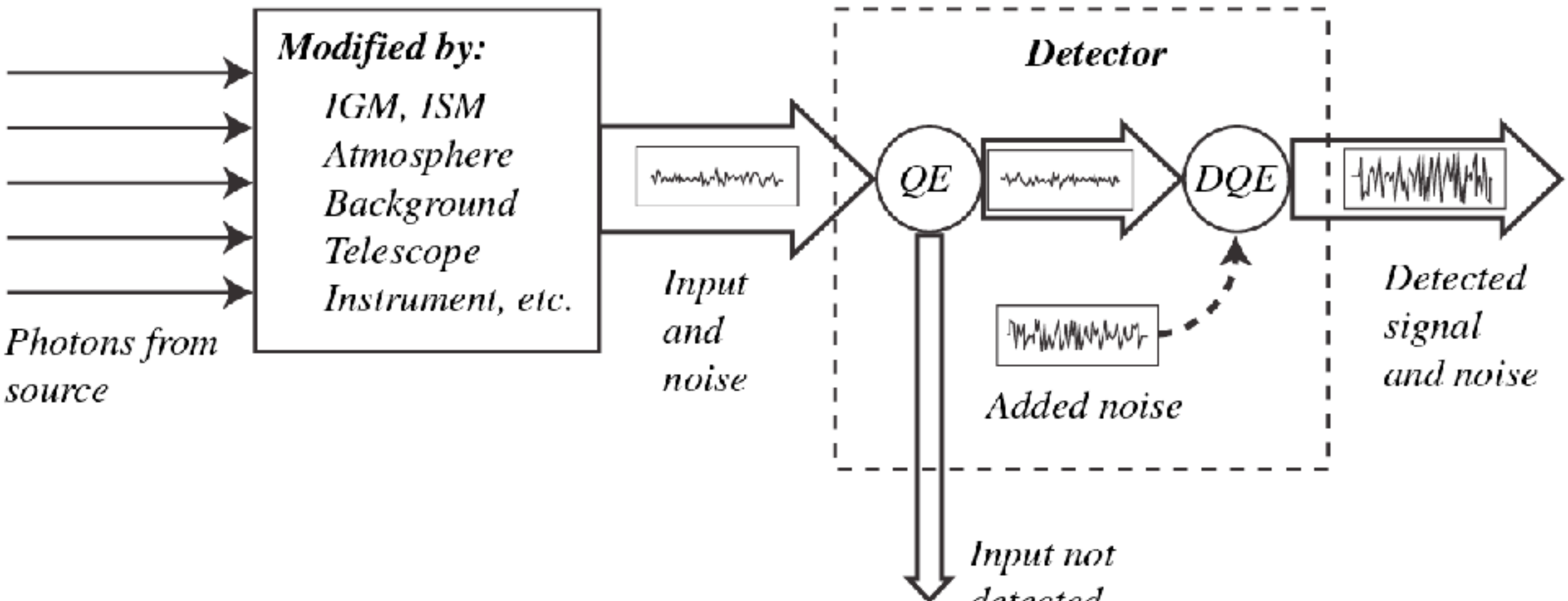
Crystals



E
Anti-bonding
Bonding
R0
Interatomic spacing
Conduction band
Gap
Valence band
Gap
Band
Gap
Band
Per atom:
Four hybrid sp3 states
Four hybrid sp3 states
Six 2p states
Two 2s states



E
Conduction
EG
Forbidden
Valence
Inner
(a)
(b)
(c)
Insulator
Metal
Semiconductor
Key:
Filled state
Permitted but unoccupied

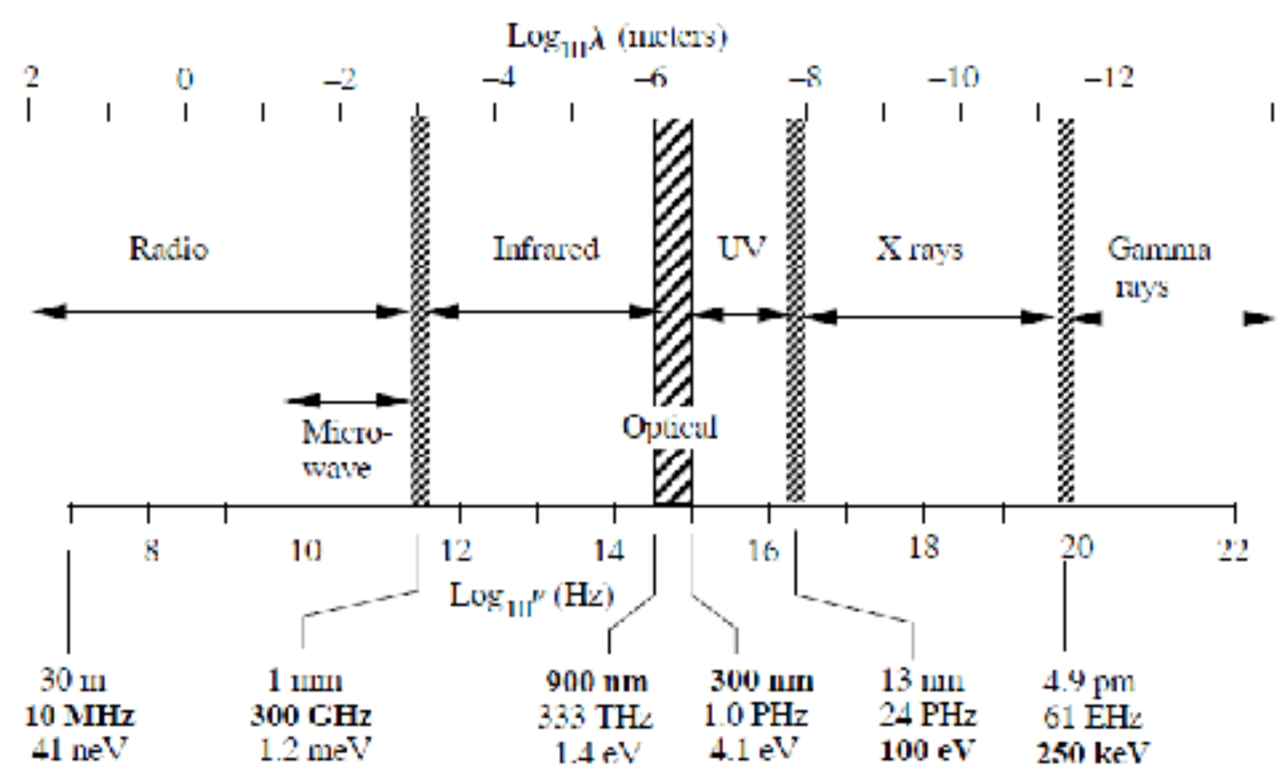


Modified by:
IGM, ISM
Atmosphere
Background
Telescope
Instrument, etc.
Photons from source
Input and noise
Detector
QE
DQE
Added noise
Input not detected
Detected signal and noise

Fundamental importance of detectors

- ***Detectors often define the science possible. They need to:***
  - Detect as many incident photons as possible.
  - Translate the detected photons efficiently to detector output.
  - Have a response matched to our wavelength or energy region of interest.
  - Ideally be configurable and adjustable – so that they can be used in optimized designs.
  - Generate as little noise as possible (i.e detector response in absence of astronomical source).
  - Be stable, and able to be calibrated.
  - And more…

Sub-band	Wavelength	Photon energy
Near-UV (NUV)	200 - 300 nm	4.1- 6.2 eV
UV	120 - 200 nm	6.2 -10.3 eV
Far-UV (FUV)	90 - 120 nm	10.3 – 13.8 eV
Extreme UV (EUV)	~5 - 90 nm	13.8 - 250 eV
X rays	0.01-10 nm	0.125 – 125 keV
Gamma rays	< 0.01 nm	> 125 keV

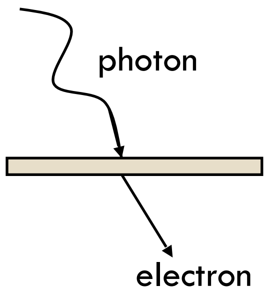


A perfect detector…

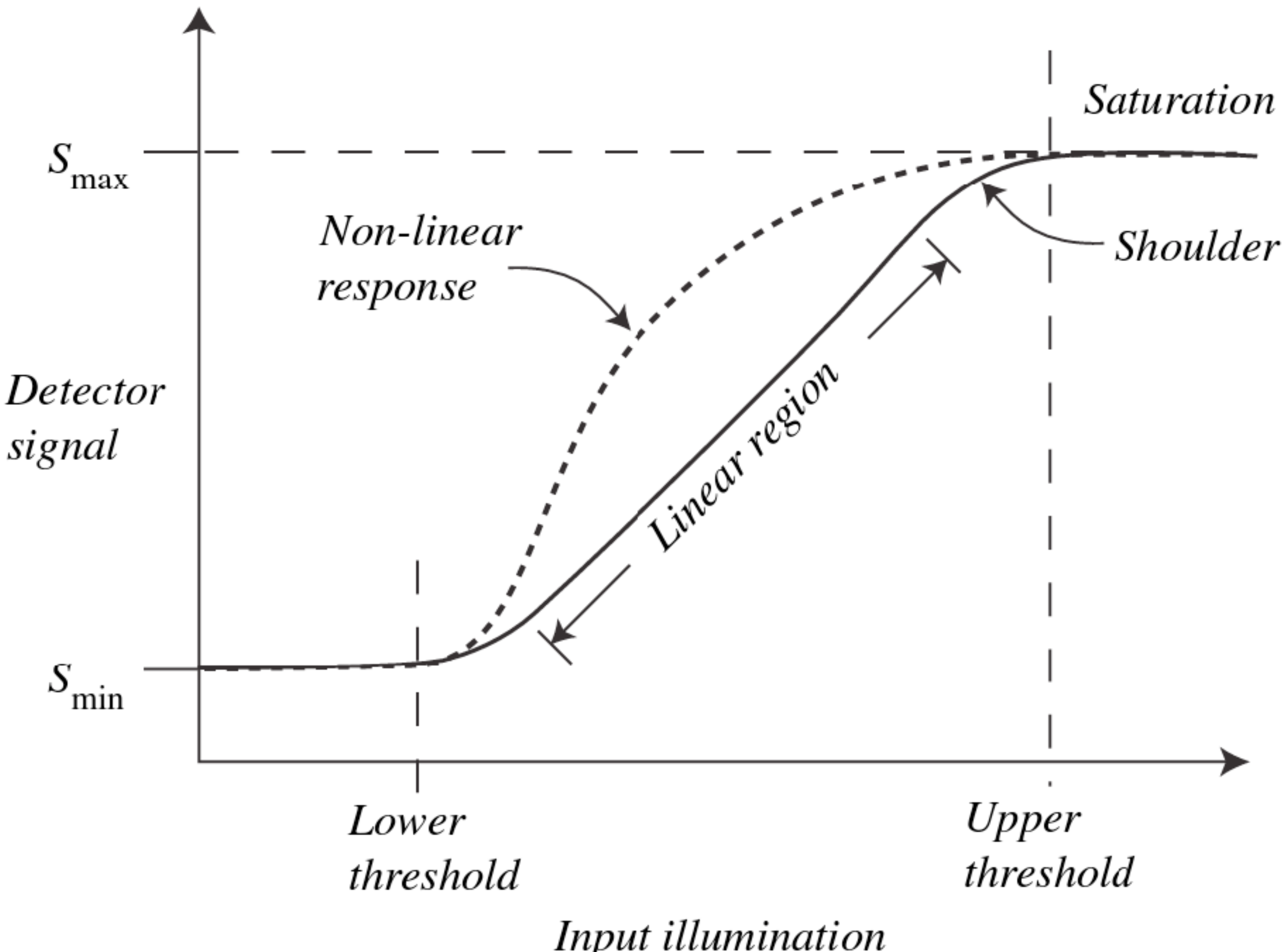
- The **perfect detector** is one with:
  - **100% efficiency** in converting photons to an induced signal.
  - If **zero** photon illumination, then **zero** detector output is obtained.
  - **Repeated measurements** of the same constant light source give the **same average value** within the ideal noise level.
  - **All** detector elements (e.g. pixels) have **exactly the same characteristics**.

  - **Sadly, does not exist, but:**
    - We can measure and understand our detector characteristics as part of the data model and correct our observed data for these instrumental effects.

Key detector properties

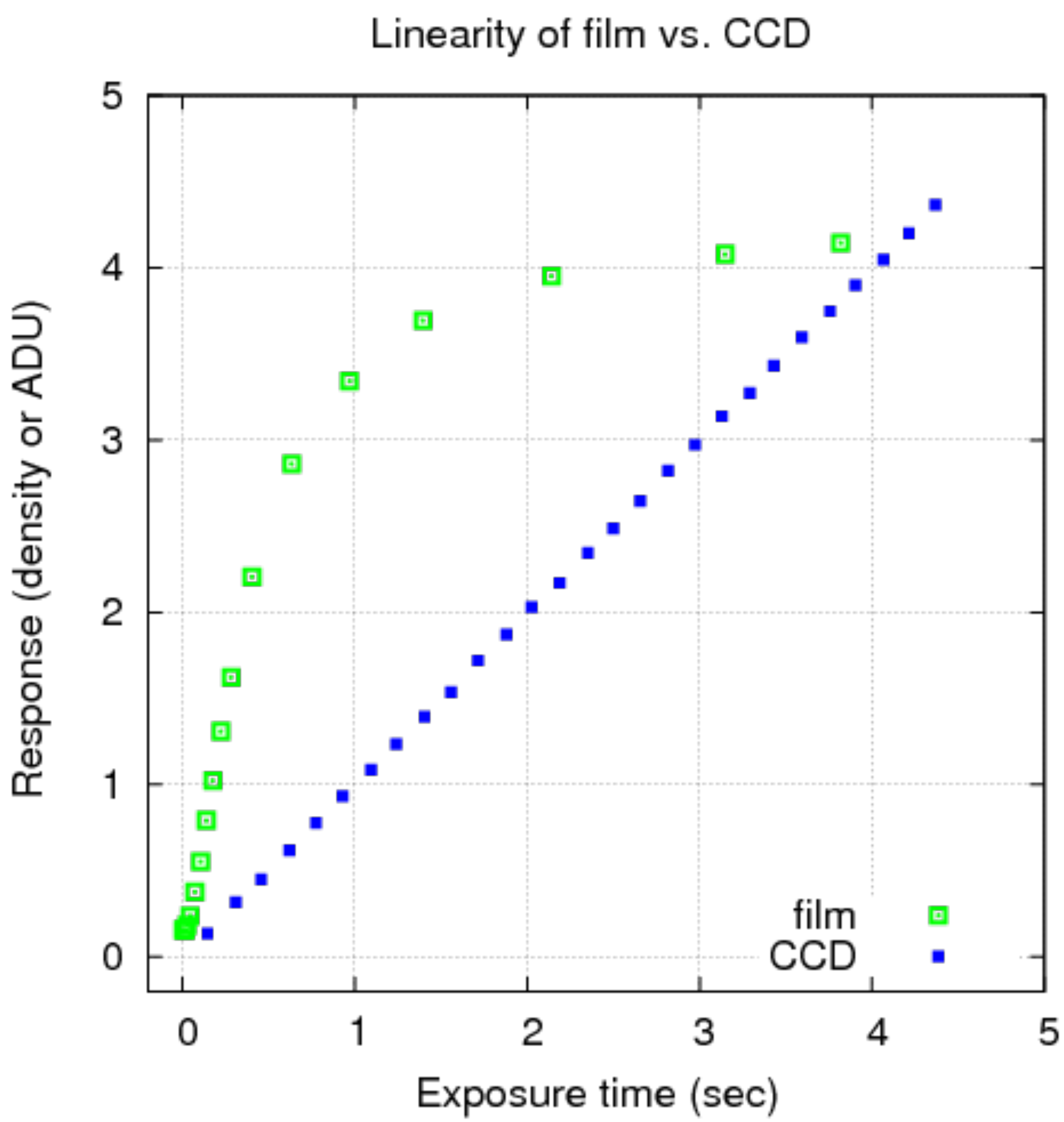


Detector property	Description
Quantum efficiency	The ratio of the number of produced photoelectrons to the number of incident photons.
Noise	The irreducible electronic fluctuations in the output signal from a detector.
Dynamic range	The ratio of the maximum detectable signal to the noise level.
Linearity	The extent to which the input signal is proportional to the incident photon rate.
Spectral response	The wavelength (or frequency) range over which the detector is sensitive.
Temporal response	The time interval over which the device cannot respond to another incoming photon.
Dark current	The output signal when not illuminated by light.



Saturation
S_{max}
Non-linear
response
Shoulder
Detector
signal
Linear region
S_{min}
Lower
threshold
Upper
threshold
Input illumination

Linearity



Linearity of film vs. CCD
Response (density or ADU)
Exposure time (sec)
0
1
2
3
4
5
0
1
2
3
4
5
film
CCD

Orion Nebula
1880 Henry Draper
51 min exposure on an 11” refractor

first photograph of Orion Nebula !

Orion Nebula
1883 Andrew Common
60 min on 36” reflector

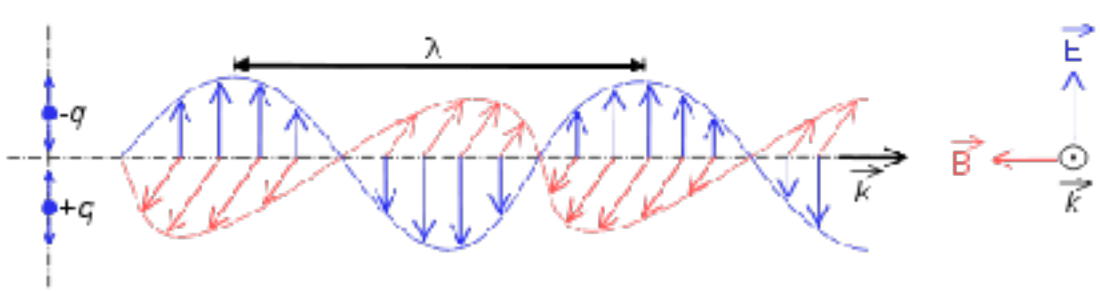
for the first time showed stars too faint to be seen by eye …

Orion Nebula
Palomar Hale 200 inch Telescope 1959

Orion Nebula
Hubble Space Telescope

Detector categories and types

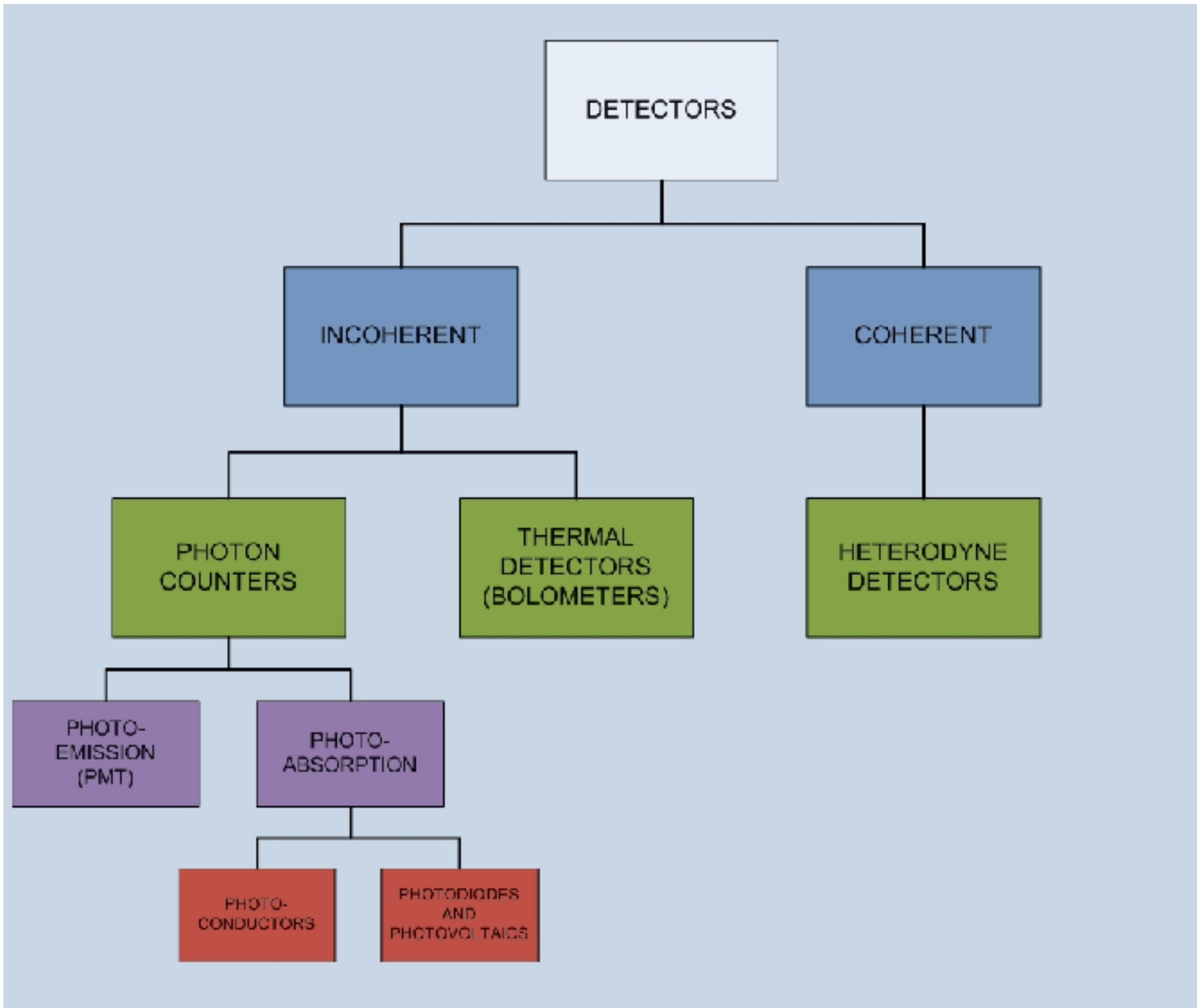
- **<u>Imaging or non-imaging:</u>**
  - Depending on whether detector measures direction-dependent radiation properties; often related to single- or multi-pixel nature.
- **<u>Detection method:</u>**
  - Photon detectors:
    - Photons release one or more electrons (or other charge carriers) on interacting with the detector material
  - Thermal detectors:
    - Photon energy heats the detector material, resulting in a change to a measurable property of the device, such as its electrical conductivity.
  - Coherent detectors:
    - Electric field of the wave is sensed directly and phase information Ψ can be preserved. All detectors that cannot measure phase are incoherent detectors.



$$\vec{E}(\vec{r},t) = \vec{E}_0 e^{j(\vec{k}.\vec{r}-wt)}\ \hat{n}$$

$$= \vec{E}_0 e^{j\Psi(\omega,\vec{k}.\vec{r},t)}\ \hat{n}$$

Families of detectors



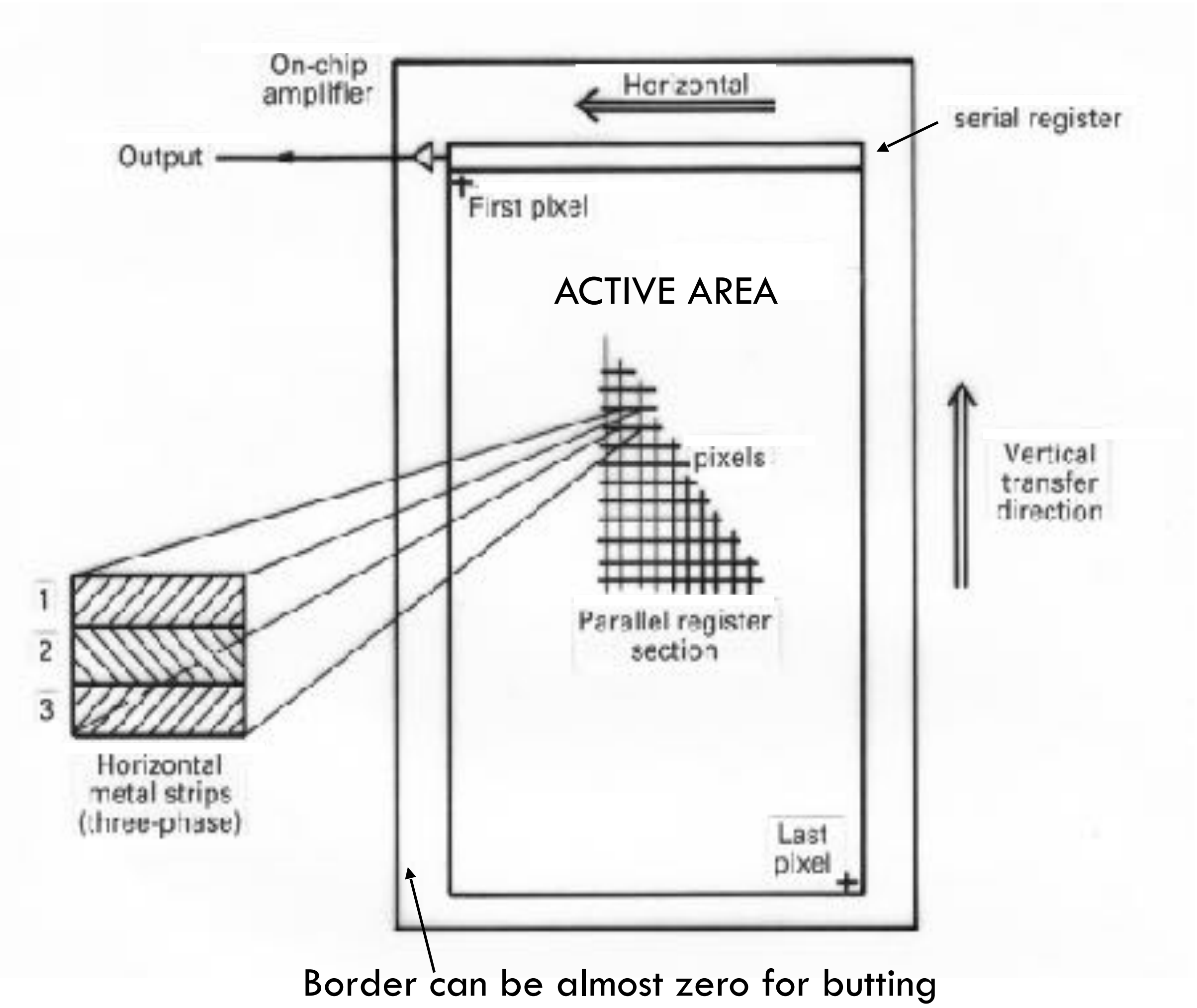
Invention and development of the CCD

- Invented in 1969 by Willard Boyle and George Smith at the Bell Laboratories (NJ)
- Image-forming devices of 100x100 pixels were not introduced until 1973
- Boyle and Smith received the basic patent at the end of 1974 and the Nobel Prize in 2009.
- CCD development was influenced by astronomical adoption and use.

Nobel Prize (1921) " The Nobel Prize in Physics 2009 was divided, one half awarded to Charles K. Kao *"for groundbreaking achievements concerning the transmission of light in fibers for optical communication"*, the other half jointly to Willard S. Boyle and George E. Smith *"for the invention of an imaging semiconductor circuit – the CCD sensor"*.

Basic principles

- A CCD is an array of pixels, each of which can absorb photons of light and release a photoelectron by the internal photoelectric effect.
- Don't want the photoelectrons to migrate away from the site of impact of the photons.
- Photoelectrons confined within a pixel by a special electrostatic field.
- Need a storage region at each pixel capable of holding many photo-generated charges.



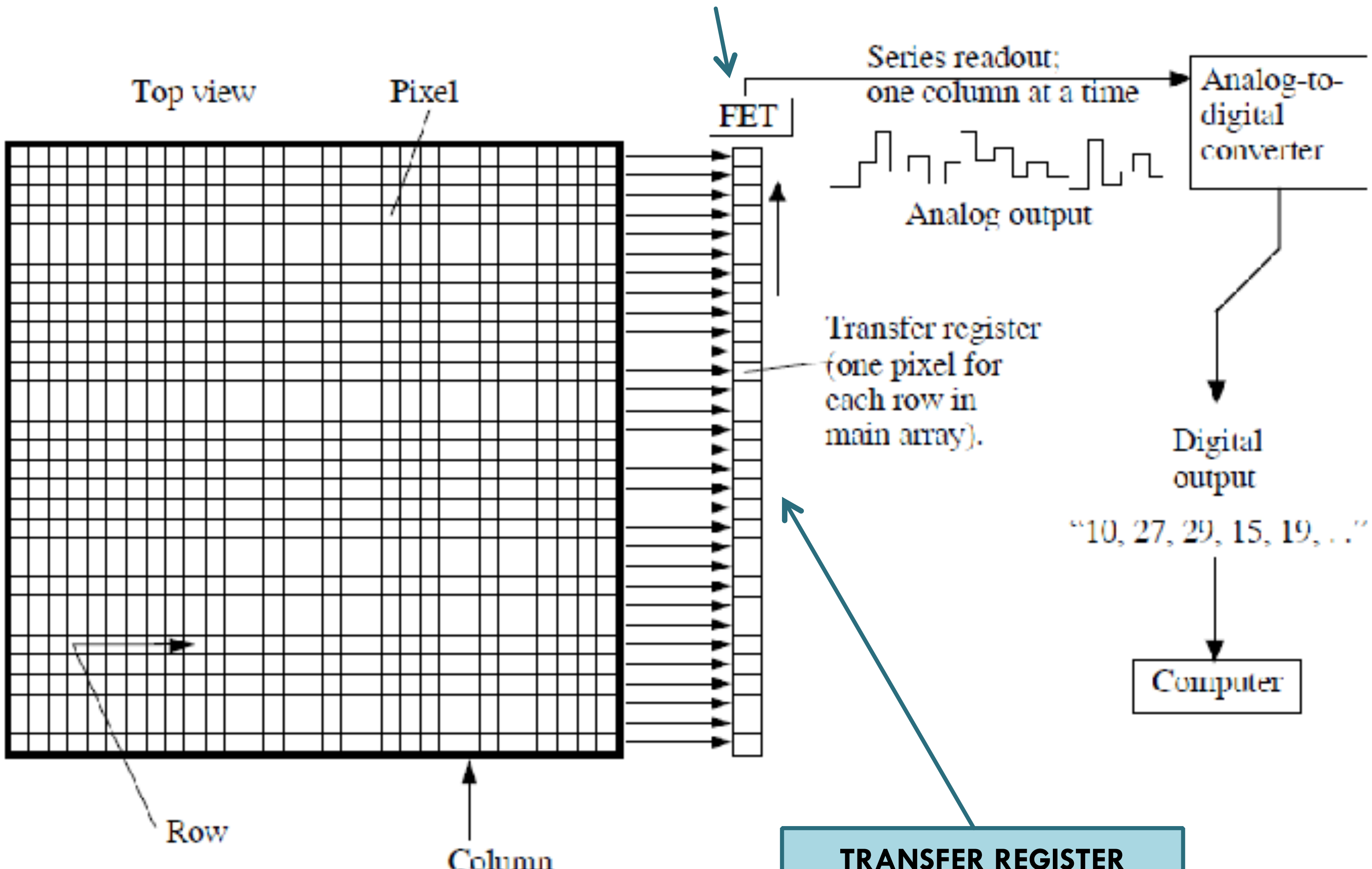
CCD readout via transfer register

FIELD EFFECT TRANSISTOR (FET): Amplifier to produce analog signal proportional to stored charge

RATE: Typically 50 kHz (50,000 pixels/s). Astronomical CCDs are ***slow-scan,*** and must be cooled.

CHARGE TRANSFER EFFICIENCY (CTE): Net loss factor in transferring stored charge.

BURIED CHANNEL TRANSFER: Transfer via a buried semiconductor channel as opposed to surface transfer. Minimizes ***electron trapping*** at impurities.



TRANSFER REGISTER

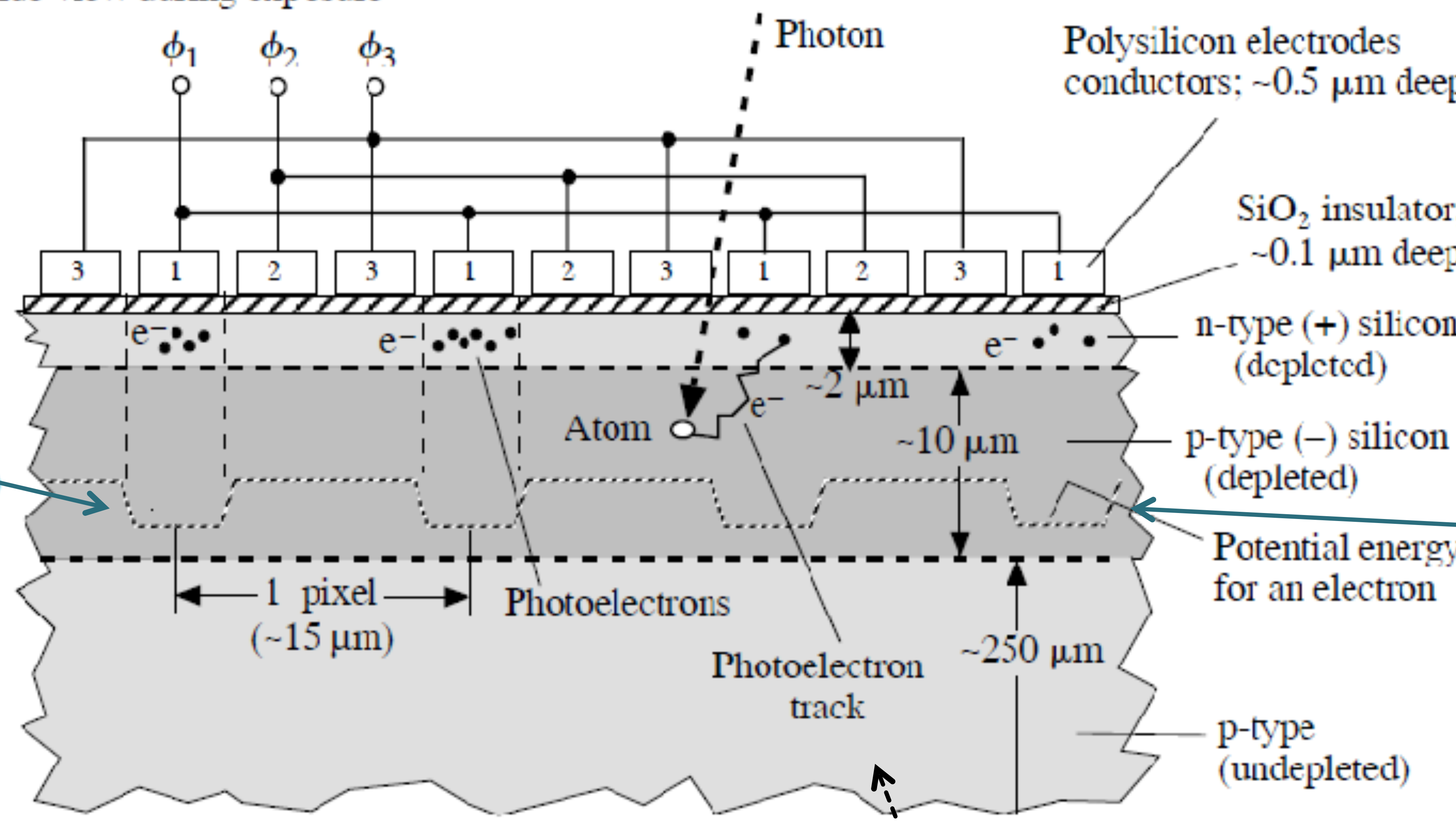
(Bradt 2004)

Structure of a CCD

COATING: May be coated for down-conversion, or anti-reflection.

FRONT-SIDE ILLUMINATION: Some loss in QE due to gates.

(a) Side view during exposure



ELECTROSTATIC POTENTIAL WELL: Depth controlled by voltages on Φ_1, Φ_2, and Φ_3

DEPLETION REGION

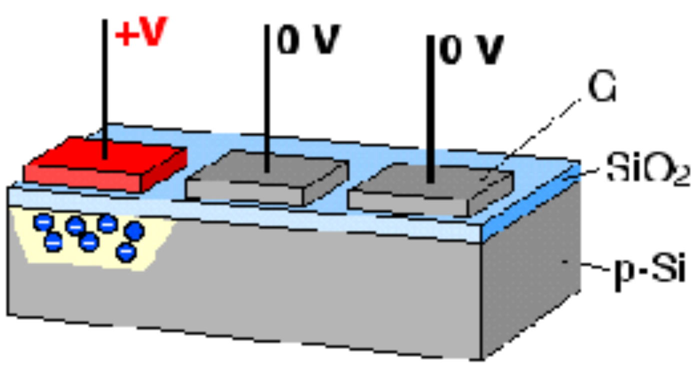
(Bradt 2004)

DARK CURRENT: Electrons promoted to conduction band by internal ***thermal*** excitation.

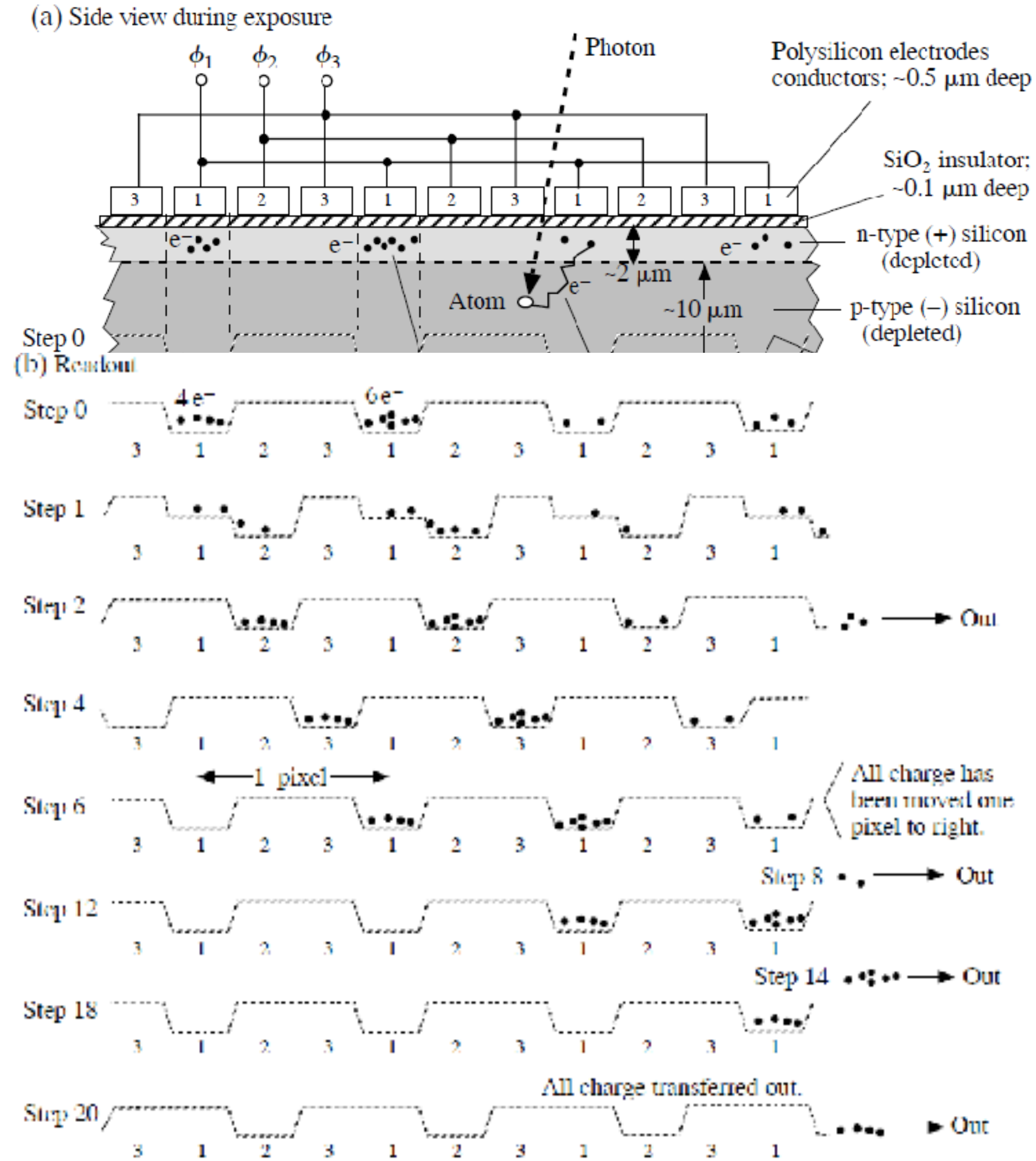
THINNED CCDs: Back-side illumination; higher QE.

Charge-coupling: a major innovation

- Can migrate stored electrons at each pixel by applying a systematic (clocked) voltage shift from pixel to pixel.
  - ***Charge-coupling***
  - Allows CCD readout
  - Multiple readouts can be used to clear the n-type electrons before each exposure.

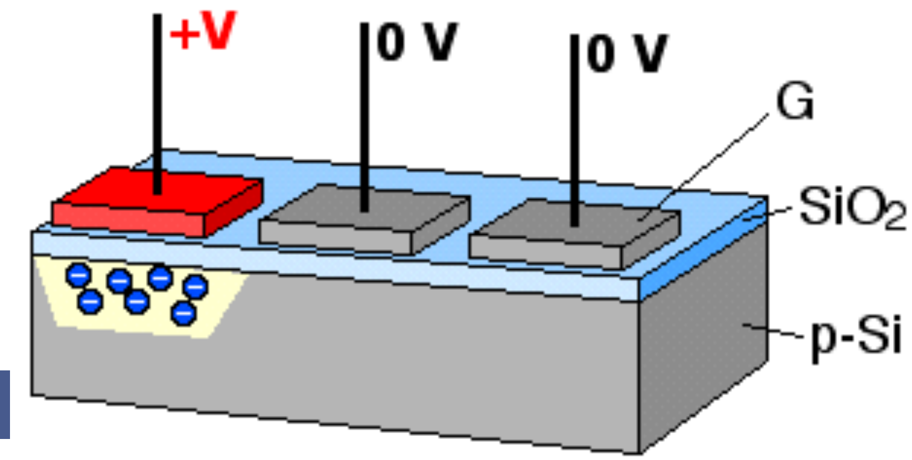


(CC)

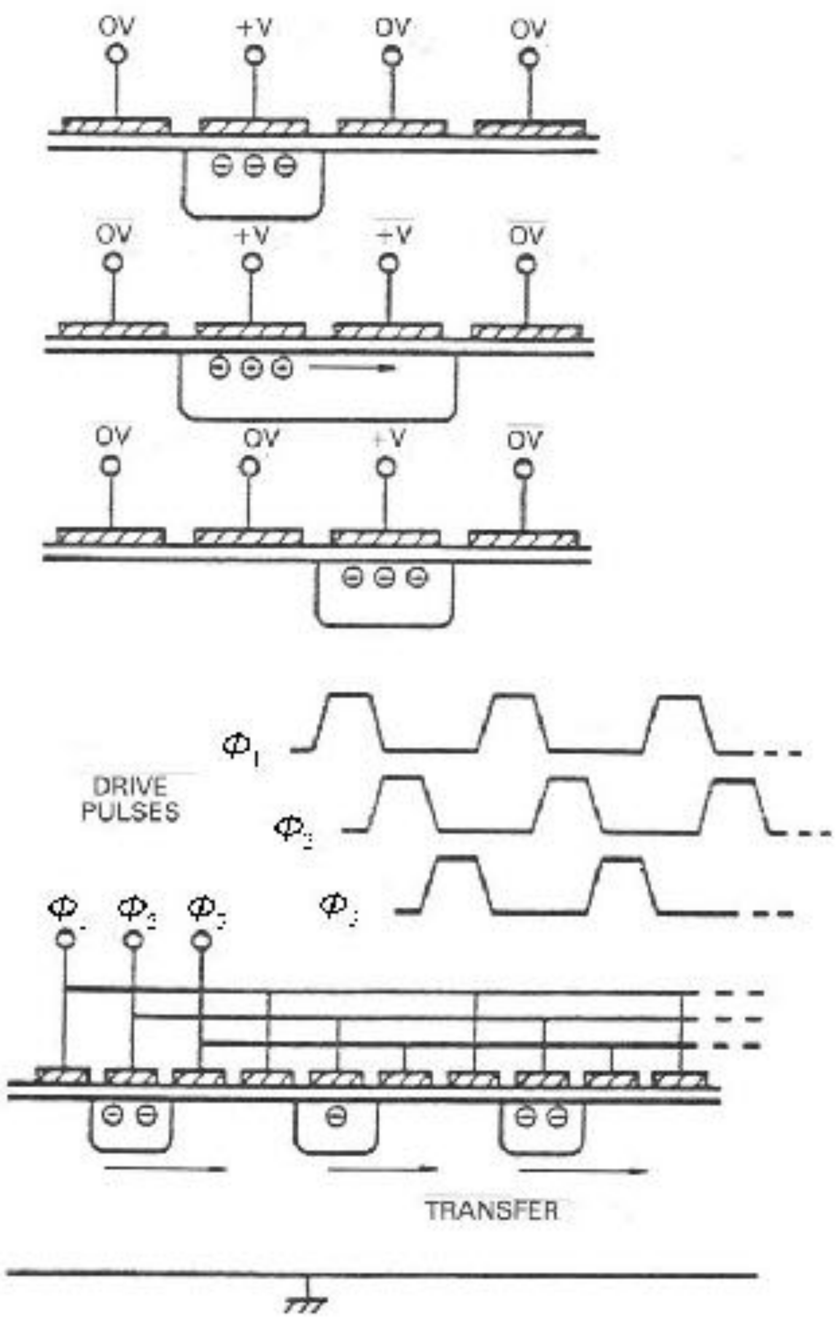


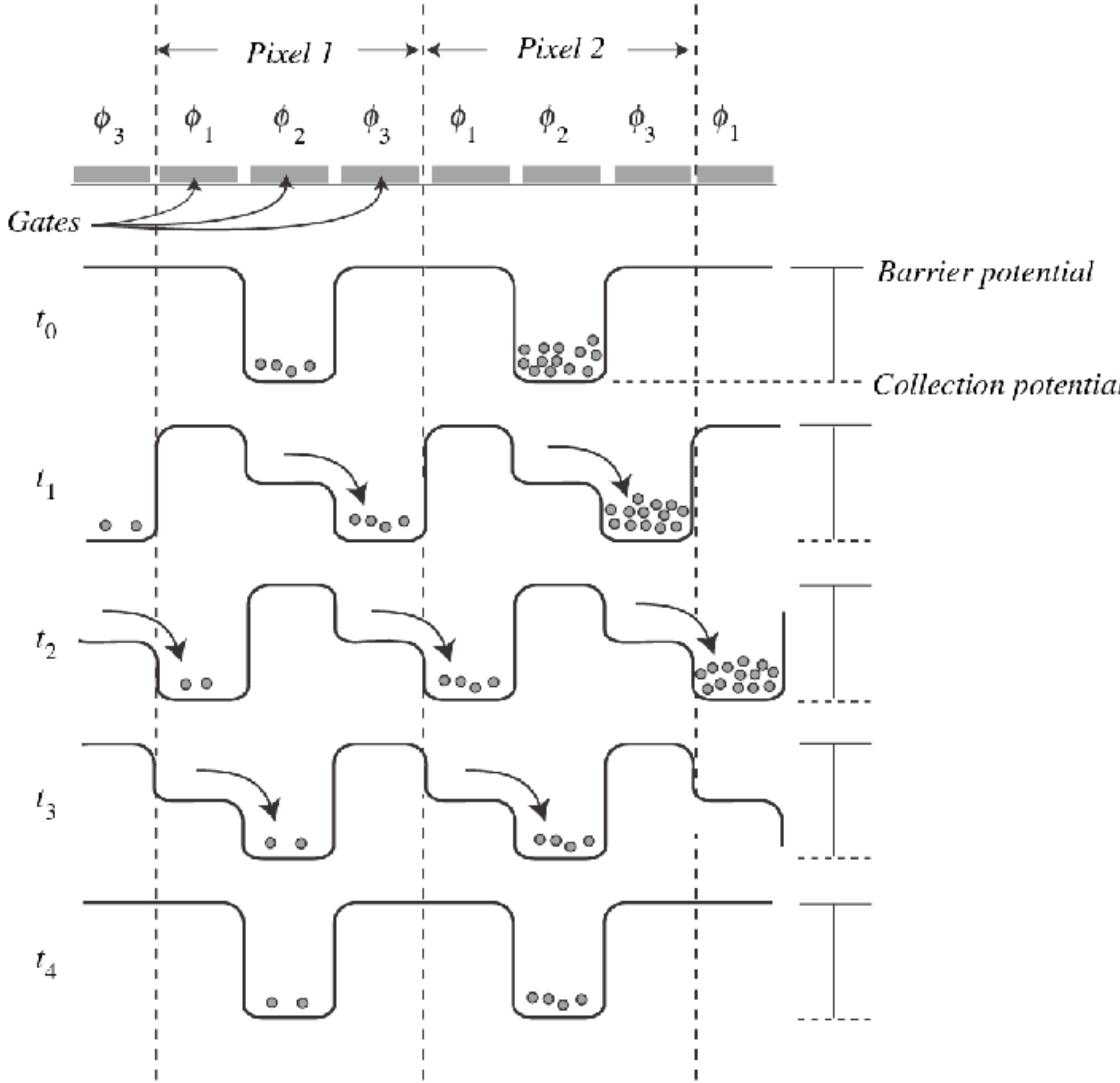
(Bradt 2004)

Charge-coupling

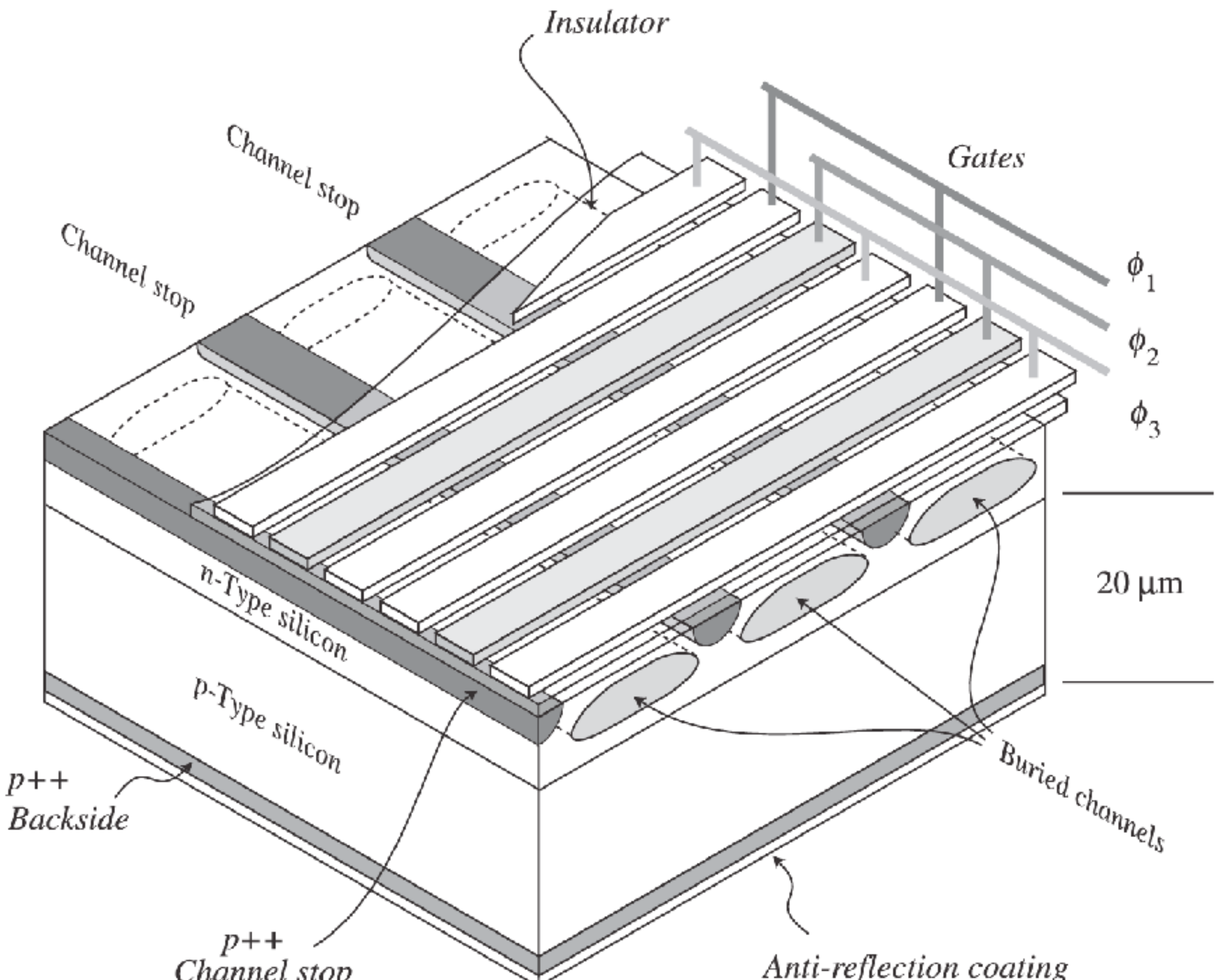


- To transfer charge to an adjacent electrode, raise the voltage on the adjacent electrode to the same value as the first one.
- By connecting sets of electrodes together, charge stored on the 2-d imaging area can be moved simultaneously in that direction.
- When the voltage on the original electrode is reduced to zero the transfer is complete.
- Because it takes 3 electrodes to define 1 pixel, three of the above transfers are required to move the 2-d charge pattern by one whole pixel along the direction at right angles to the electrode strips.
- The process of raising and lowering the voltage is known as **clocking**.
- The drive or clock pulses can be described in a diagram called a **timing waveform** (Fig. 7.7).





Pixel 1
Pixel 2
ϕ_3 ϕ_1 ϕ_2 ϕ_3 ϕ_1 ϕ_2 ϕ_3 ϕ_1
Gates
Barrier potential
Collection potential
t_0
t_1
t_2
t_3
t_4


Insulator
Channel stop
Channel stop
Gates
ϕ_1
ϕ_2
ϕ_3
20 μm
n-Type silicon
p-Type silicon
Buried channels
p++
Backside
p++
Channel stop
Anti-reflection coating

Early astronomical adoption

- From 1973 to 1979 Texas Instruments (TI) developed CCDs of 100 × 160 and 400 × 400 pixels, then 500 × 500 pixels and finally 800 × 800 pixels.
- **1974:** Jim Janesick attached a 100x100 Fairchild CCD to his small 20.3 cm (8-in) telescope and imaged the Moon (Janesick 2001).
- **1976**: Smith and Janesick obtained the first astronomical imagery with a charge-coupled device on a professional telescope; the 61-inch telescope on Mt. Bigelow near Tucson (Arizona).
- CCD images of Jupiter and Saturn were obtained using a methane filter.
- Brad realized they were observing "limb-brightening" of Uranus in the methane band for the first time.
- **1981:** 5-20 astronomical groups were working on CCD imagers.
- Large, custom astronomy CCDs typically ordered directly from CCD foundries.

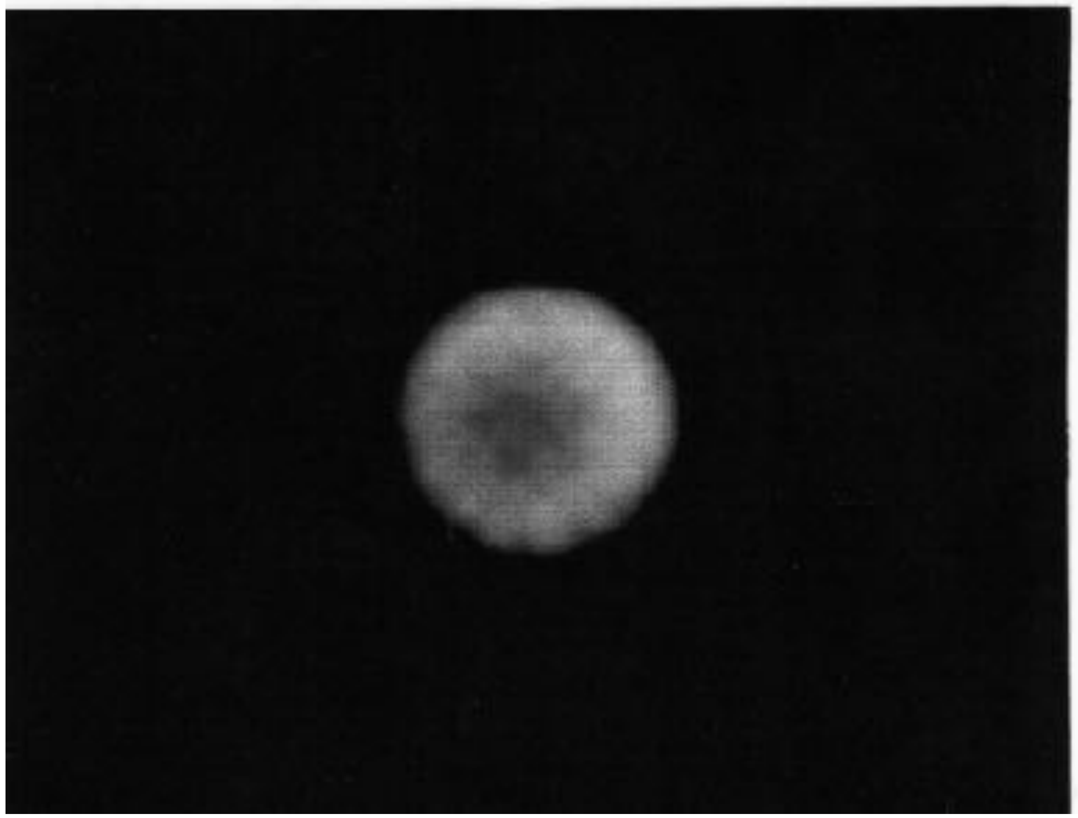


Figure 7.2: Uranus observed with one of the first CCD cameras in 1976.

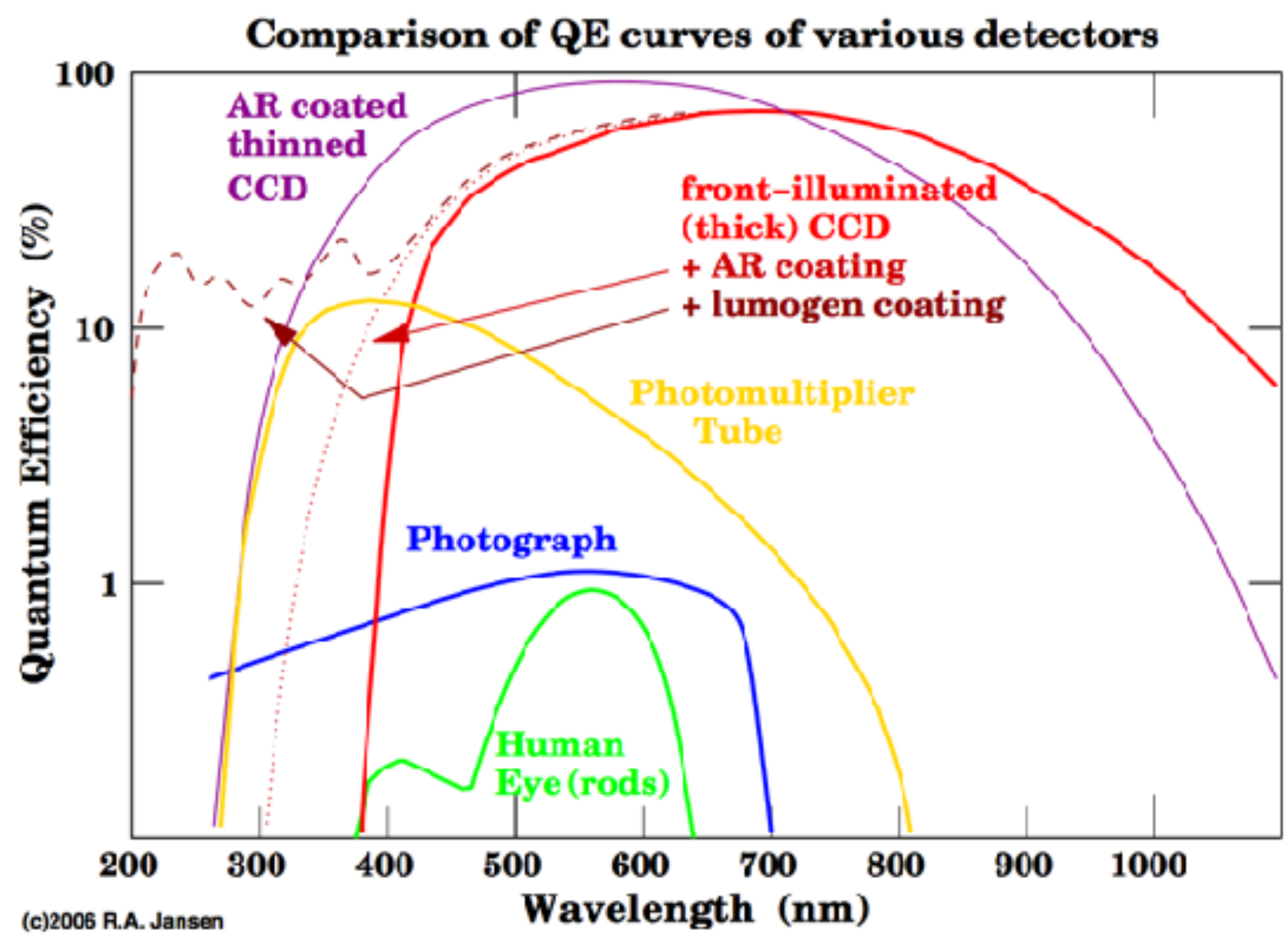
CCD performance

Pros:

- Can be designed for UV/O/IR, even X-ray bands.
- Very high quantum efficiency relative to other detectors.
- Highly linear over dynamic ranges of 10^5:1.
- Digital !

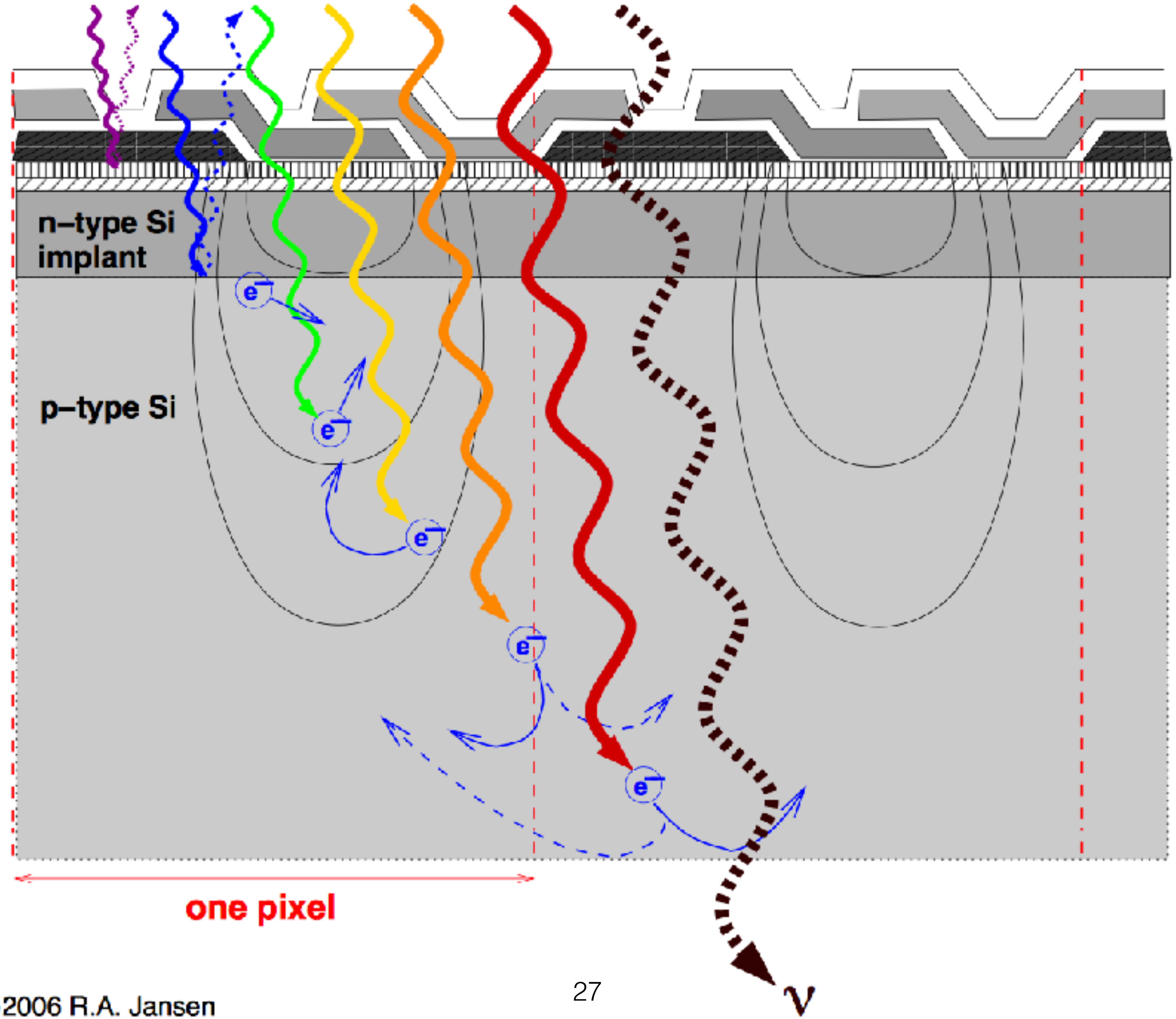
Cons:

- CCD charge storage can be contaminated by cosmic ray ionization events – can be reduced by ***anti-blooming*** measures.
- Electron storage does eventually saturate.
- Readout time; noise added by amplification and digitization.
- Each pixel does not have perfectly uniform response.
- Pixels can be <15 μm in size, but still larger than fine grain photographic film.
- Filling large areas was difficult. Only now can arrays of CCDs match the photographic plate areas of the 1950s.



The dominant imaging detector in astronomy today.

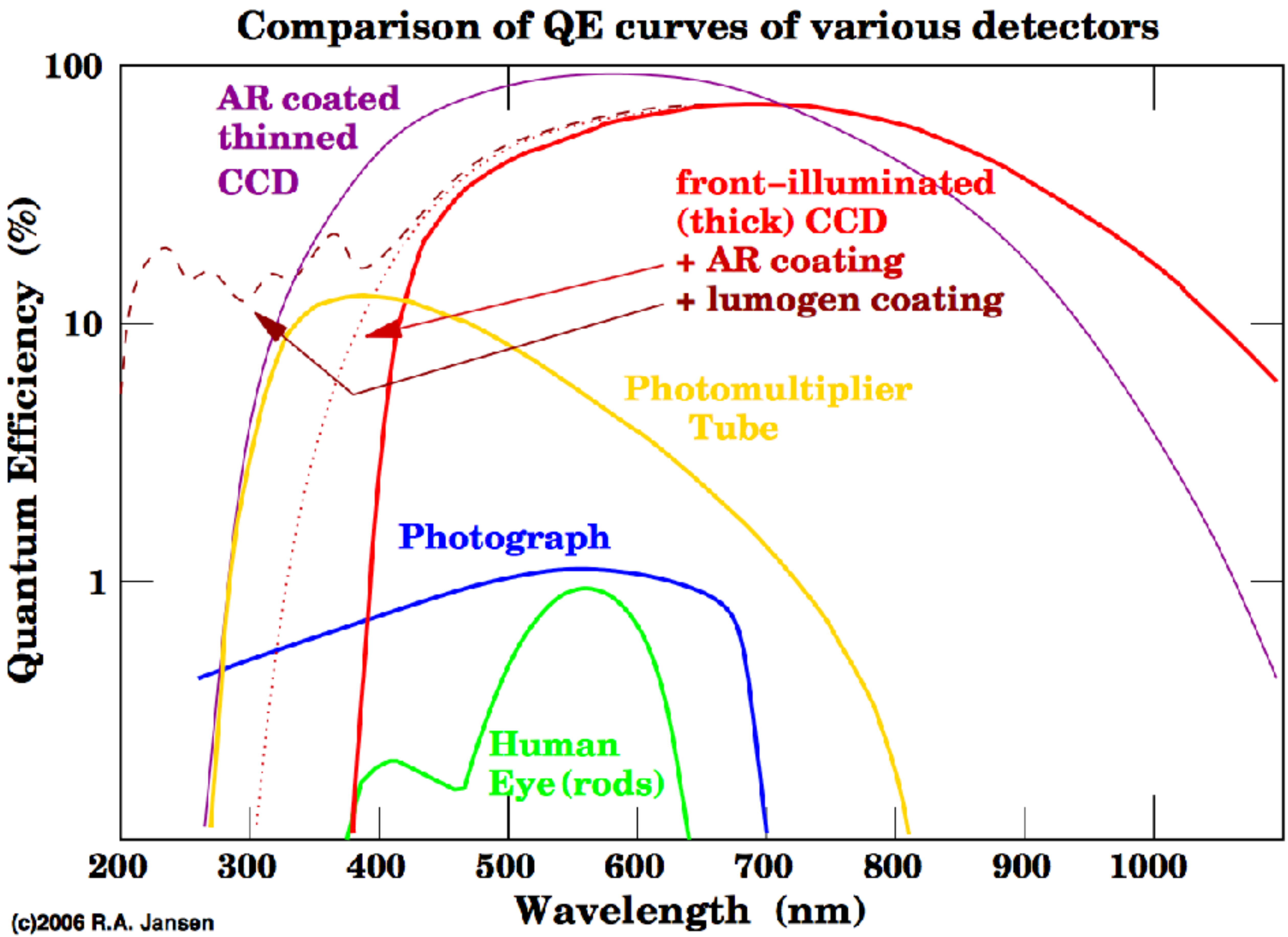
Wavelength–dependent absorption depth



(c)2006 R.A. Jansen

Back-illuminated CCDs

- In **front-side illumination**, photons are compelled to pass through the overlying electrode structure in order to reach the depletion region in the silicon.
  - In practice, this approach results in severe absorption of blue light in the electrodes.
- **Virtual phase** CCDs are one approach to solving this problem.
- Alternatively, the CCD can be turned over and illuminated from the back side!
- Before this becomes effective however, the thick silicon substrate must be reduced in thickness (either mechanically or chemically) to only 10 µm or so.
- **Thinned, backside-illuminated** CCDs have excellent response in the blue and violet.
- If thinned too much they lose red response because red photons need more absorption length and if it's not there, they will pass right through the silicon!

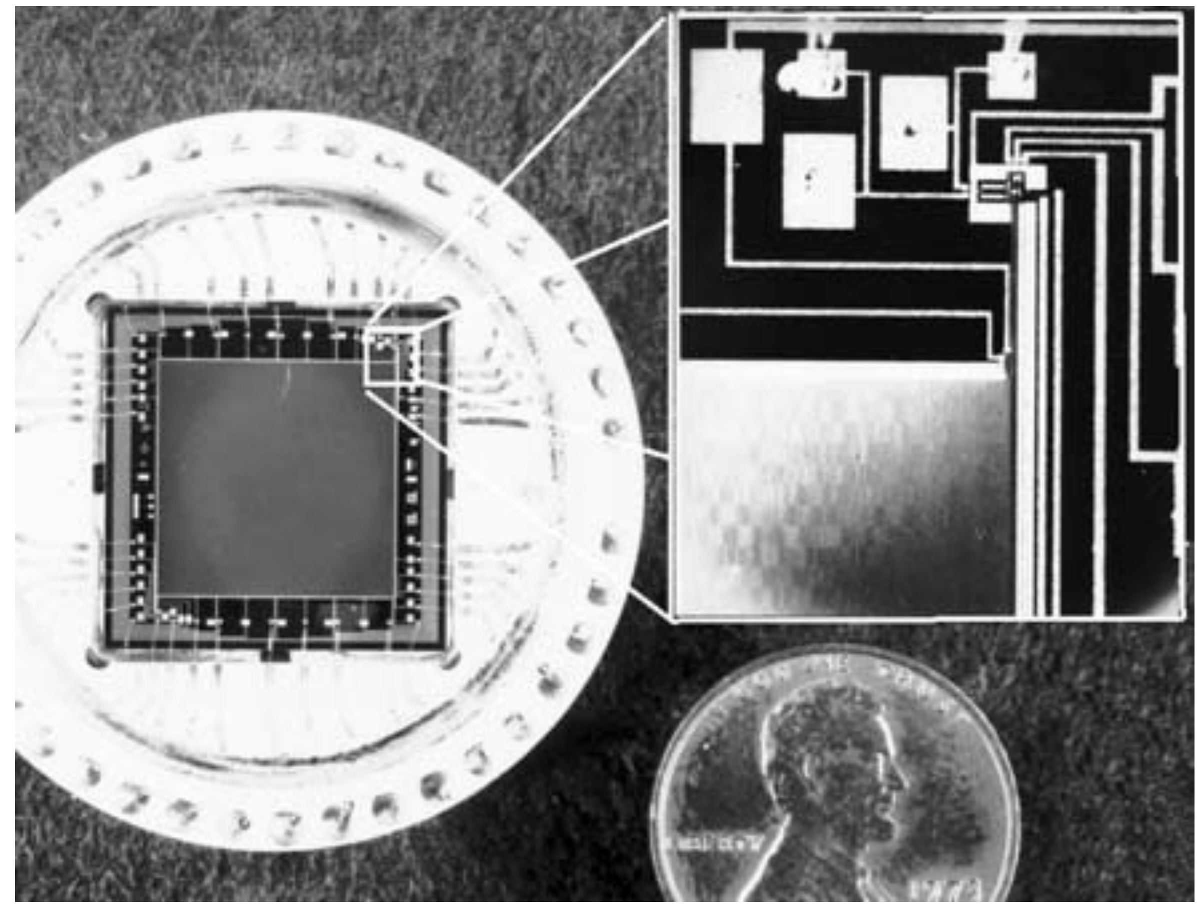


Comparison of QE curves of various detectors
100
10
1
Quantum Efficiency (%)
AR coated
thinned
CCD
front–illuminated
(thick) CCD
+ AR coating
+ lumogen coating
Photomultiplier
Tube
Photograph
Human
Eye (rods)
200
300
400
500
600
700
800
900
1000
Wavelength (nm)
(c)2006 R.A. Jansen

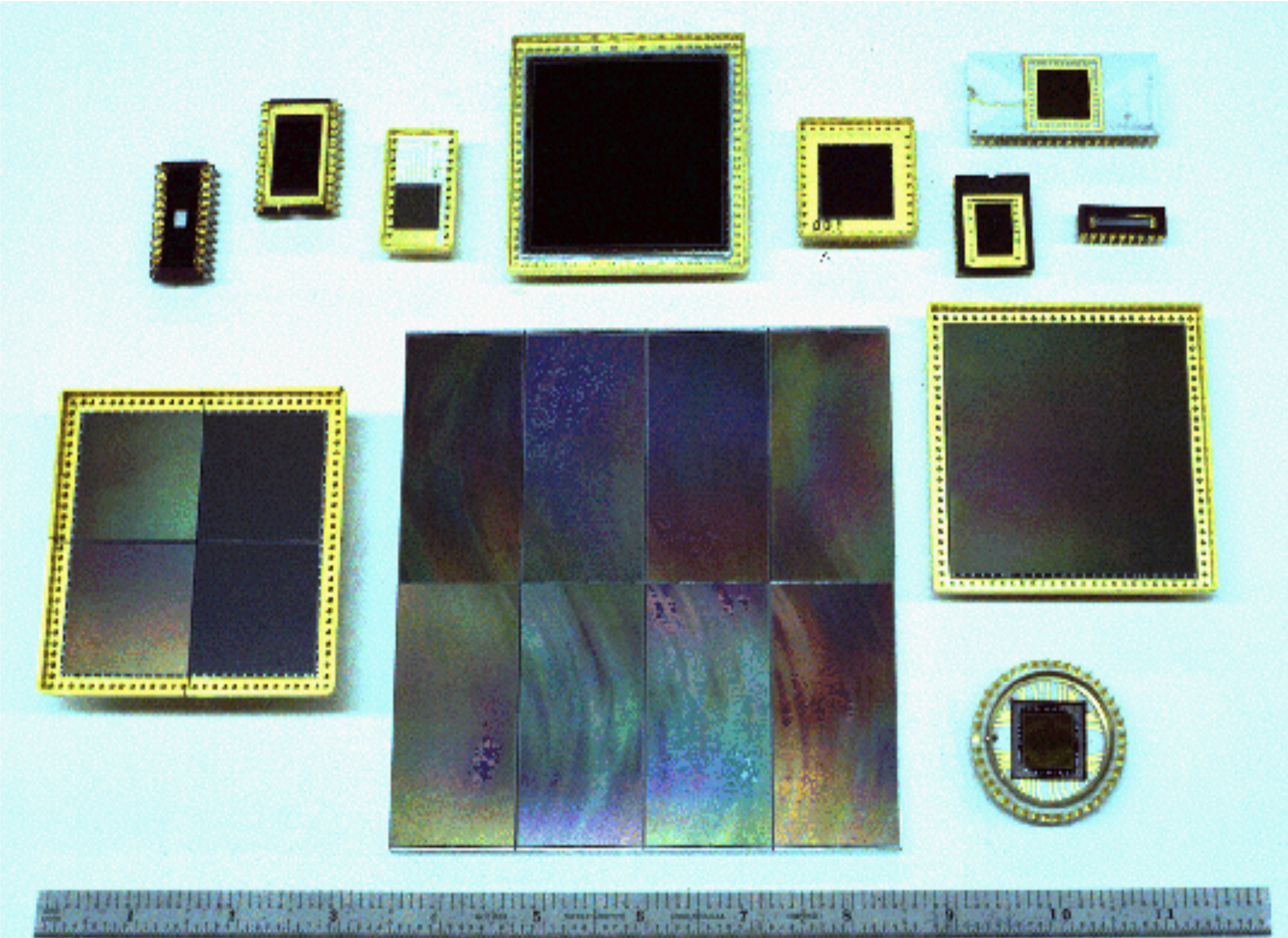
CCD for HST/WFPC

Texas Instruments
mid 1980s
800x800 pixels (<1MP)

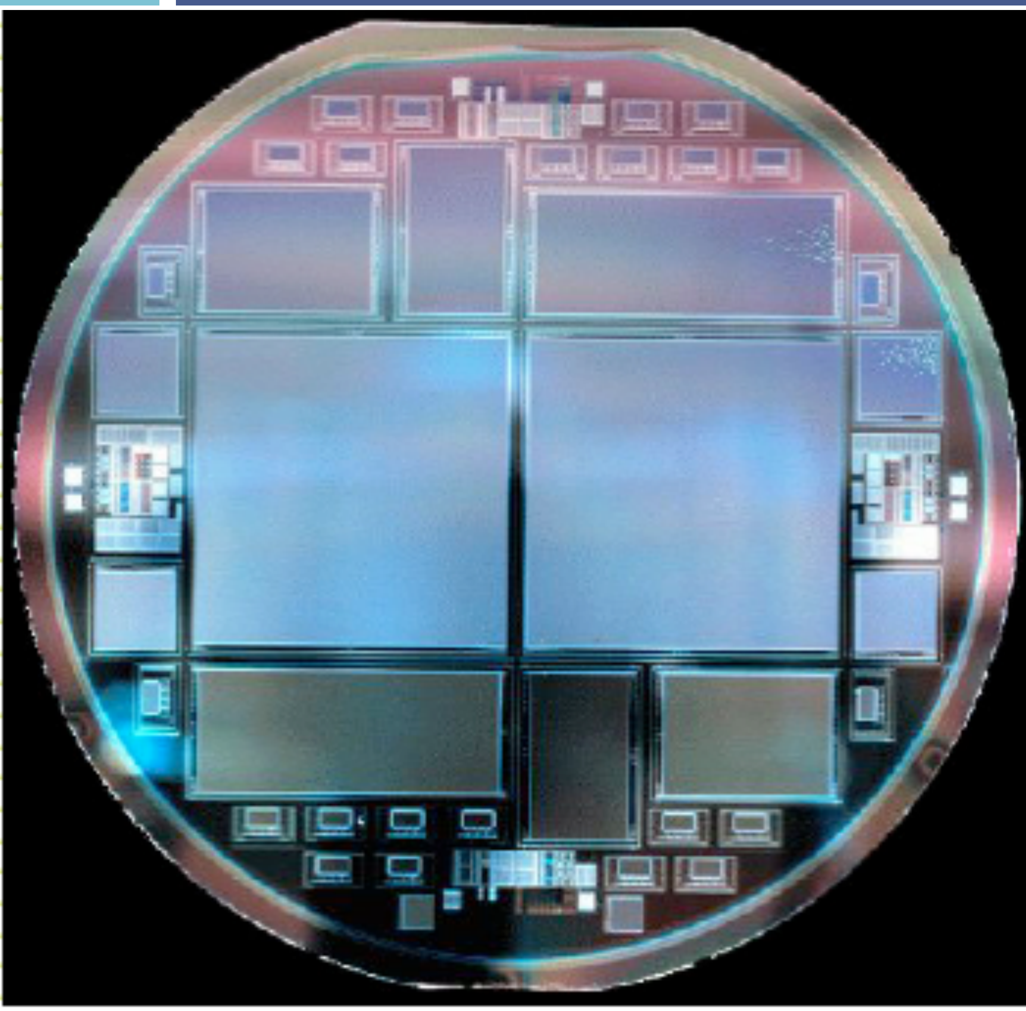
Image shows the front side, but this
detector was back-side illuminated



Historical collection of CCDs including some large format devices.



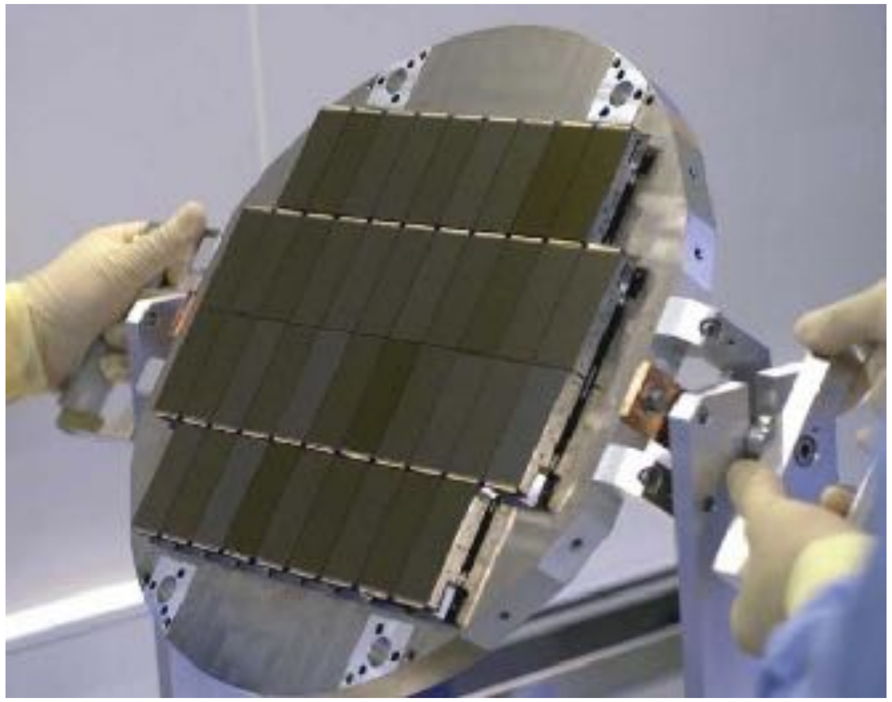
CCDs today

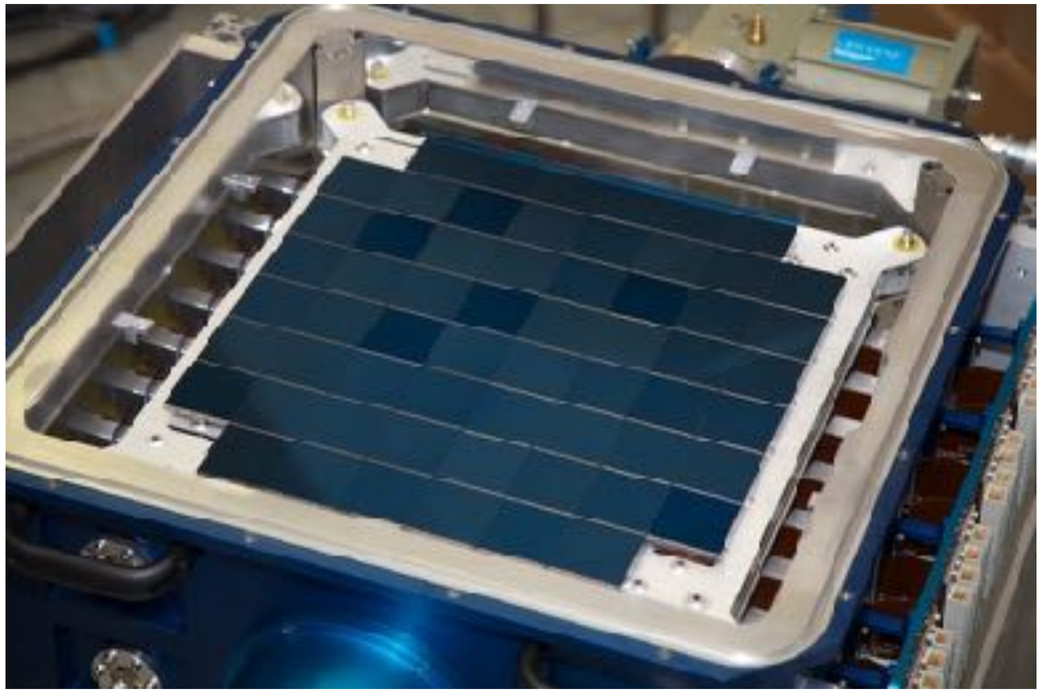


A huge wafer of silicon about 6 inches in diameter with multiple CCDs and other devices laid out on it. The large chips are 4k x 4k pixels

A ***mosaic*** of 40, 2048 x 4612 pixel CCDs yielding a camera with 378 megapixels on the CFHT this camera has a 1 degree field with 0.18” resolution (“buttable” CCDs).

CCD Mosaics



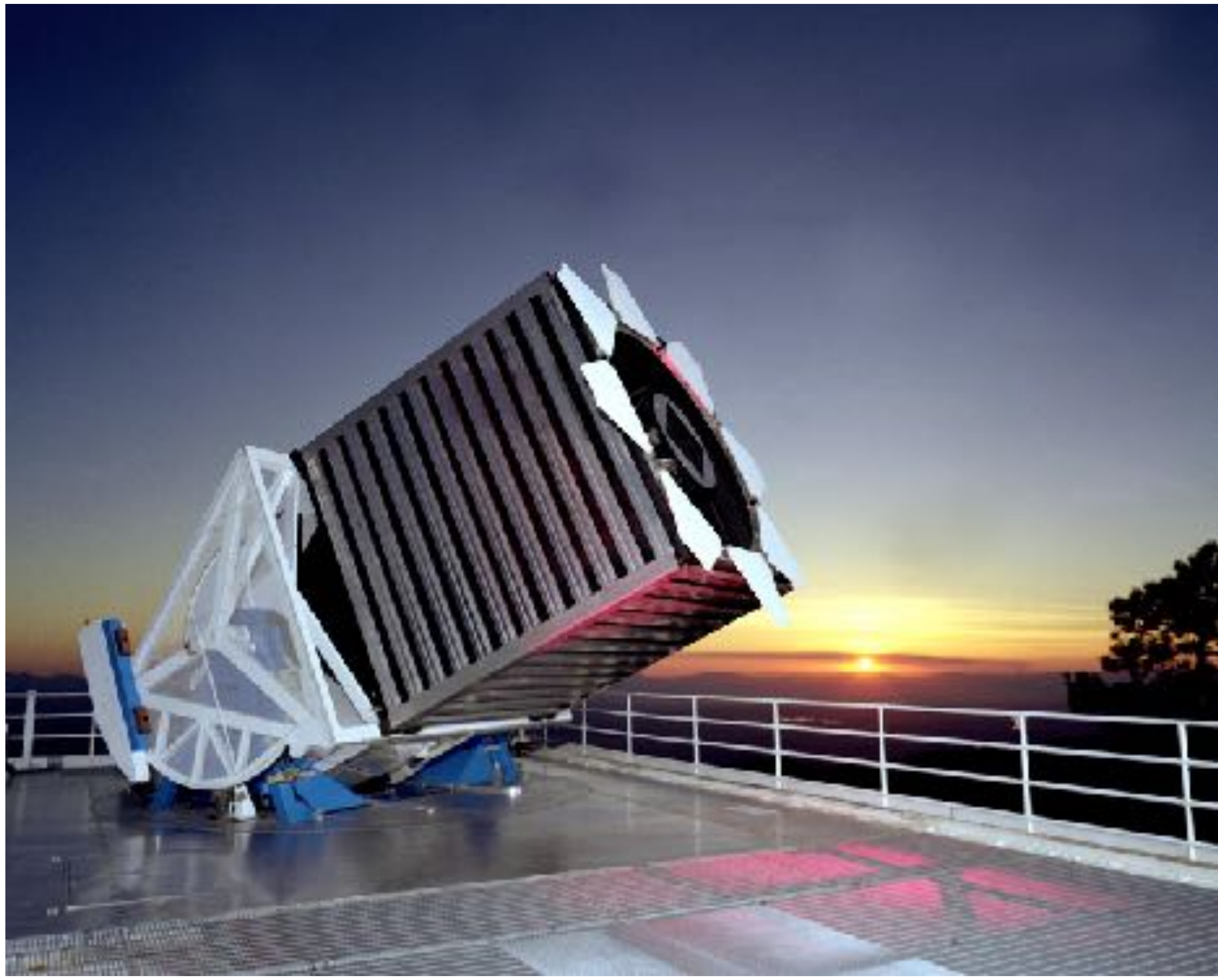


First operational wide-field CCD camera was MegaCam on the Canada-France-Hawaii Telescope (CFHT) in 2002. Using 40 2K x 4K CCDs from e2v, this camera's field of view = 0.9 square degrees (Fig. 7.14 – left above).

LSST will have 3 gigapixel camera; Pan-STARSS 1.4 gigapixel

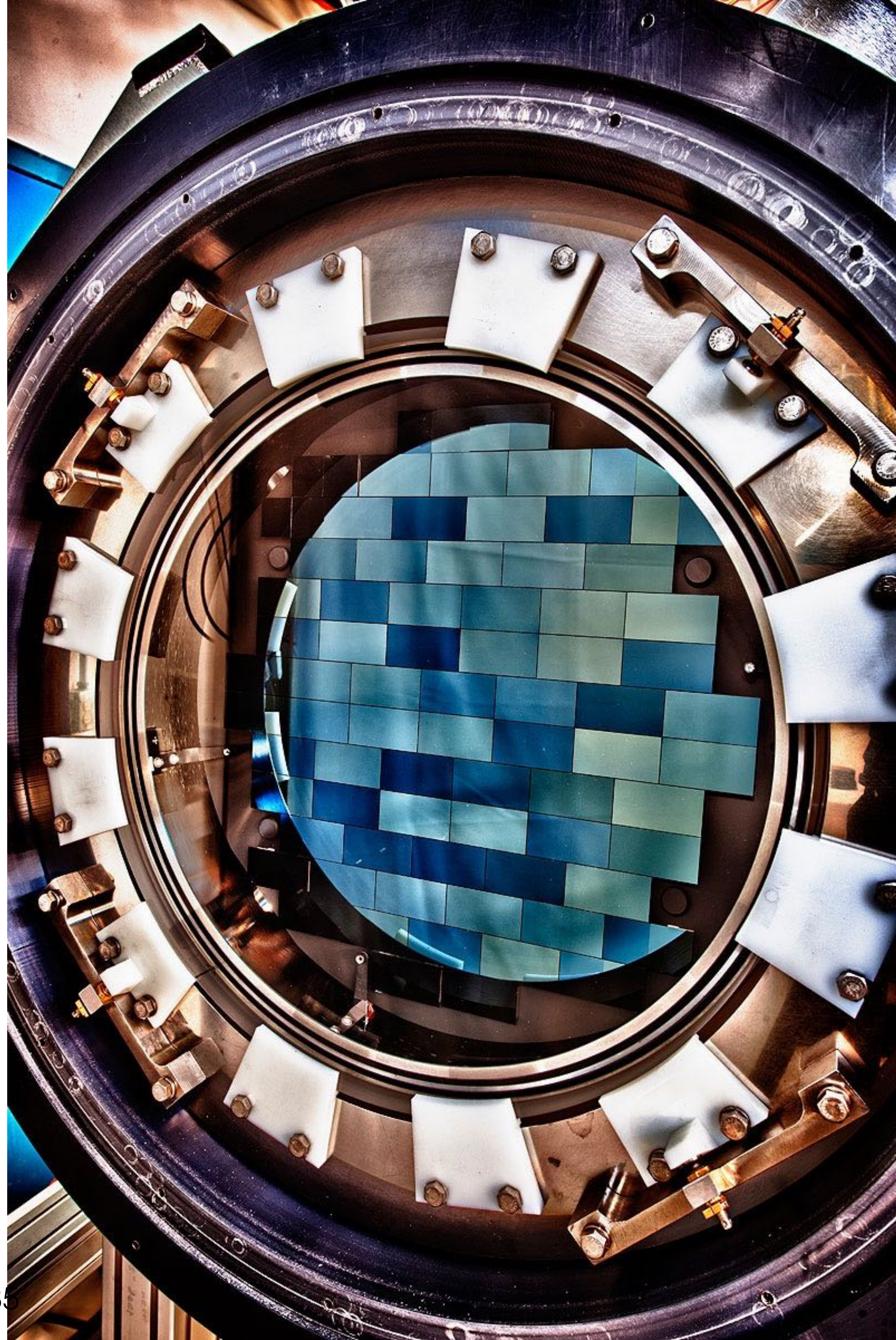
SDSS

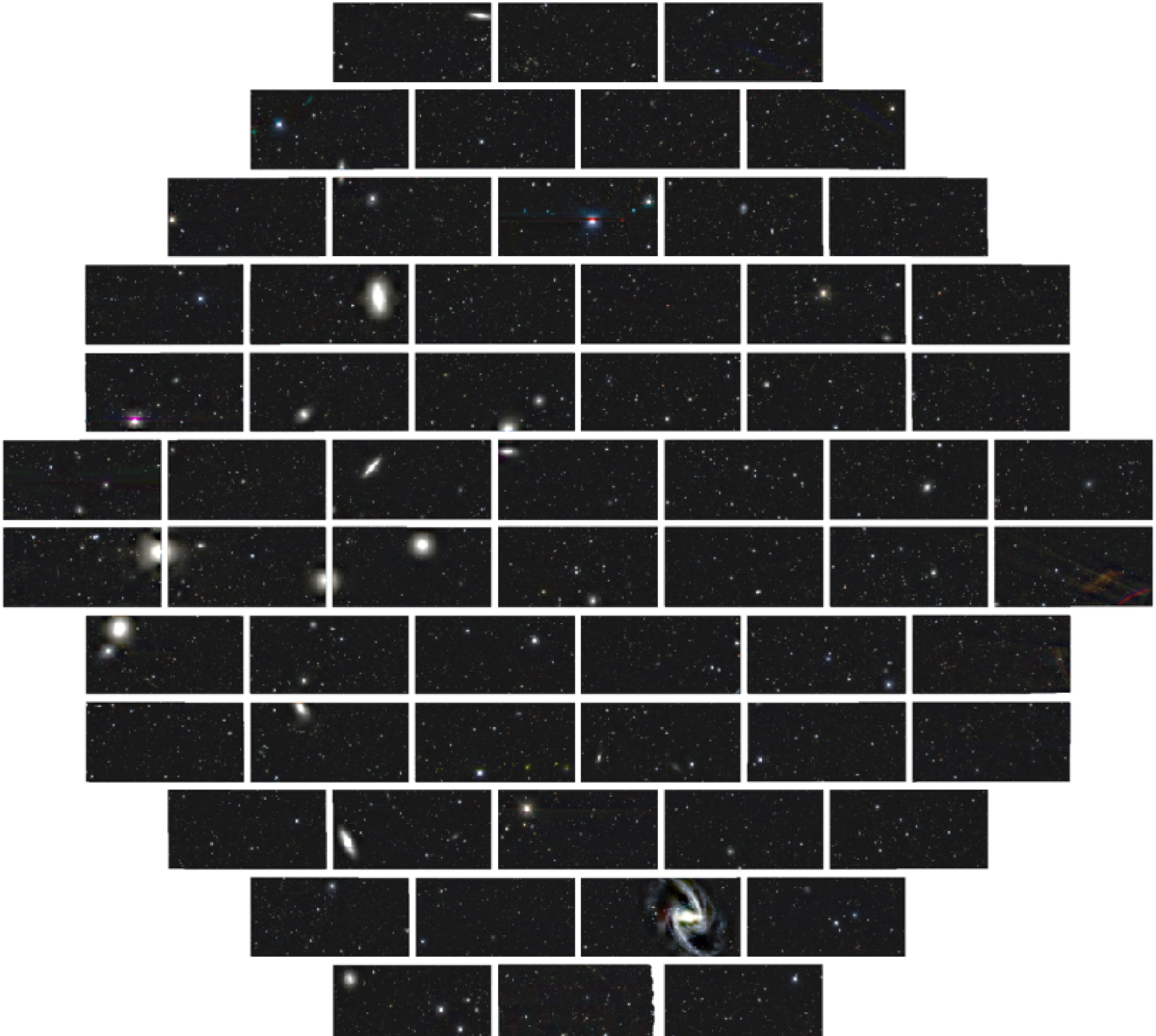
- Sloan Digital Sky Survey
- 2000-2009
- 2.5 m telescope at Apache Point, New Mexico
- 120 Mpix, 200GB every night
- *u,g,r,i,z* simultaneously in drift scan mode
- 35% of sky, photometry of 500M objects, spectroscopy for 3M



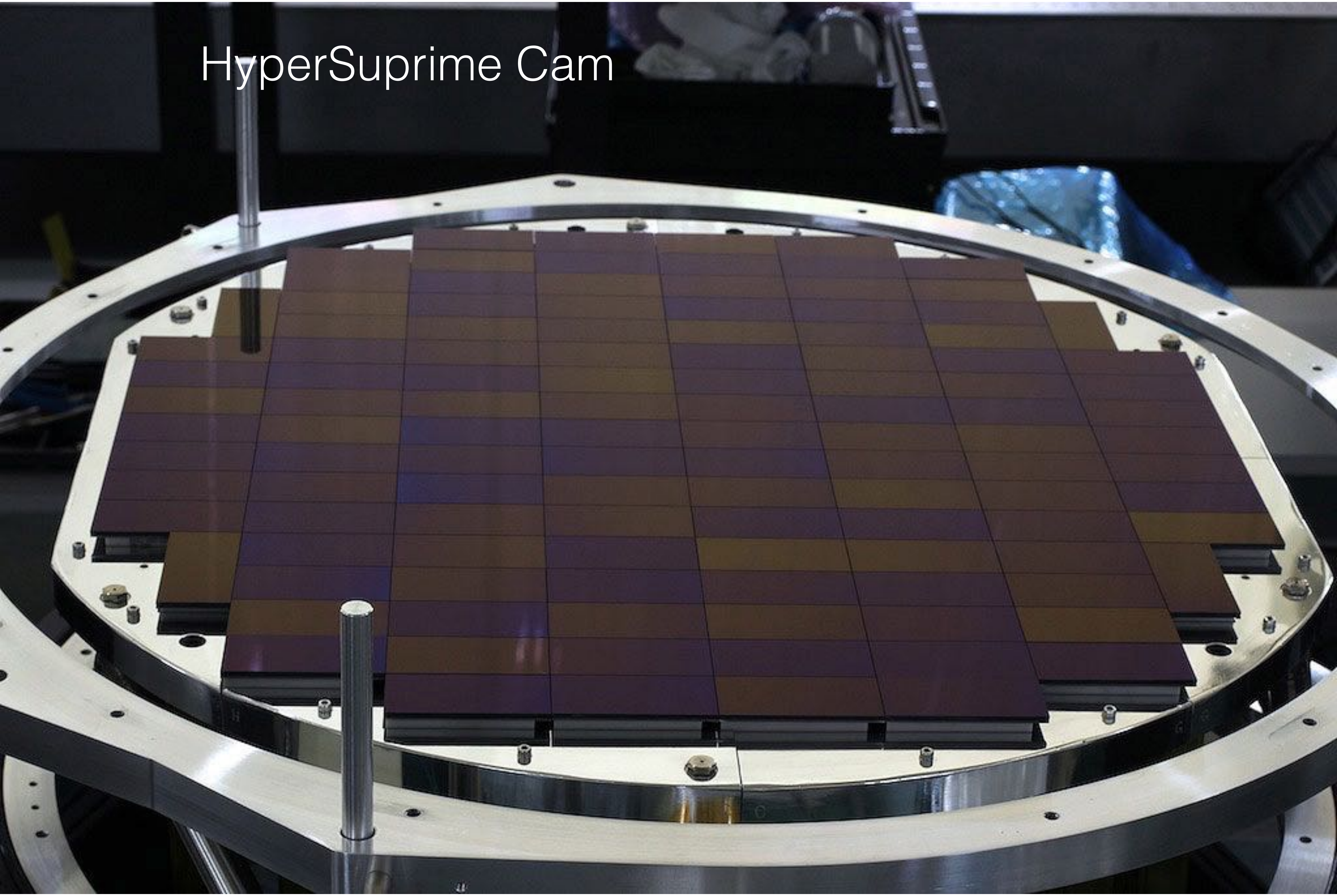
DES

- Dark Energy Survey
- 2013—2019
- 4m Blanco Telescope at CTIO in Chile
- 2.2 deg diameter, 3 deg^2 FOV
- 62 2048x4096 (8 Mpix) back-illuminated CCDs with 15 um pixels
- (for comparison, iPhone is ~5 Mpix with ~2um pixels)
- 570 Mpix
- 5000 deg2 survey in grizY

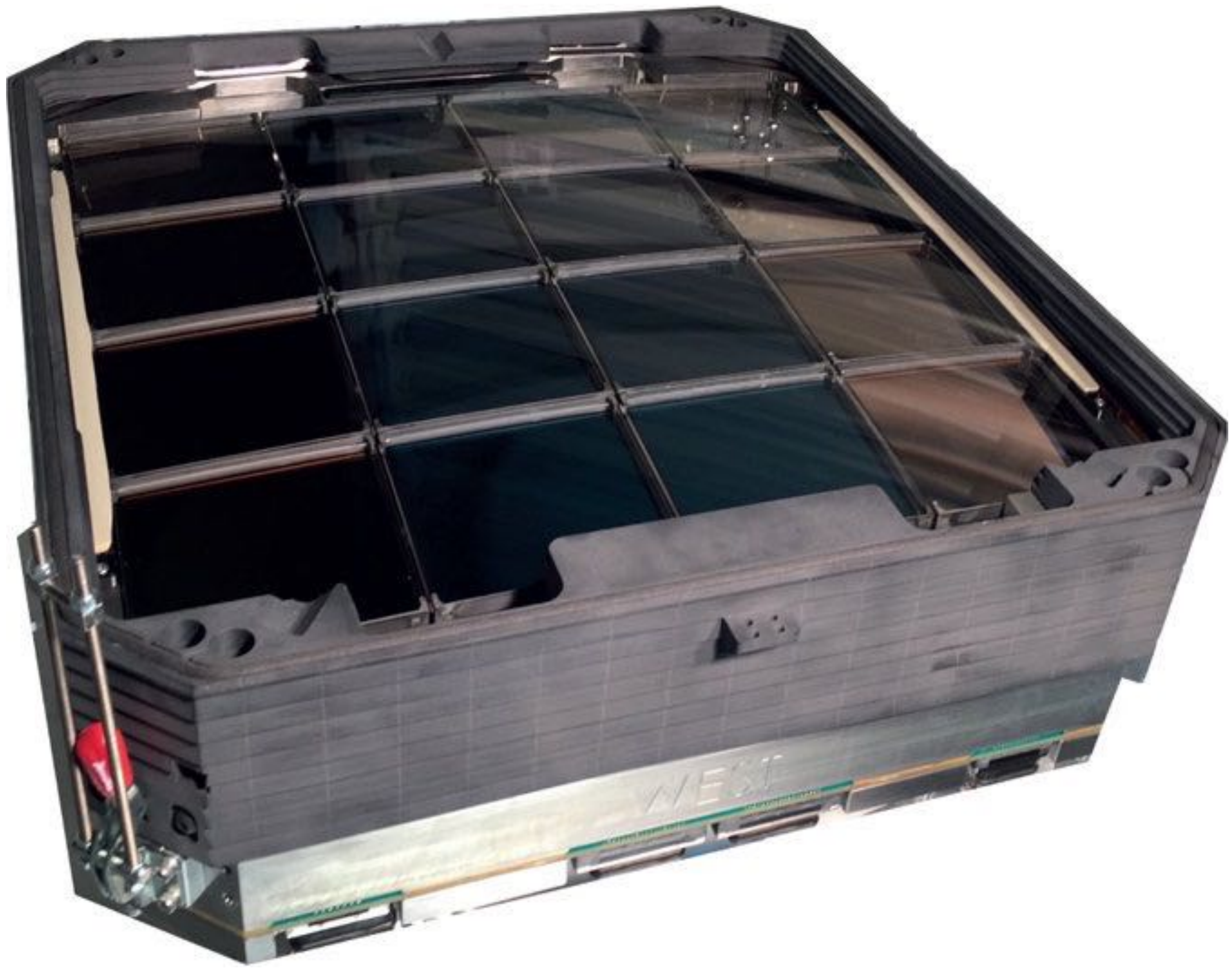




HyperSuprime Cam



Zwicky Transient Factory focal plane



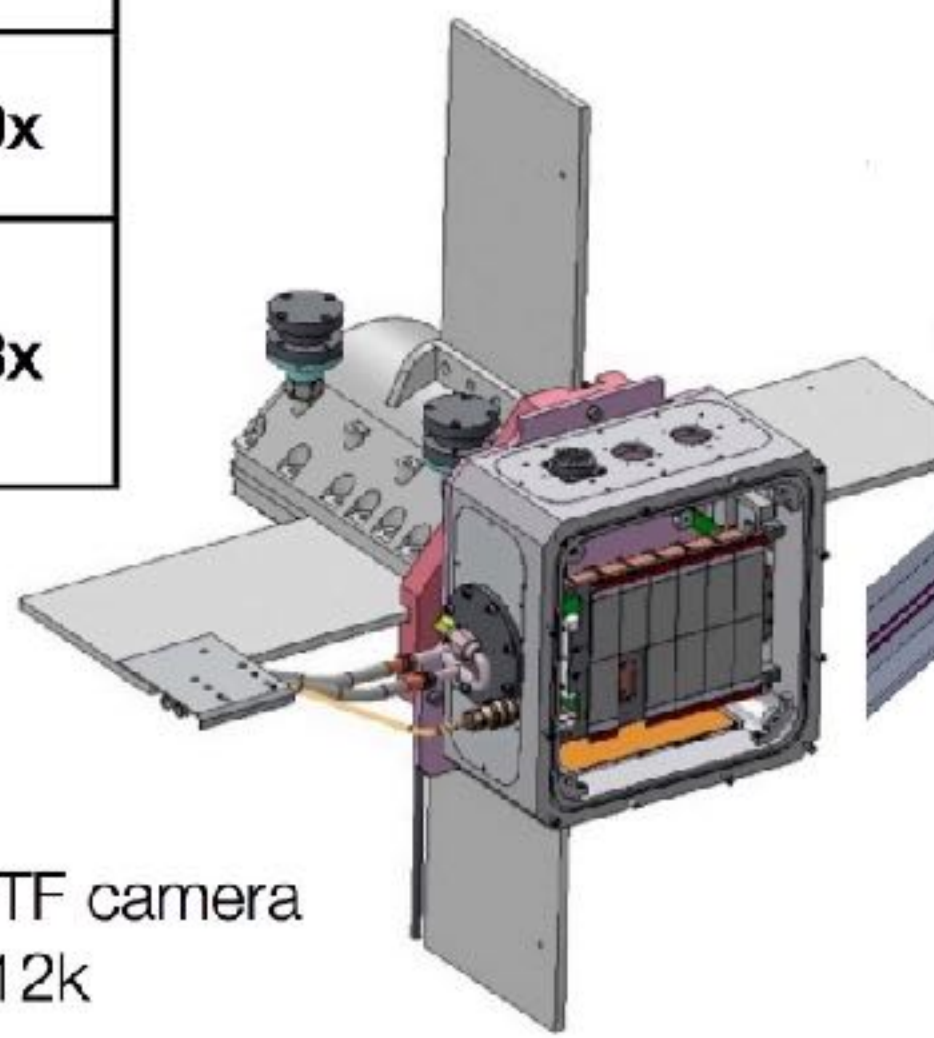
ZTF will survey an order of magnitude faster than PTF.

	PTF	ZTF
Active Area	7.26 deg^2	47 deg^2
Overhead Time	46 sec	<15 sec
Optimal Exposure Time	60 sec	30 sec
Relative Areal Survey Rate	1x	**15.0x**
Relative Volumetric Survey Rate	1x	**12.3x**

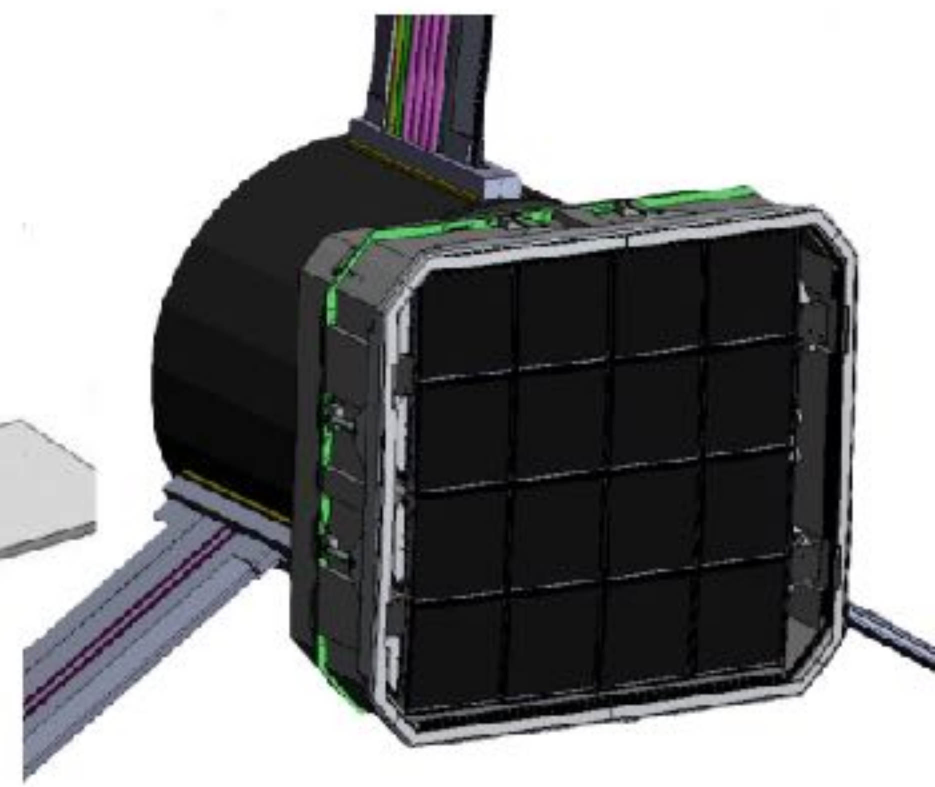
3750 deg^2/hour

⇒ 3π survey in 8 hours

>250 observations/field/year
for uniform survey



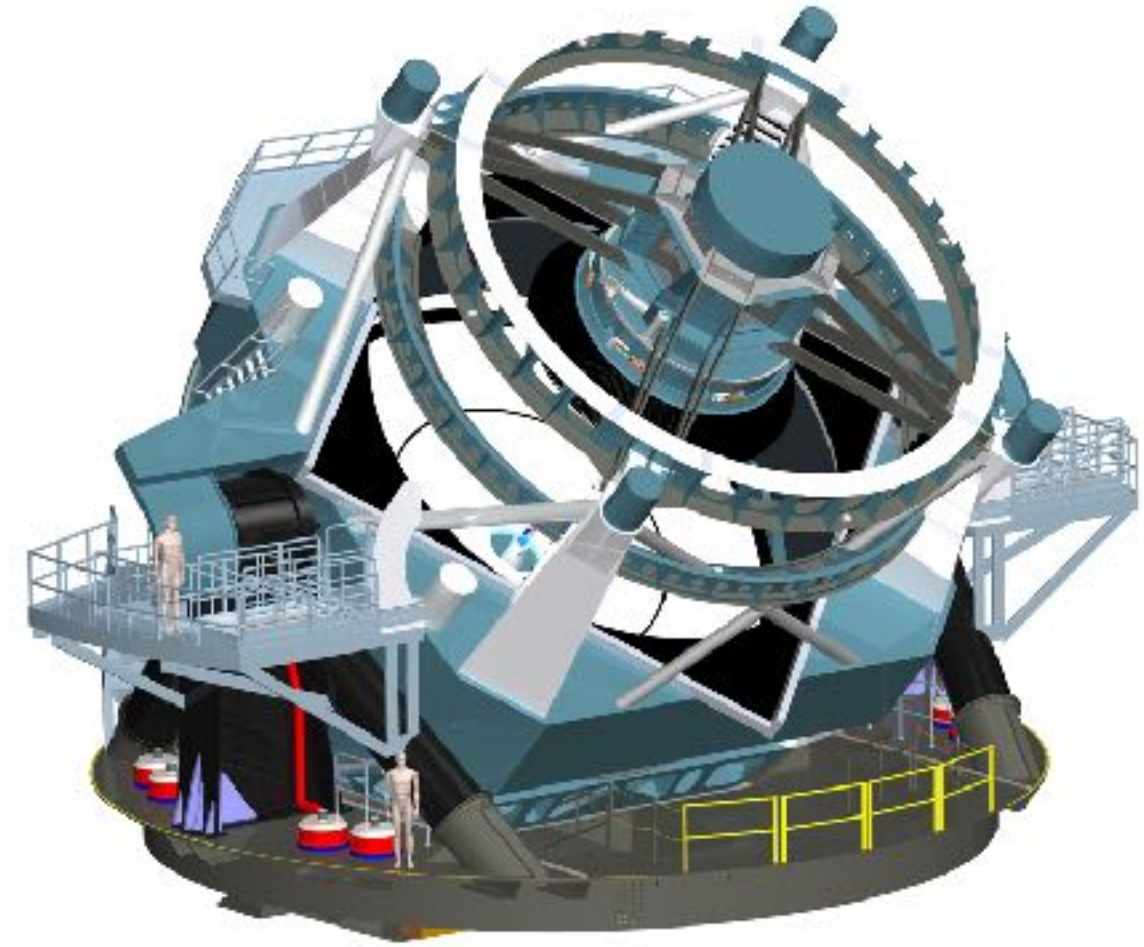
Existing PTF camera
MOSAIC 12k



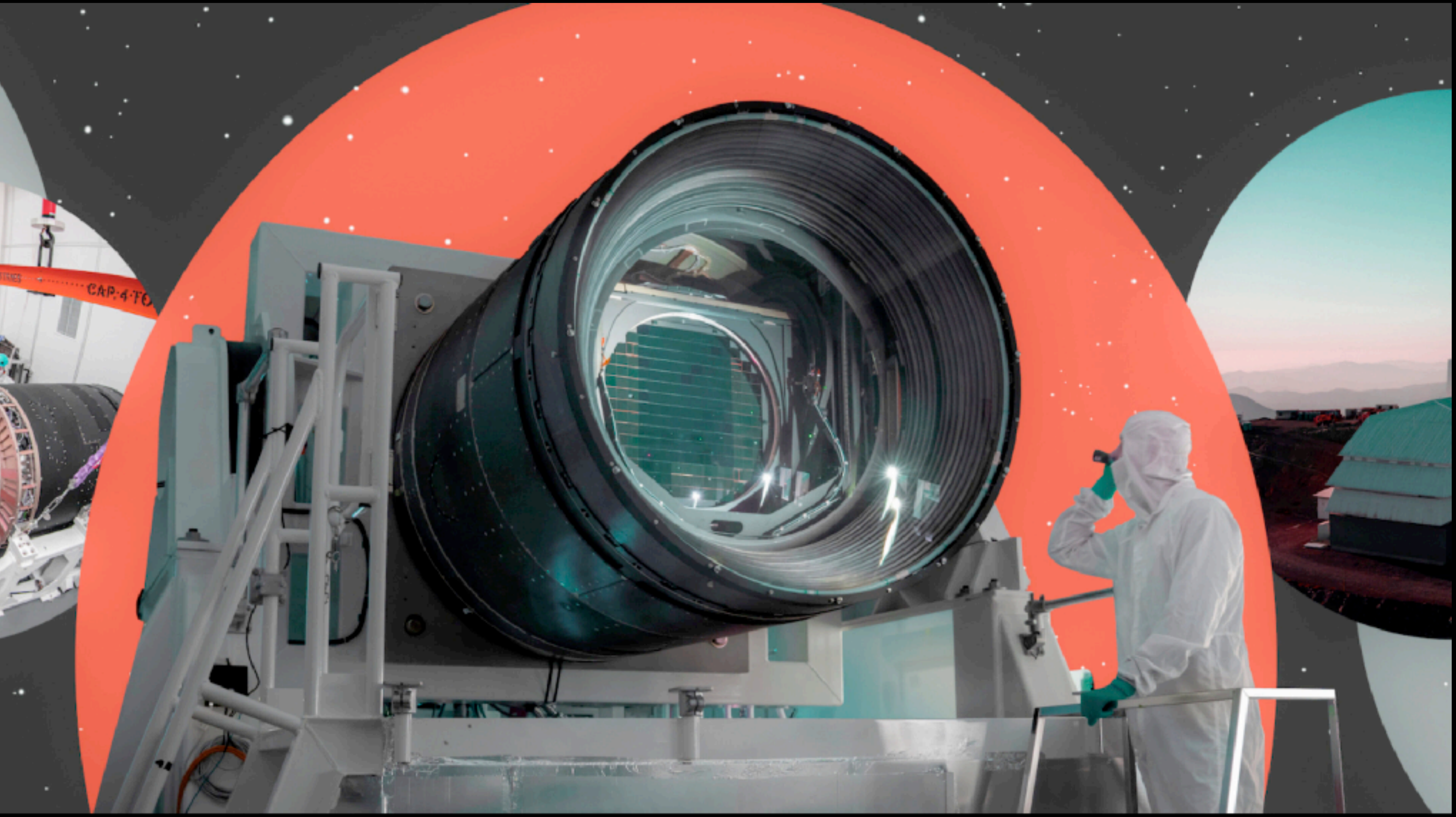
New ZTF camera:
16 6k x 6k e2v CCDs

LSSTCam

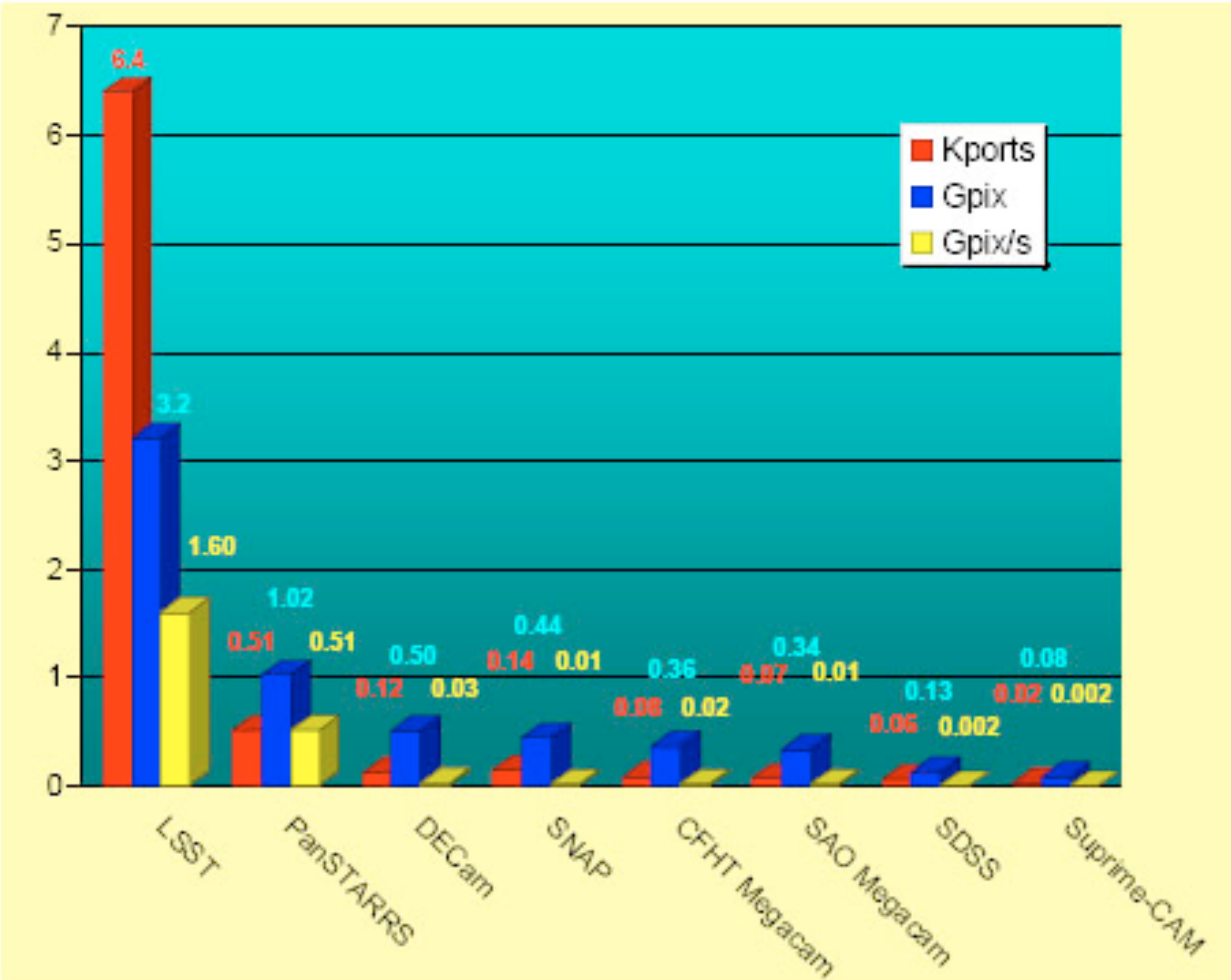
- Commissioning Camera - 1 raft now on the telescope
- 2025-2035
- 3.2 Gpix CCD
- 3.5d diameter. 9.6 deg^2 FOV
- 15s exposure every 20s —> time domain !
- 15-20 TB/night



Size of LSST camera

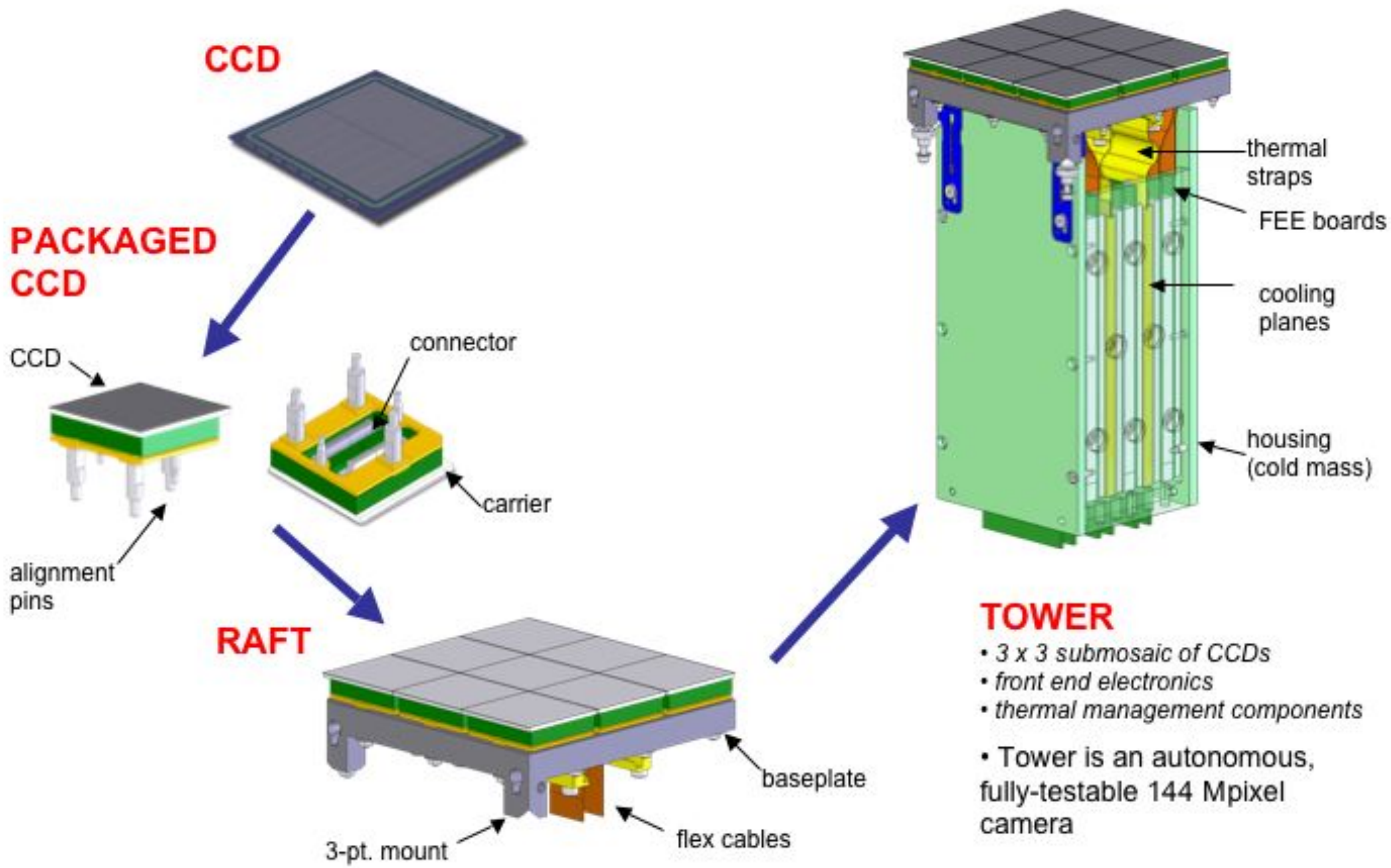



NOTICE



Kports
Gpix
Gpix/s
6.4
3.2
1.60
0.51
1.02
0.51
0.12
0.50
0.03
0.14
0.44
0.01
0.08
0.36
0.02
0.07
0.34
0.01
0.06
0.13
0.002
0.02
0.08
0.002
0
1
2
3
4
5
6
7
LSST
PanSTARRS
DECam
SNAP
CFHT Megacam
SAO Megacam
SDSS
Suprime-CAM

Packaging CCD Arrays

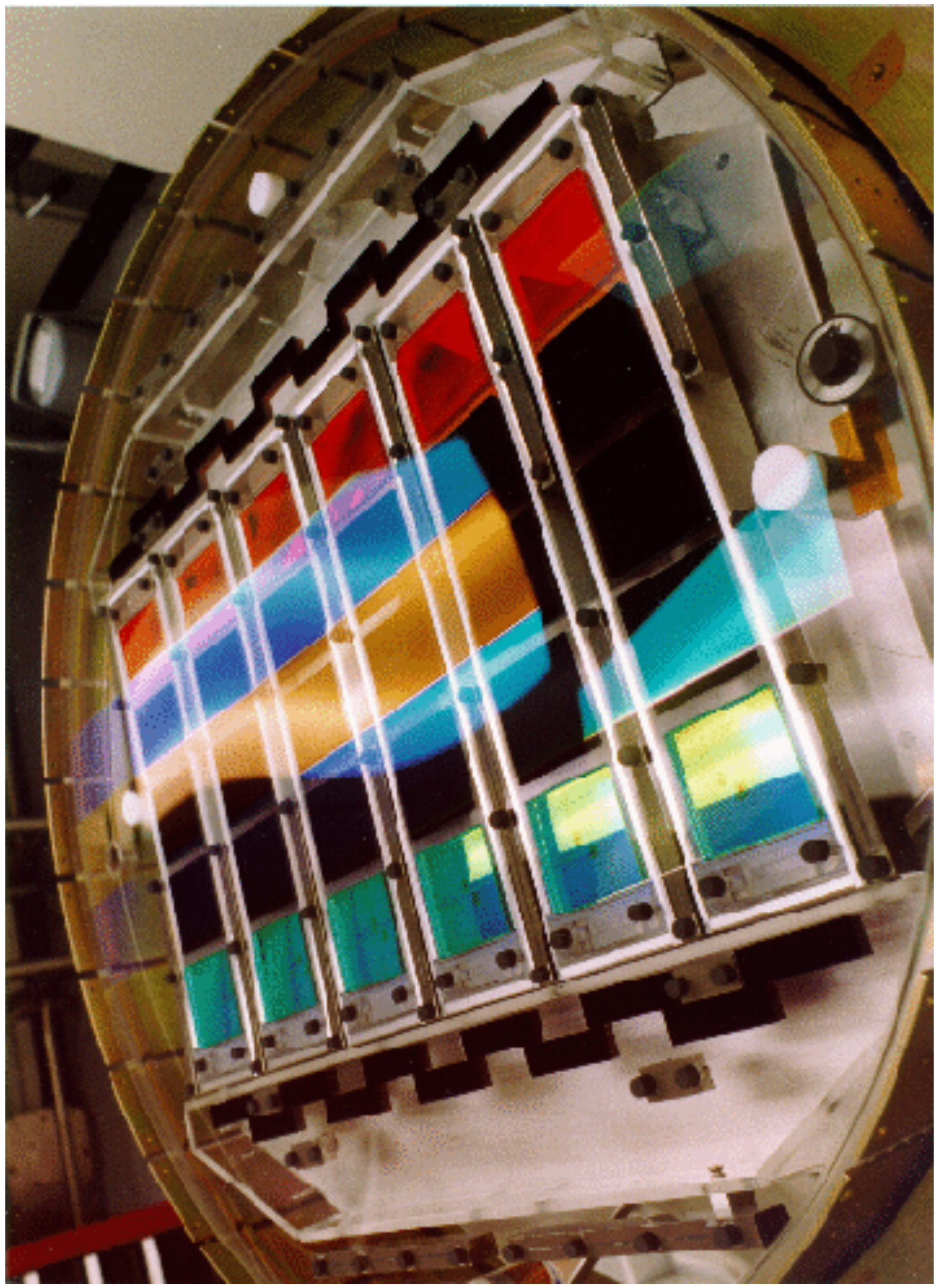


CCD Cameras

name	telescope	pixels	# CCDs	total pixels	area	data rate	data volume	date
PFUEI	Palomar 5m	400 x 400	1	160k	18 mm^2			1976
WFPC	HST 2.4m	800 x 800	8	5M	1k mm^2			1990
SDSS	Apache Point 2.5m	2k x 2k	30	126M	73k mm^2	0.2 TB/night	50 TB	1998
MegaCAM	CFHT 3.6m	2k x 4k	36	340M	62k mm^2			2003
Pan-STARRS	PS1 1.8m	5k x 5k	60	1.4G		5 TB/night	5 PB	2010
GAIA	GAIA 1m	2k x 4k	106	1G	300k mm^2	40 GB/day	73 TB	2013
DECam	CTIO 4m	2k x 4k	62	520M	118k mm^2	0.5 TB/night	1.5 PB	2013
hyper-SuprimeCam	Subaru 8m	2k x 4k	104	870M	210k mm^2	0.25 TB/night	160TB	2014
ZTF	Palomar 48" (1.2m)	6k x 6k	16	600M	157k mm^2	5.65 TB/night	5 PB	2017
LSST	8m	4k x 4k	189	3.2G	320k mm^2	15 TB/night	60 PB	2021

Characterizing array performance

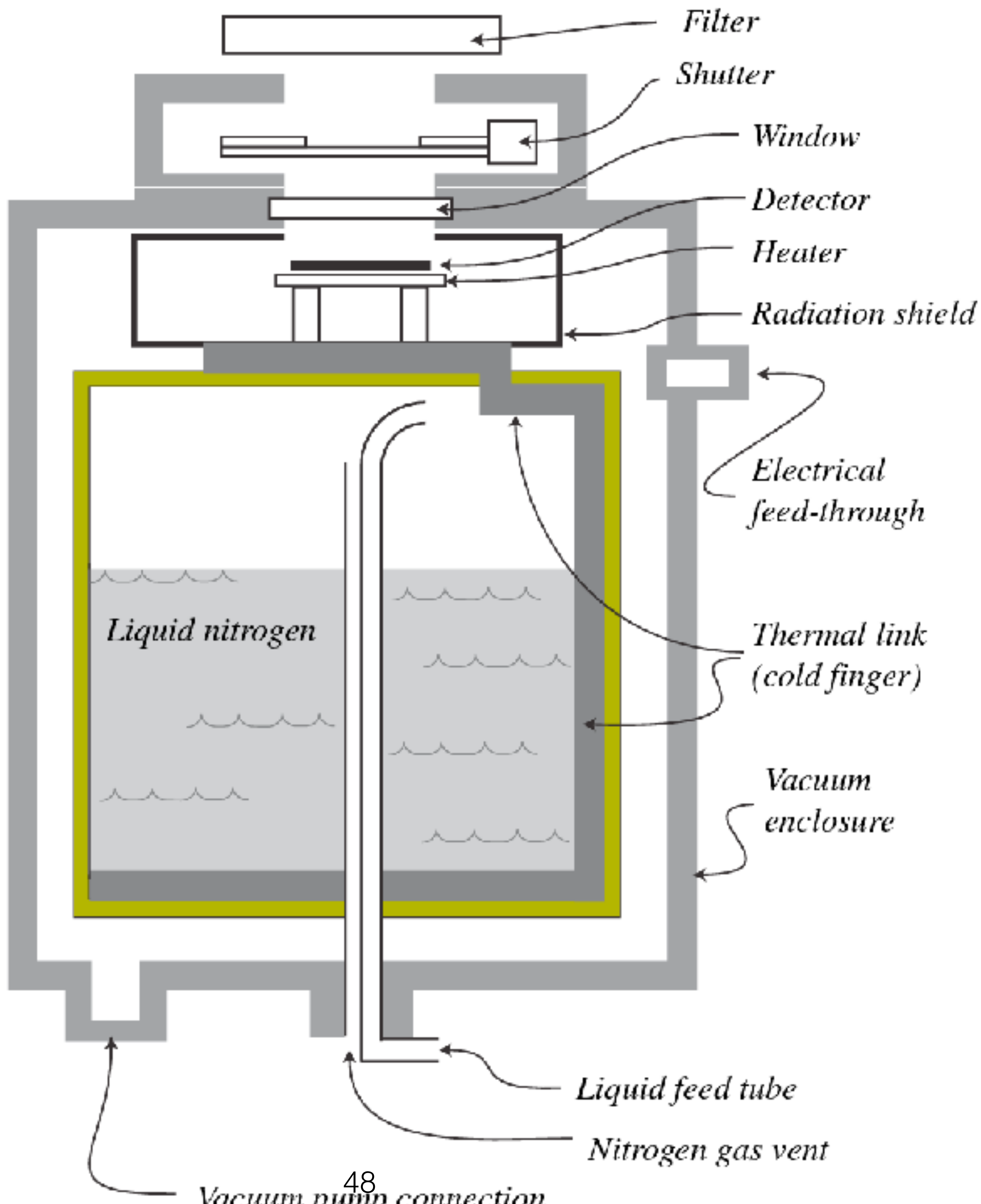
- Quantum efficiency:
  - The efficiency of converting incident photons to photoelectrons.
- Charge diffusion:
  - Migration of stored electrons to adjacent pixels.
- Charge transfer efficiency:
  - Net loss factor in transferring stored charge.
- Readout noise:
  - Number of electrons introduced per pixel on readout. Produced by noise in amplifier and output electronics.
- Dark current:
  - Electrons promoted to conduction band by internal thermal excitation.
- Gain:
  - Number of electrons needed for each unit increase in A/D output.
- Dynamic range:
  - Ratio of full-well capacity to read noise.
- Bias:
  - Intentional positive offset introduced in output electronics.
- Digitization noise:
  - Noise added by discrete number of bits in analog-to-digital (A/D) converter.
- Saturation / non-linearity:
  - Regime in which CCD becomes non-linear or saturated.



SDSS Camera

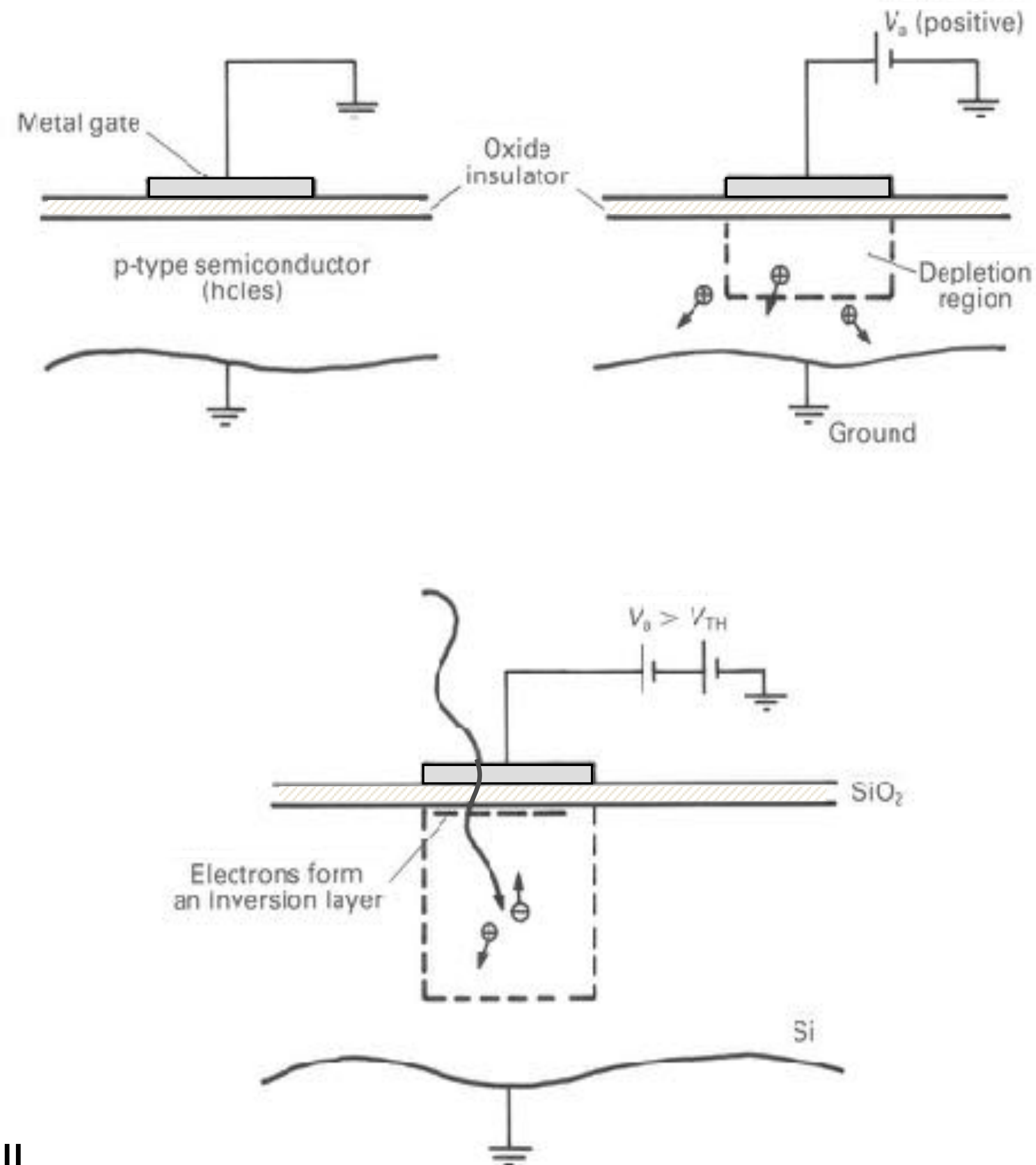
Astronomical requirements

- In astronomical applications:
  - CCDs are not used in video mode – i.e. frames every 1/30th of a second.
  - Long exposures to build up a charge image from a faint source.
  - Readout process cannot be too rapid, because the charge transfer efficiency will be impaired and the electronic noise is greater at higher readout rates.
  - "**Slow scan**" mode:
    - Typically, rates of about 50,000 pixels per second (50 kHz) are used which implies that it will take 20 seconds to read out an array with one million pixels.
    - Requires shutter.
- With such long exposures however, the second problem arises—dark current!
- To permit long exposures, astronomical CCDs must be **cooled** to temperatures well below the freezing point of water, implying a vacuum chamber to avoid frosting.
- The advent of MPP CCDs helped considerably in reducing the cooling requirements, but for the most stringent applications in astronomical spectrographs, more cooling is required and **most CCD cameras at professional observatories use modified liquid nitrogen cooling systems**.


Filter
Shutter
Window
Detector
Heater
Radiation shield
Electrical
feed-through
Liquid nitrogen
Thermal link
(cold finger)
Vacuum
enclosure
Liquid feed tube
Nitrogen gas vent
Vacuum pump connection

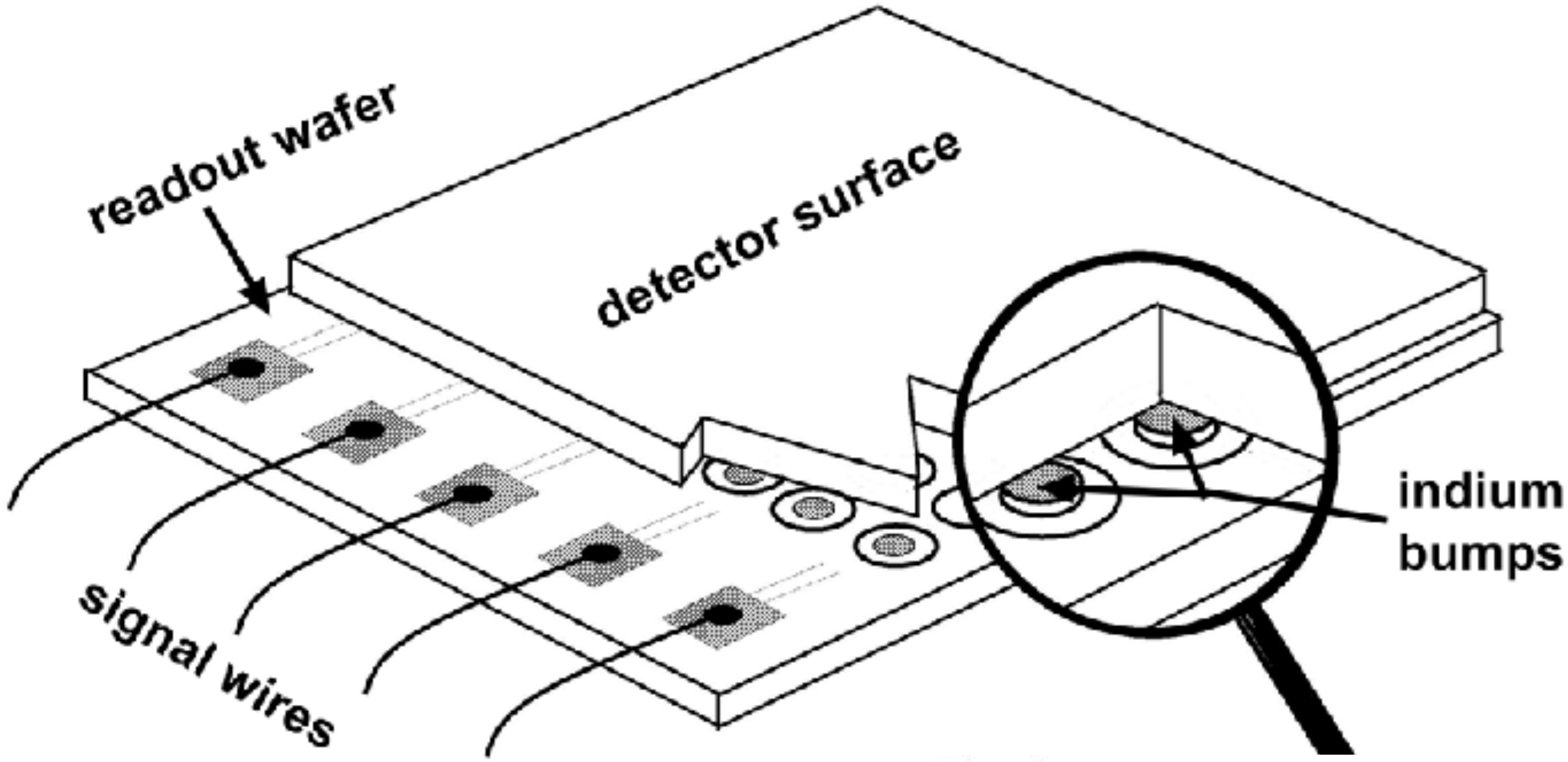
Metal-oxide-semiconductor (MOS) structure

- Apply metal electrodes to the semiconductor silicon together with a thin (100 nm) separation layer made from silicon dioxide, an electrical insulator.
- Structure behaves like a parallel plate capacitor and can store electrical charge. It is called an ***MOS*** (metal-oxide-semiconductor) structure.
- An electric field is generated inside the silicon by the voltage applied to the metal electrode.
- If the material is p-type (excess holes) then a positive voltage on the metal gate will repel the holes which are in the majority and sweep out a region **depleted** of charge carriers.



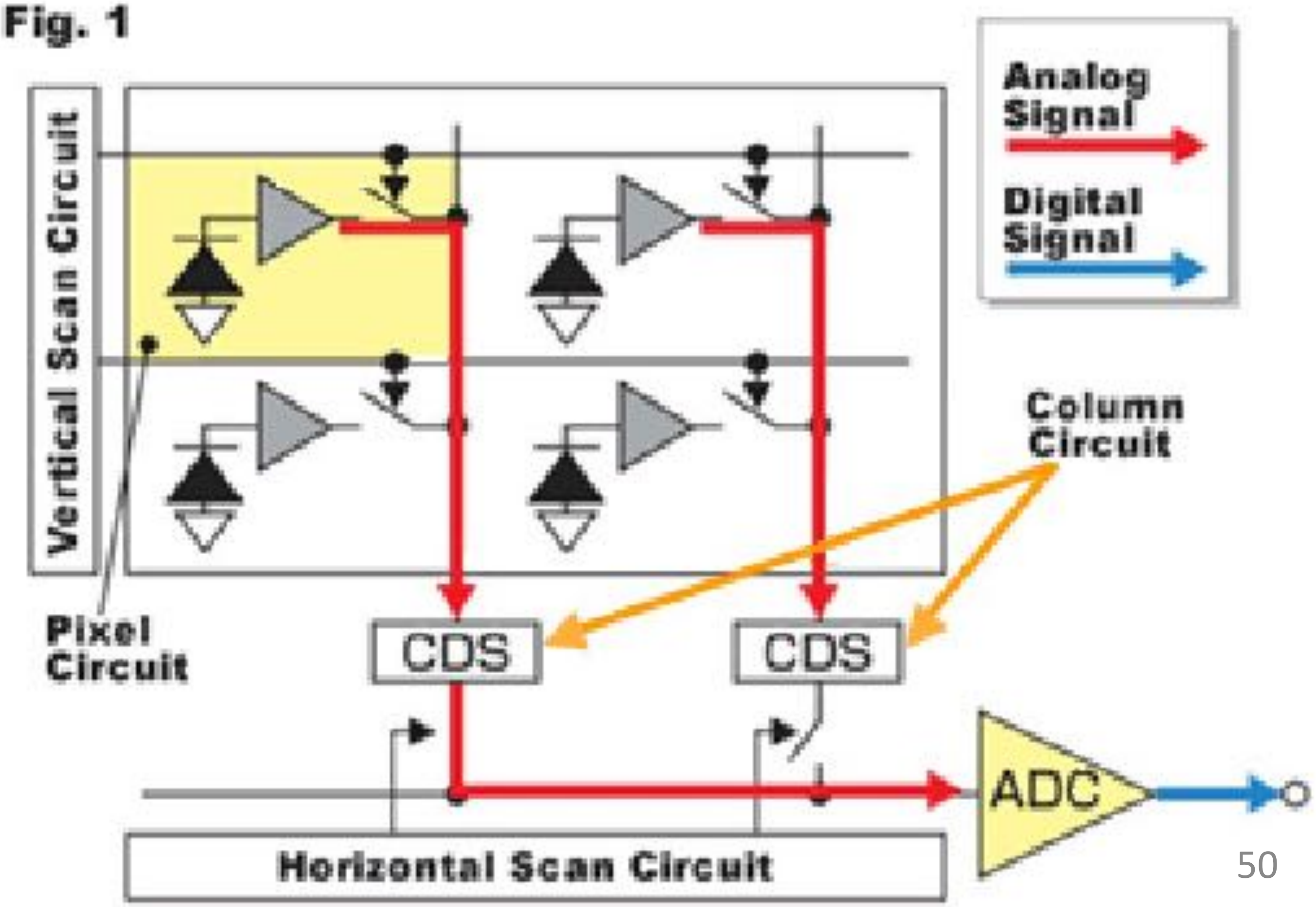
Formation of a MOS potential well.

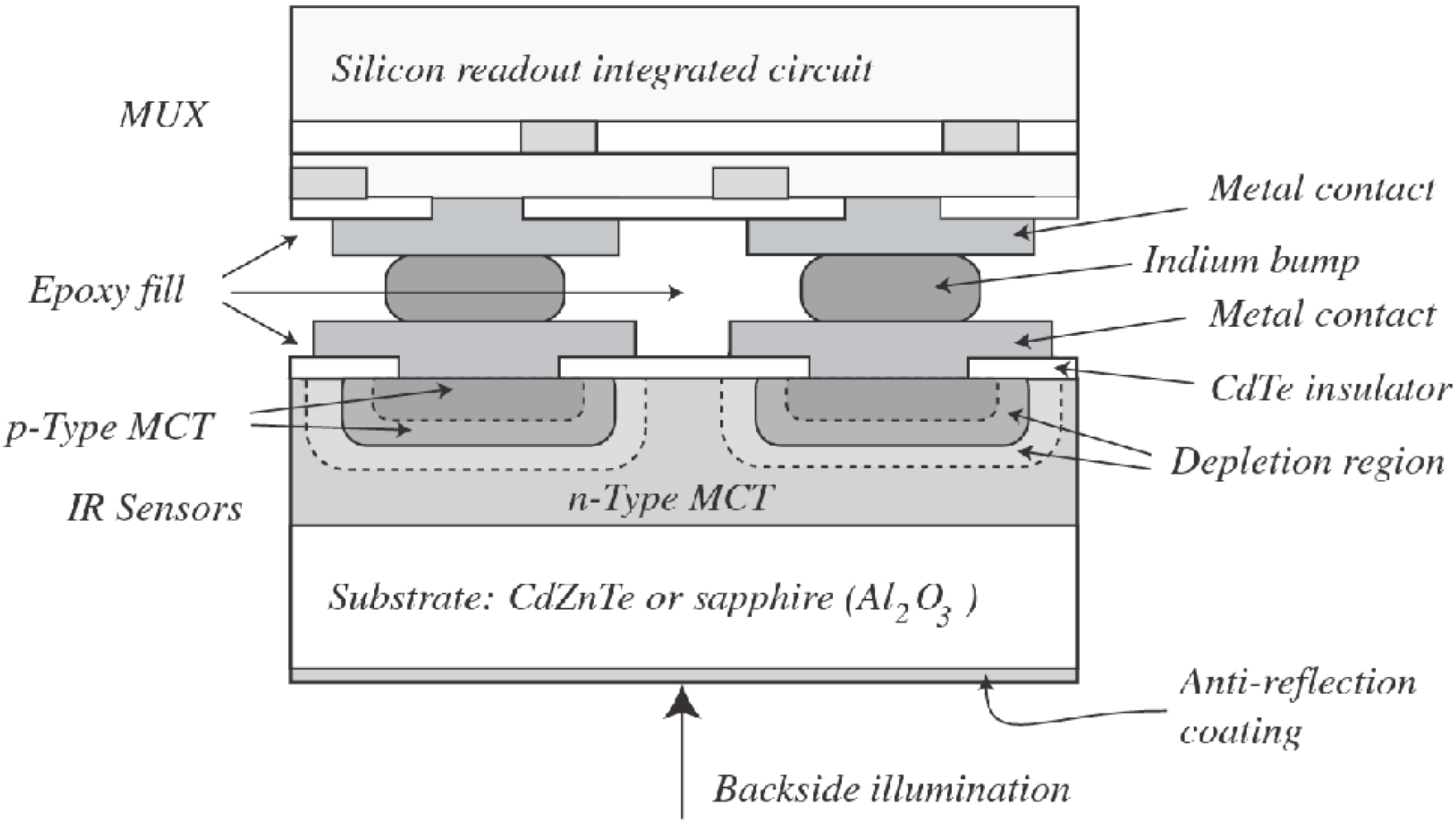
CMOS (complementary metal oxide semiconductor) detectors



CMOS detectors are like taking just the readout for an infrared array and placing photodiodes on the inputs of the amplifiers.

Because they can be made in standard integrated circuit foundries, they are relatively cheap. They can also be made with a lot of the support electronics on the same silicon chip. To the right is Sony's diagram of how one works.




Silicon readout integrated circuit
MUX
Metal contact
Indium bump
Metal contact
Epoxy fill
CdTe insulator
p-Type MCT
Depletion region
n-Type MCT
IR Sensors
Substrate: CdZnTe or sapphire (Al_2O_3)
Anti-reflection
coating
Backside illumination

CMOS detectors can be really big!

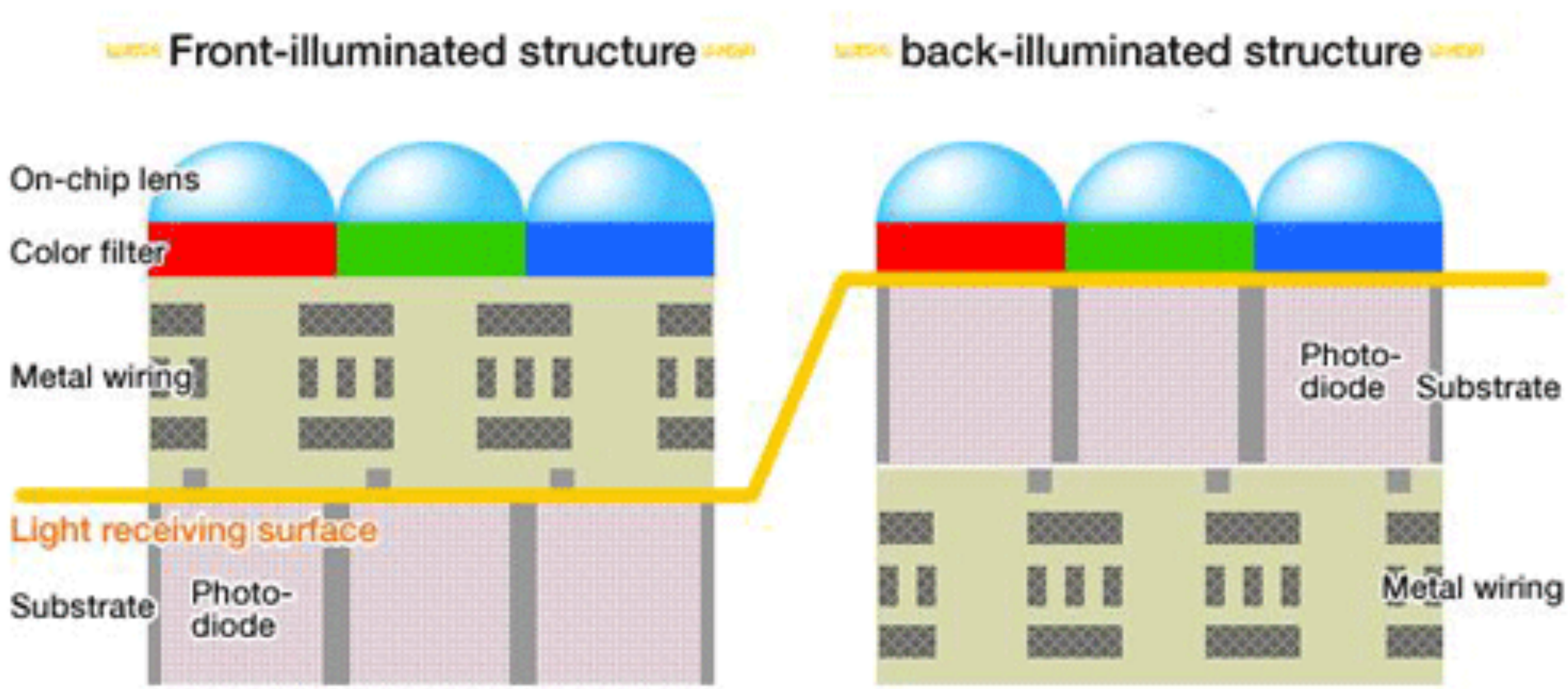
To the left, an X-ray detector; to the right a garden variety 18 Mpix camera detector.

However……..

For very low light levels (e.g., use in astronomy):

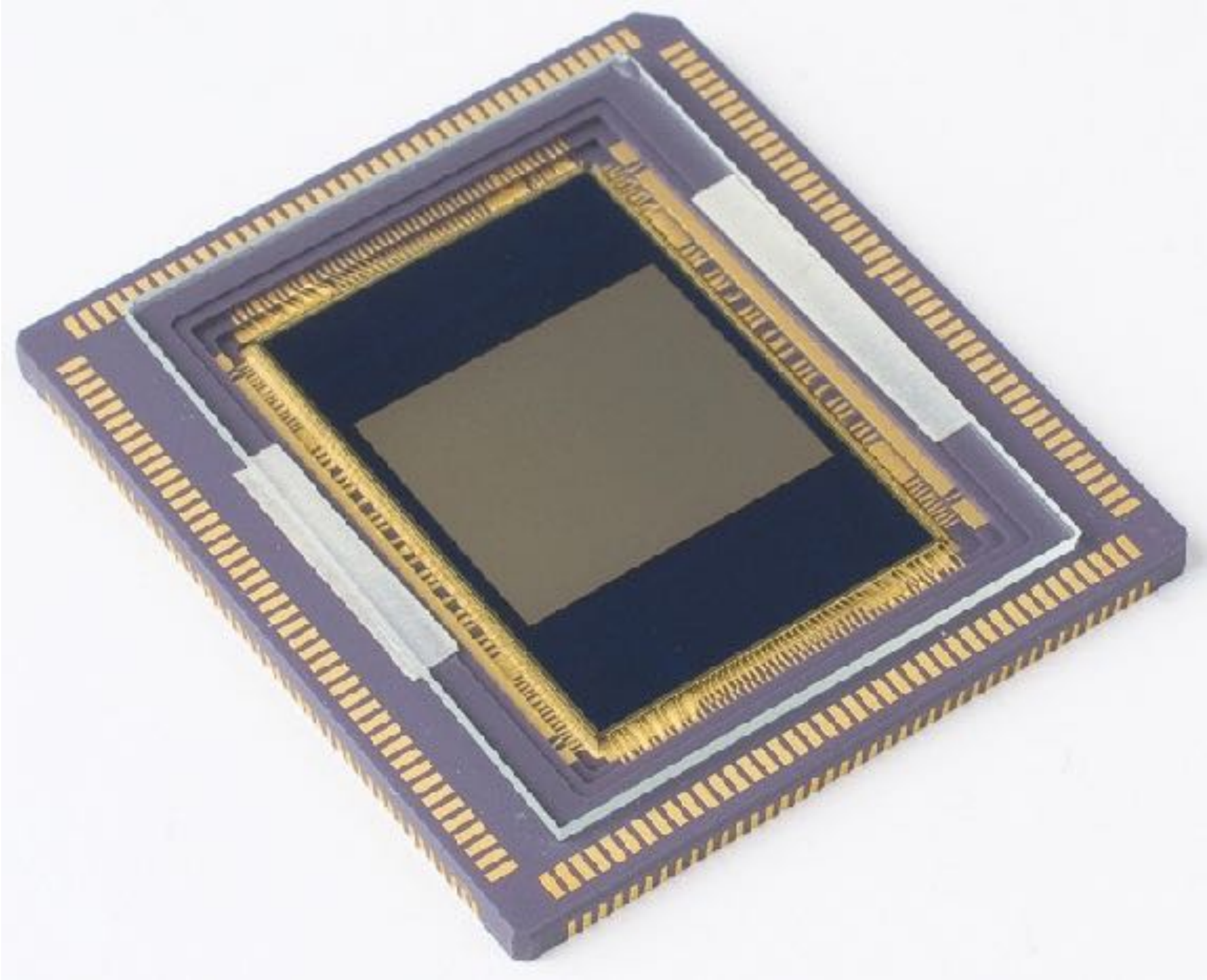
• CMOS detectors have poor fill factors (amplifiers compete for real estate on the chip with the detectors)

• Questions about cryogenic performance: the fill factor issue can be fixed by putting a grid of microlenses over the array (Canon does this for some of its cameras), but it is not clear that such a device can be cooled. Thanks to Sony, camera arrays are now available that are back illuminated:



An example: Fairchild CIS2521F

- 2560 X 2160 pixels
- 6.5 μm X 6.5 μm pixel pitch
- Readout speed maximum: 100 frames per second
- Read noise < 1.5 electrons rms
- Peak quantum efficiency > 52%



Calibrating Frames

- We assume that the CCD electronics system is perfect and so there are no other sources of unwanted noise save the irreducible readout noise. The simplest approach is to construct a "final frame" by performing the following steps to subtract dark current and normalize with a flat-field:

$$FINAL\ FRAME = \frac{OBJECT\ FRAME\ \text{-}\ DARK\ FRAME}{FLAT\ FRAME\ \text{-}\ DARK\ FRAME}$$

- The dark-subtracted flat-field frame in the denominator is usually "normalized" beforehand by dividing each pixel value with a constant equal to the mean or median of the entire frame; this step is included in the noise formulation below.
- For infrared arrays, the (brighter) night sky is often used as the flat-field source and so another version of this equation would substitute SKY FRAME for FLAT FRAME because the “sky-flat” is derived from many dithered frames of a relatively blank field of sky.

Dark Current & Cooling

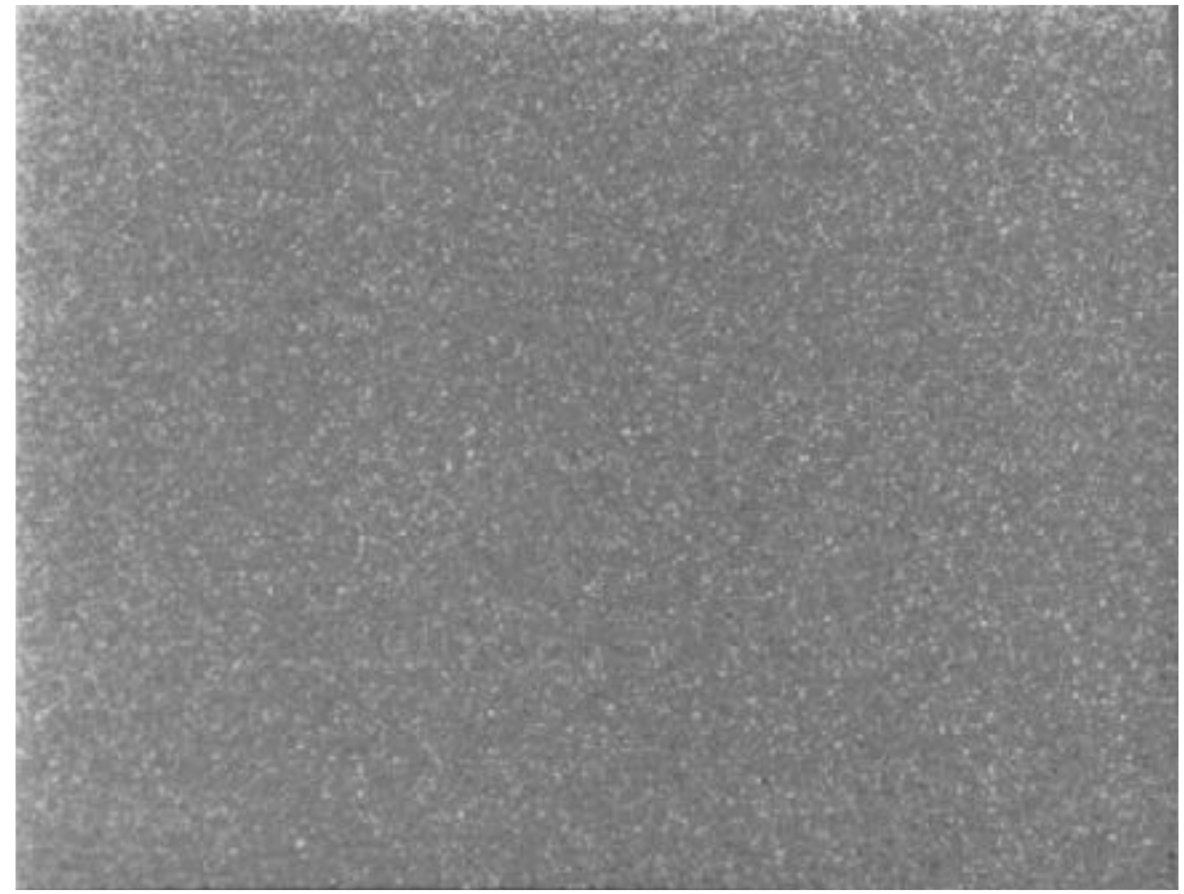
- Random (Brownian) motions of atoms at normal room temperatures within the silicon lattice will release sufficient energy to give rise to a continuous stream of electron-hole pairs in the absence of light.
  - This process for electron-hole pair production is called "thermal” because the energy source is heat and the “dark current” produced can be very substantial.
- At room temperature the dark current of a typical, non-inverted CCD is about 100,000 electrons per second per pixel (or equivalently, 1.8×10^{-9} A/cm^2 for 30 µm pixels) which means that the CCD storage wells will fill up, saturate and spill over in just a few seconds.
- If the CCD is read out rapidly and continuously at a high rate (frequency) such as 5 MHz then dark current is cleared after only a brief accumulation time (~1/60th of a second), so the CCD can be used for TV applications.
- Most astronomical sources are much too faint to yield a good image in such short snapshots, especially as there is a small but definite penalty of increased noise for every read out.
- Fortunately, the solution is straightforward. The CCD needs to be cooled to a low temperature.

Dark current sources

- For charge-coupled devices there are three main sources of dark current. These are
  - thermal generation and diffusion in the neutral bulk silicon
  - thermal generation in the depletion region
  - thermal generation due to regions called "surface states" at the junction between the silicon and the silicon dioxide insulation layer.
- Of these sources, the contribution from surface states is the dominant contributor for multi-phase CCDs.

Bias Frame

With no photons, the CCD (or IR array) electronics system will still produce a small non-zero reading for each pixel. This electronic signature is known as the **bias level** (given the symbol b in Eq. 9.5). It can be determined by taking a short unexposed frame (~zero exposure time, shutter closed), also called an “erase” frame. An example of a bias frame is shown in Fig. 9.6.



A bias frame should contain very little “fixed-pattern” structure; it should be dominated by random readout noise. The standard deviation of a good-sized patch of the array detector therefore gives an immediate estimate of the readout noise (R/g) in data numbers. If there is unavoidable fixed patterns in the bias frame then it is straightforward to take the difference between <u>two</u> bias frames to eliminate this effect. The measured noise distribution in the <u>differenced</u> image is simply $\sigma = (\sqrt{2})(R/g)$. It is larger by root 2.

Flats

The point of the flat is to calibrate the **relative** response (aka gain, calibration) of the array (ie each pixel).

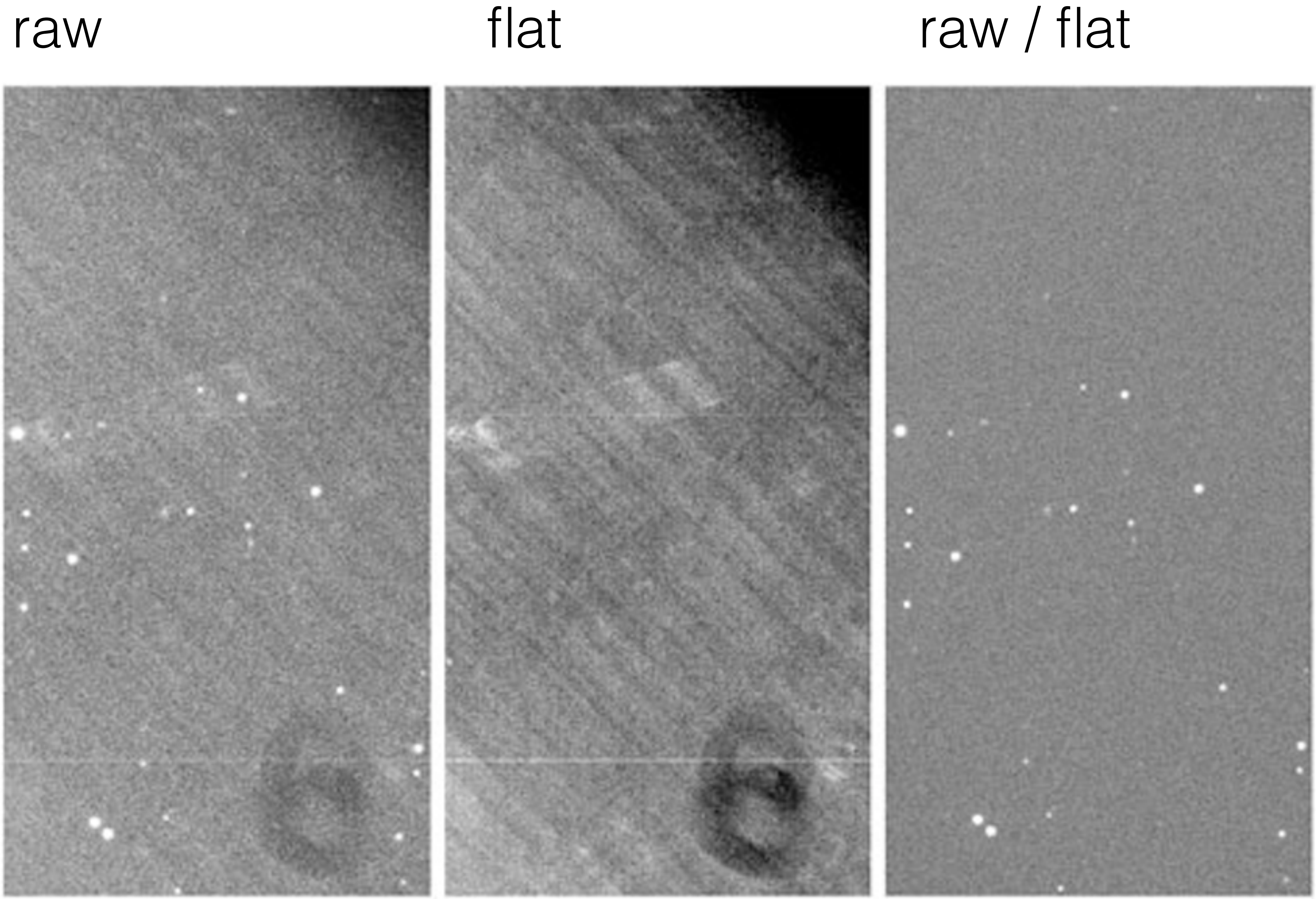
Twilight flats — Take image of the sky at twilight. This is a roughly uniformly bright source, fills the entire FOV. Can get field uniform to ~1%. Not doing anything at this time, anyways. BUT there is usually a gradient across the field and the spectrum is different than night sky.

Dark-sky flats — Take the image in the conditions in which you observe. BUT stars ! Also, need long exposure times to get ~10^4 photoelectrons. 10^2—10^6 times longer exposures than twilight flats.

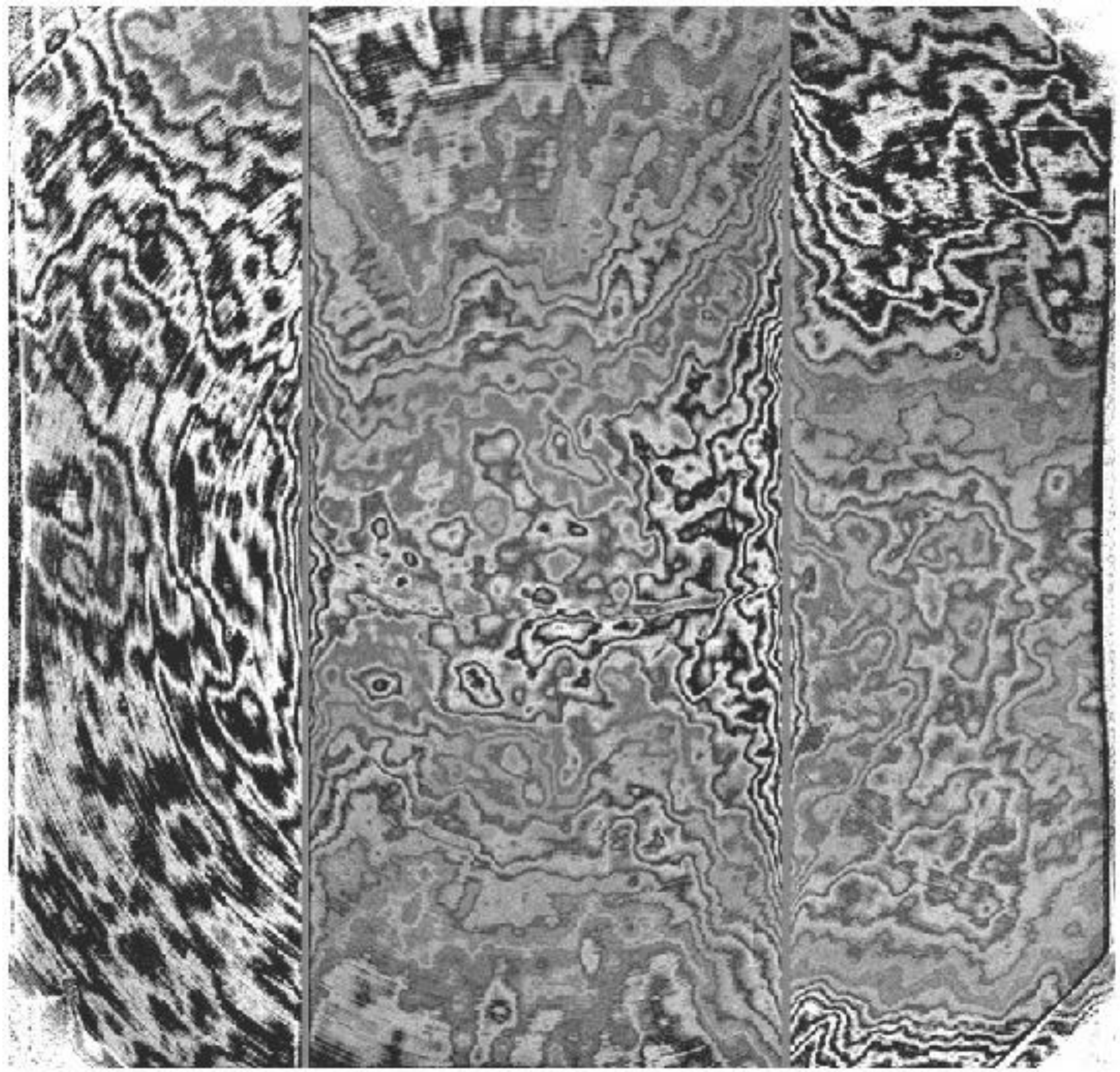
Dome flats — Put a white screen up in the dome, illuminate with a known source. Repeatable ! Unwanted light from scattering and reflections.

Averaged flat — Average all of your science observations. Works for surveys of blank fields. Increasingly common.

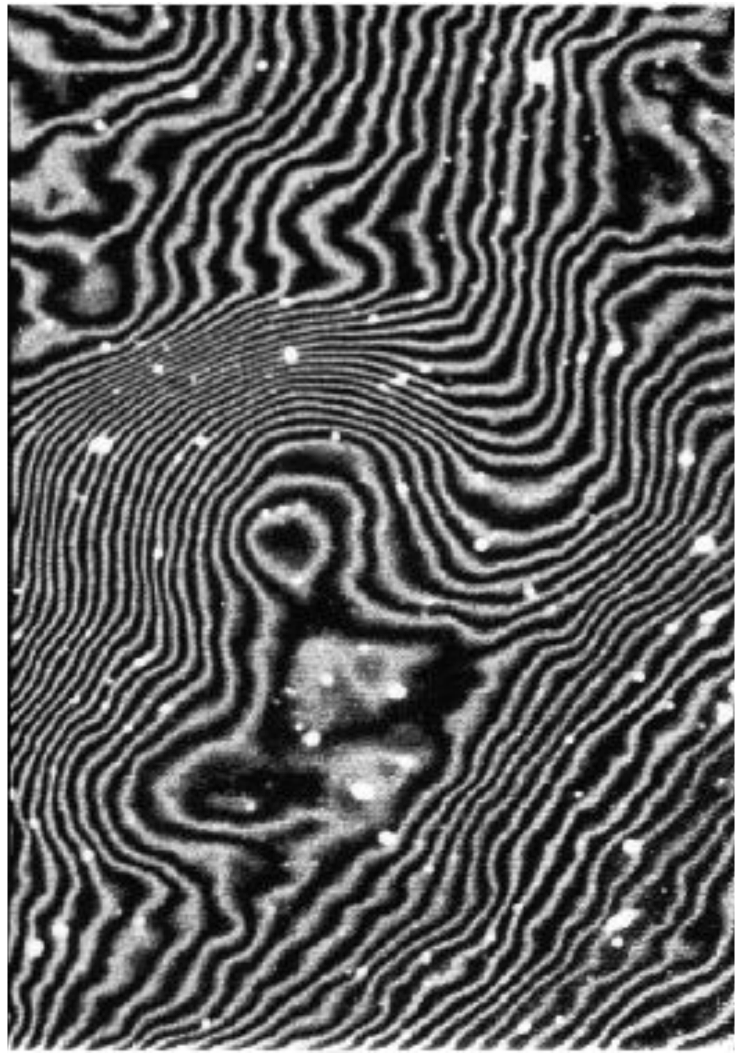
Compound flats — Do all of the above, combine them in a sensible strategy. e.g. dome or twilight flat for small spatial scales, smoothed version of dark or averaged flats for large scales,

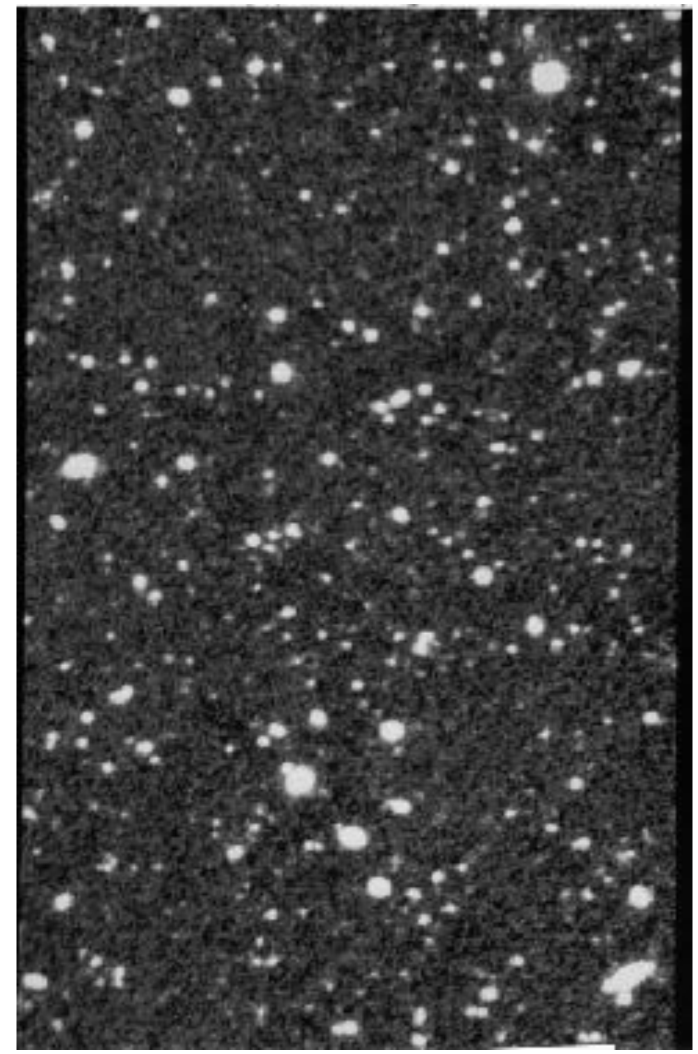

raw
flat
raw / flat

Fringing from OH sky lines



Interference Fringes



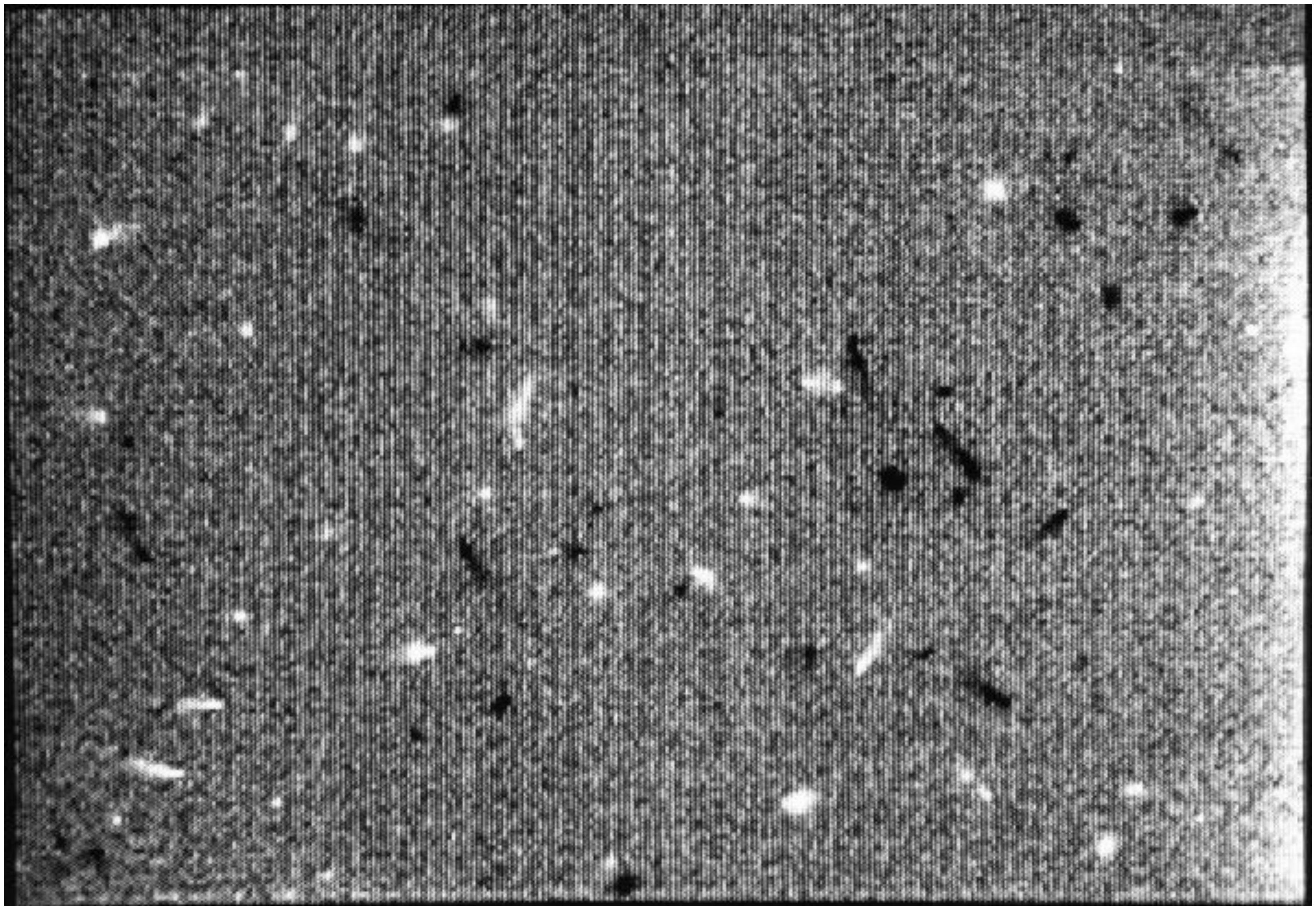


Adaptive modal filtering: computes the absolute difference between the mean and the median of values associated with a pixel over all the images in a set and rejects deviant values until this difference falls below a certain value or a maximum number of values have been rejected. A given pixel is then median-filtered over all the images. This technique is suitable for relatively un-crowded fields in which 2/3 of the actual CCD area is occupied by sky. The “fringe frame” thus created must be scaled by trial-and-error and subtracted from the object frame until a patch of sky is entirely flat. Figure 9.9 shows a 500-s CCD exposure in a far red band on a 4-m telescope before and after fringe removal. The correction effect is remarkable!

Cosmic Rays

- High-energy sub-atomic particles (protons) entering the Earth’s atmosphere from space generate a shower of secondary particles (muons) which can be "stopped" by a sufficiently thick layer of silicon.
- The energy released can generate ~80 electrons per pixel per micron of thickness in the silicon. With a collection depth of 20μm, a muon event is seen on a CCD image as a concentrated bright spot of a few thousand electrons.
- Thinned backside illuminated CCDs and those with epitaxial (thin) collection layers are less prone to cosmic rays than thicker frontside illuminated arrays.
- High altitude observatories may encounter more problems.
- In space the problem is even more severe.
- The Galileo CCD camera would have arrived at Jupiter in a saturated state, but neutron testing before launch revealed large dark spikes in the array. The problem was solved by operating the CCD at -120 C instead of -40 C.
- Likewise, the operating temperature of the Cassini and Hubble CCDs was selected to be -90 C to lessen the effect of traps induced by proton impacts.
- CCDs on the Chandra X-ray Observatory experienced a proton-related radiation problem that seriously degraded charge transfer efficiency soon after launch (Janesick 2001).

Figure 8.7 is the “difference” of two long exposures showing multiple hits in each exposure.



The only approach is to take multiple frames of the same region and compare them, as there is very little chance of a cosmic ray event hitting the same pixel two times in quick succession.

Steps to reduce an optical CCD image

1. **Subtract** the offset frame from the data, dark and response frames to obtain data', dark', and response'
2. **Scale** dark' to exposure time of data and response and subtract from them to get data'' and response''
3. **Divide** data'' by response''

data' = data - offset

dark' = dark - offset

response' = response - offset

data'' = data' - αdark'

response'' = response - βdark'

final = data'' / response''

offset = amplifier offset

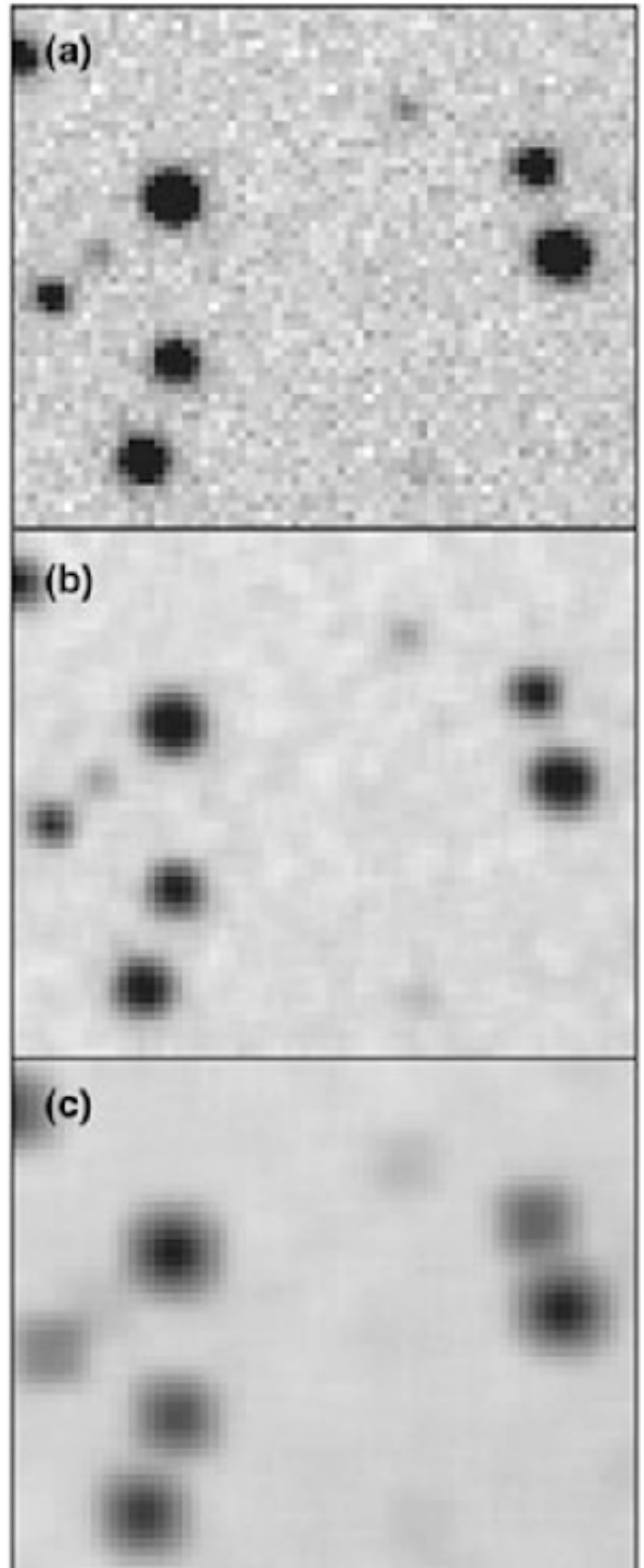
dark = dark current

response = flat field

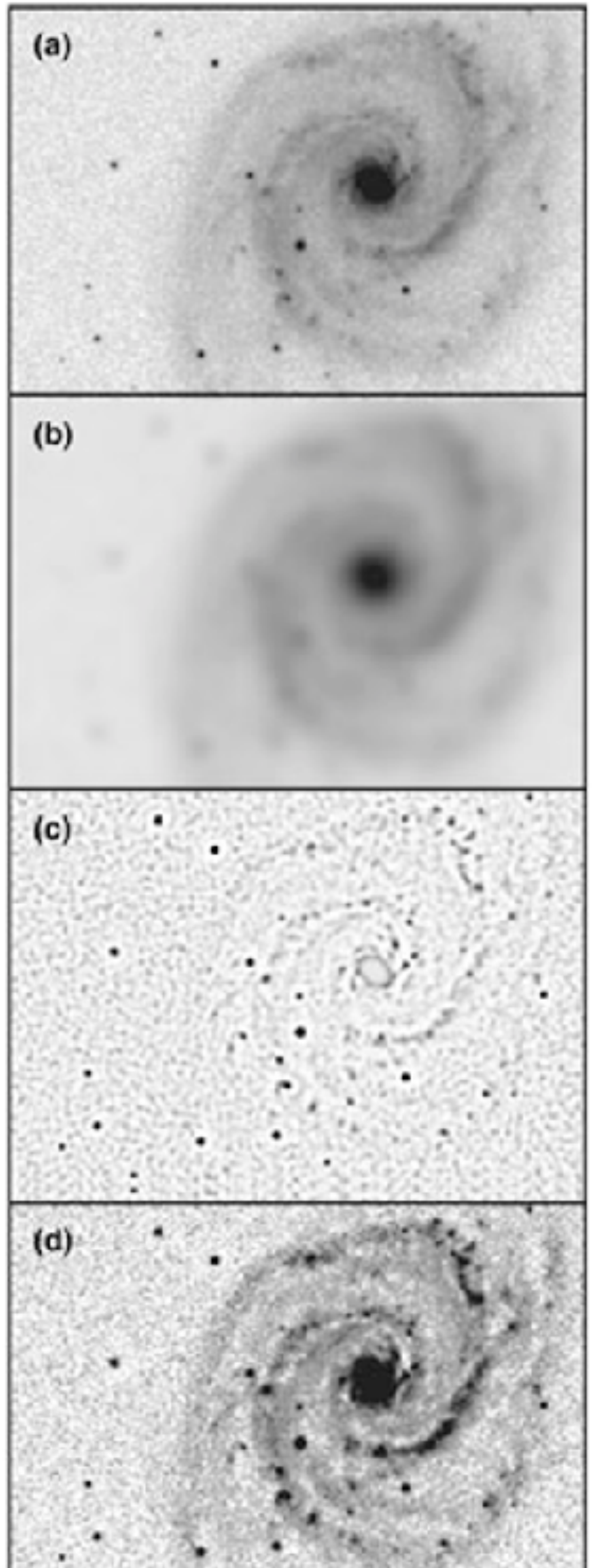
α = scaling to make dark field the same exposure as data

β = scaling to make dark field the same exposure as flat field

smoothing



filtering



M51

7-pixel FWHM Gaussian filter

high-pass (Laplacian) filter

“unsharp mask”
(to emphasize point sources)

Detective Quantum Efficiency (DQE)

- *DQE is defined as the quantum efficiency of an idealized imaging system with no readout noise but which produces the same signal-to-noise ratio as the actual CCD system in question*.
- In the ideal case, the signal-to-noise ratio for a CCD pixel is given by

$$\frac{S}{N} = \frac{\eta N_p}{\sqrt{(\eta N_p + R^2)}} \tag{9.3}$$

- where η is the quantum efficiency at the wavelength of concern, N_p is the total number of photons incident on the pixel in the exposure time, and R is the root mean square (rms) value of the readout noise. (For simplicity we ignore any correction to R due to digitization noise.) An *idealized* detector with *no* readout noise would have QE equal to η' and a noise, given by Poisson statistics, of the square root of $\eta' N_p$. Equating the two signal-to-noise ratios yields an expression for the DQE (η') of

$$\eta' = \eta \frac{1}{\left(1 + \frac{R^2}{\eta N_p}\right)} \tag{9.4}$$

- which shows that η' is less than η, and that the DQE of a CCD is dependent on the signal level (N_p) (Table 9.1). To keep the DQE within 10% of the QE, N_p must be greater than $10 \times R^2/\eta$.

Assuming you have done everything right, you need to calibrate your raw data:

Calibration must take into account the differing properties of the detectors in the array:

- pixel-to-pixel variations in amplifier offset
- pixel-to-pixel variations in dark current
- pixel-to-pixel variations in responsivity

Three unknowns require three sets of data:

- Offset frame (sometimes called bias frame): very short exposure, no signals
- Dark current frame: long exposure, no signals
- Response frame (sometimes called flat field): uniform illumination

Image data reduction consists of:

- Subtract offset from data, dark, and response frames to obtain data’, dark’, and response’.
- Scale dark’ to exposure time of data and response and subtract from them to get data’’ and response’’
- Divide data’’ by response’’

The result: if the data frame has a uniform exposure, then the product will be a uniform image at a level corresponding to the ratio of the exposure on the data frame to the exposure on the response frame (exposure = level of illumination multiplied by the exposure time). Sources will appear on top of this uniform background.

For best results:

- Dark current and response frames may need to be obtained close in time to the data frames
- It may be necessary to use identical integration times for dark, response, and data frames
- Response and data frames should be taken with illumination of identical spectral character
- Need a minimum of 3 frames on source – 5 or more is better – to be sure there are no transient bad pixels (e.g., cosmic ray hits)
- Permanent bad pixels can be masked out by replacing values with the average of those from surrounding pixels. However, doing so can give bad data that looks good. It is better to take multiple exposures and move the source on the array between them to fill in the source structure with all good data – bad pixels then just reduce the integration time at some points in the image.

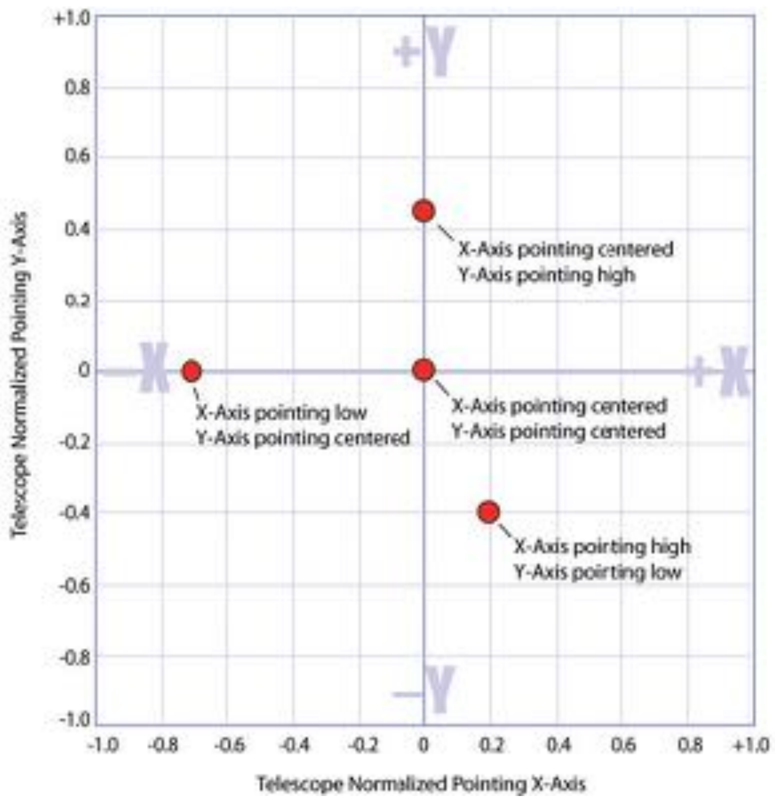
A good strategy for imaging is:

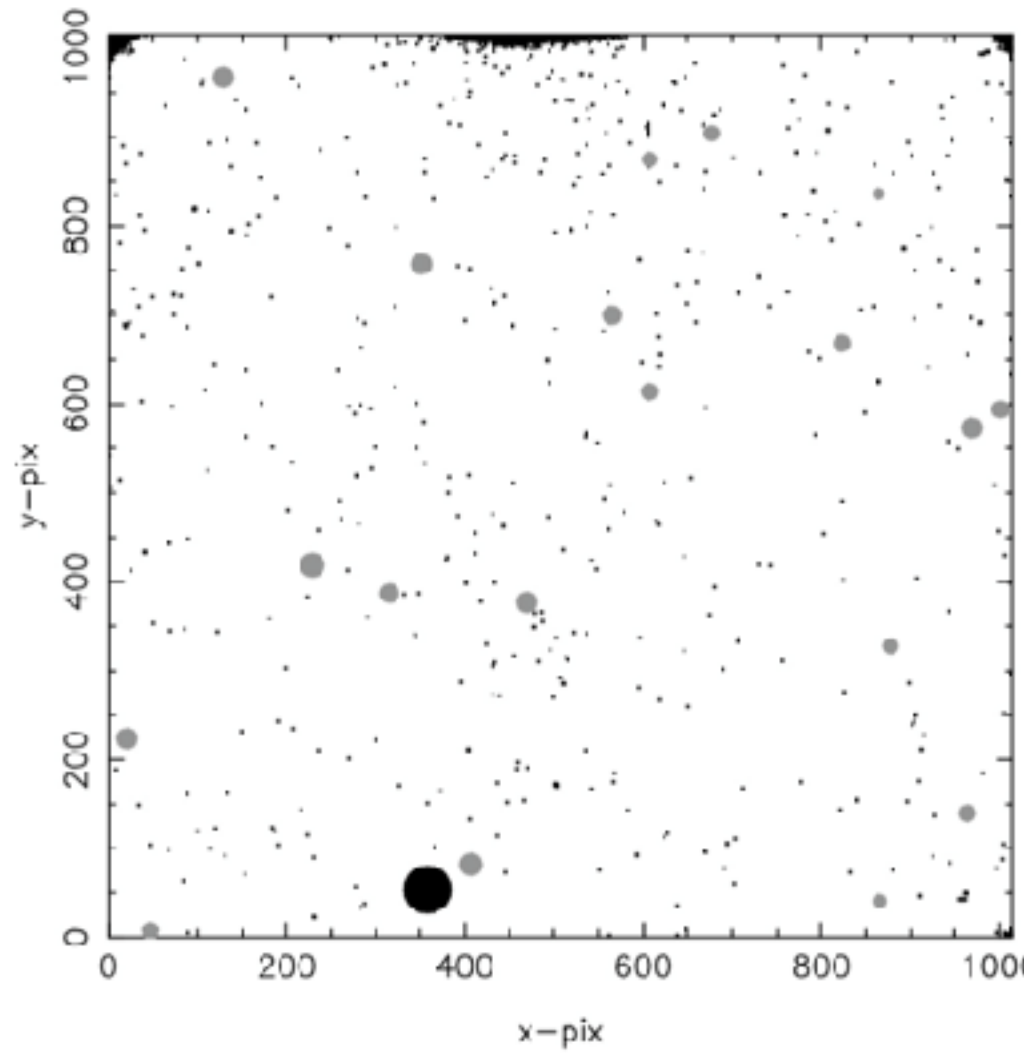
- Take repeated exposures of the field, moving the source on the array between exposures
- Generate the response frame by a median average of these frames – sources will disappear because they do not appear at the same place on any two frames
- Obtain dark frames with the same exposure time as used for the data and response frames
- Subtract dark from data and response (also takes out offset); divide corrected data by corrected response
- Shift frames to correct for frame-to-frame image motions
- Median average again to eliminate bad pixels and cosmic rays, while gaining signal to noise on the source image

What are we going to do?

- Open the files with python.
- Look at the images.
- look at the values of the histograms.
- trim the data image (x,y)
- remove the offset from the images
- figure out how to scale the dark and flat
- make a pretty picture
- calibrate the image and extract a flux for a source

CCD Observing: Dithering





Dithering is a technique where an observation is divided into multiple exposures that are spatially offset by moving the telescope by a small shift compared to the detector size.

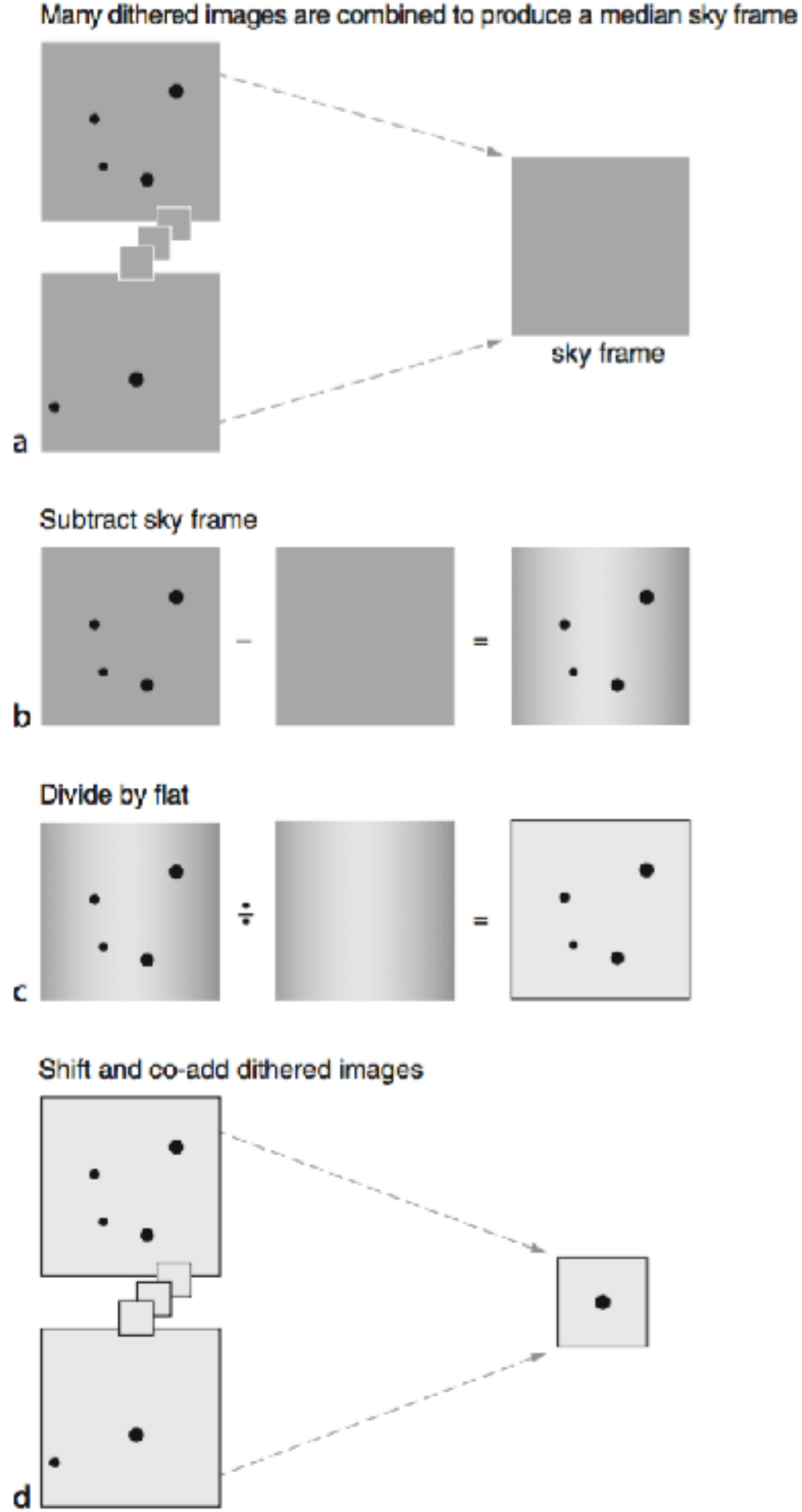
Using large offsets comparable to the size of the detector to increase the field-of-view is normally referred to as mosaicking.

PROS:

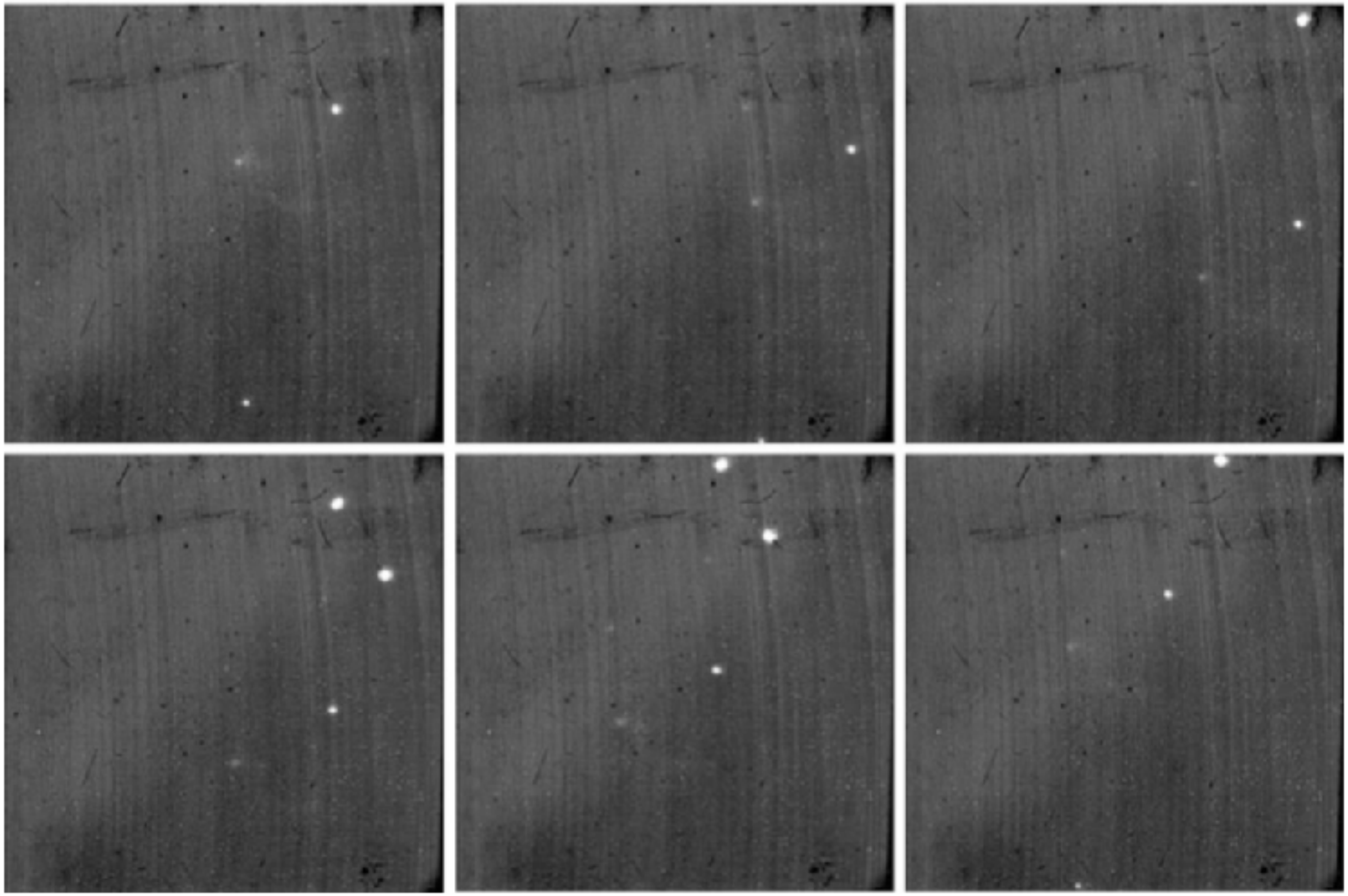
- This places the target on different locations on the detector, which helps to increase the quality of the combined image. Moving the detector an artifact-specific distance makes it possible to remove defects or non-optimal pixels.
- These could be individual hot or dead pixels, bad rows or columns, other defects.
- Also, effects of pixel-to-pixel errors in the flat-field or pixel-to-pixel sensitivity over the detector are reduced. For detectors with a chip gap, dithering can be used to fill the gap in order to produce a complete image.
- Using sub-integer dithers helps improve the spatial sampling of the point-spread function (PSF), which is particularly important for HST since the PSF is under-sampled in most of the instruments.

CONS:

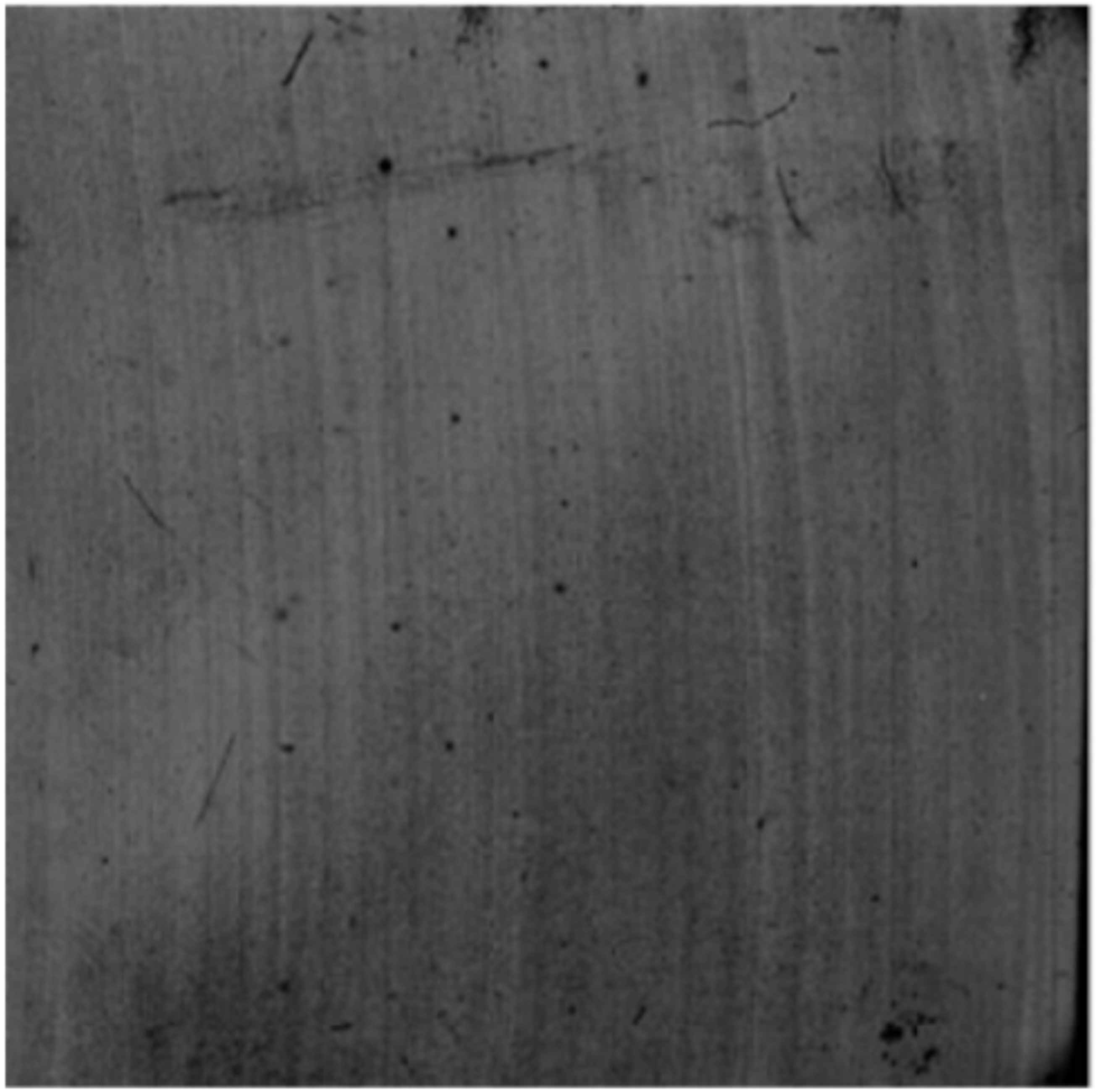
- Executing the dithers with adequate accuracy (at a sub-pixel level) may be difficult or impose overheads.
- Splitting exposures may lead to additional noise (e.g., readout noise in cases where the images are detector rather than sky noise limited) and dithering inevitably leads to a smaller final field of the deepest imaging.
- Finally the precise measurement of shifts between images and the reconstruction of the final co-added image can be time consuming.


Many dithered images are combined to produce a median sky frame
sky frame
a
Subtract sky frame
−
=
b
Divide by flat
÷
=
c
Shift and co-add dithered images
d

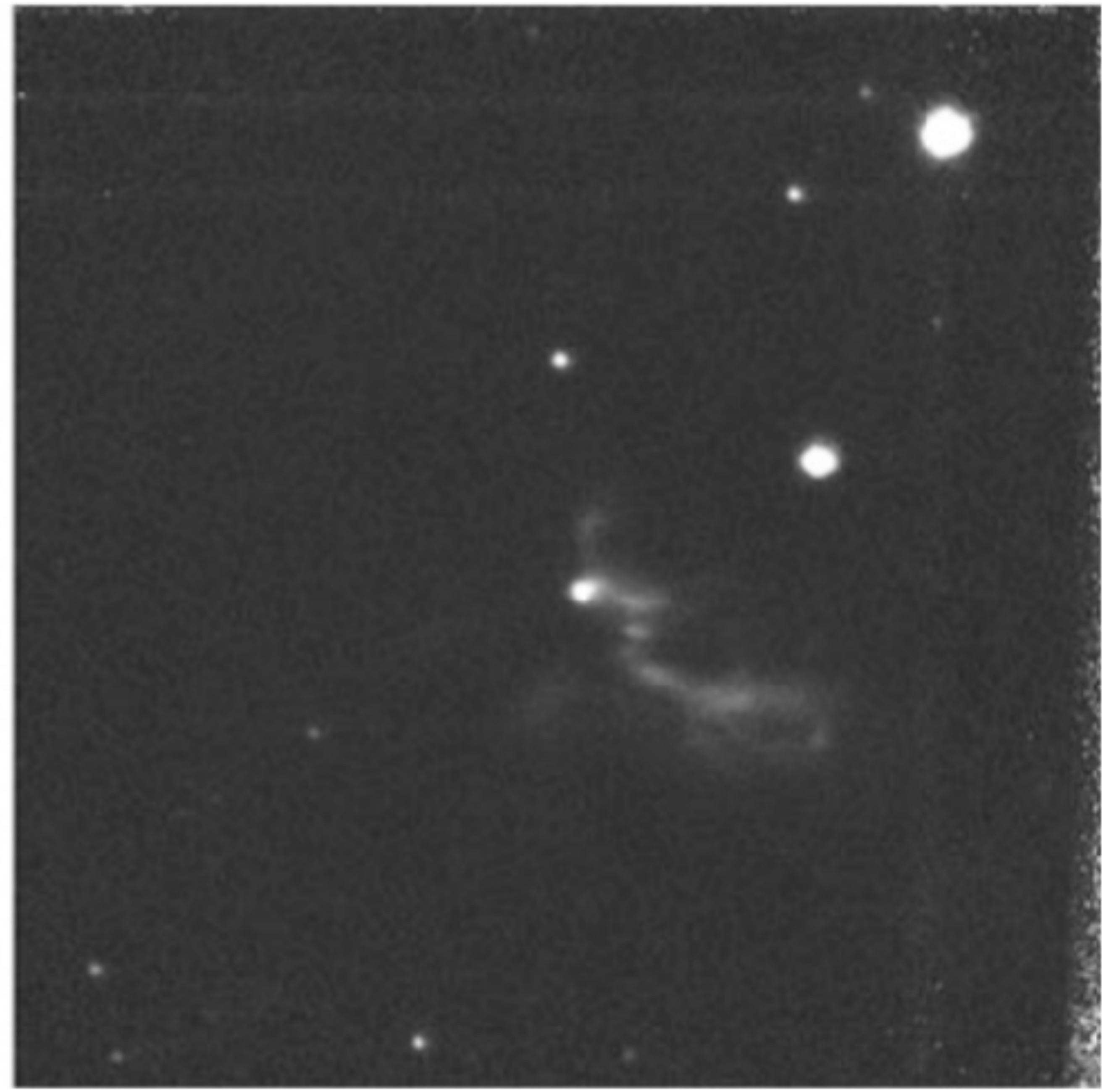
Many dithers



Sky flat

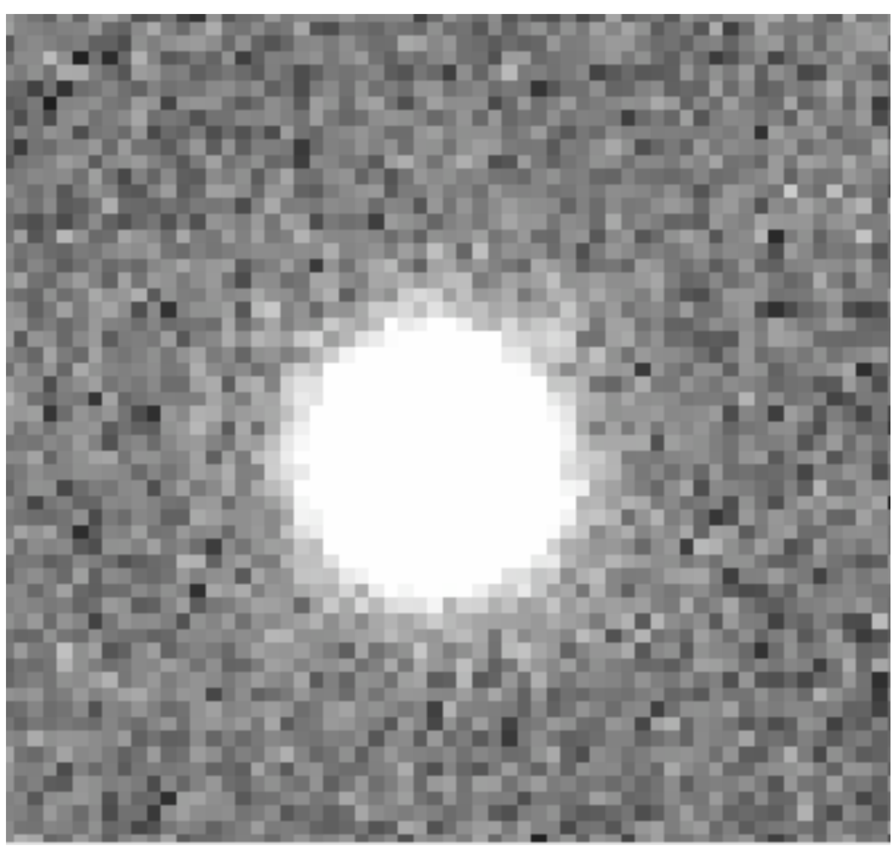


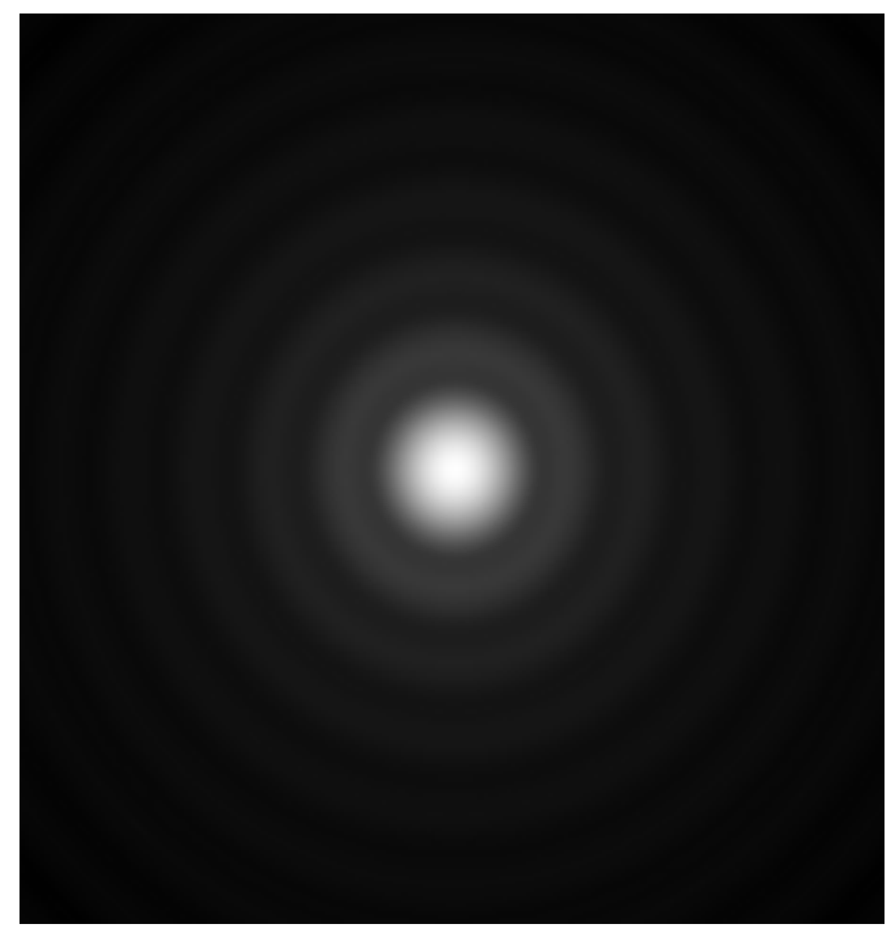
final image



Photometry: Basic Questions

- How do you identify objects in your image?
- How do you measure the flux from an object?
- What are the potential challenges?
- Does it matter what type of object you're studying?

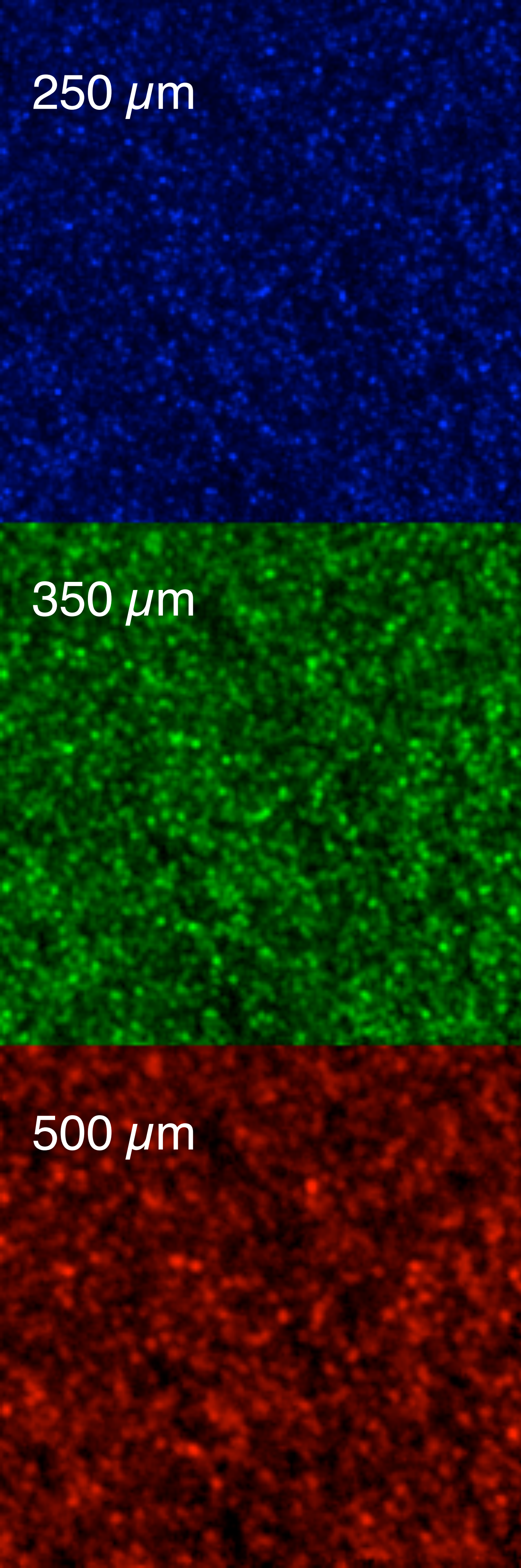


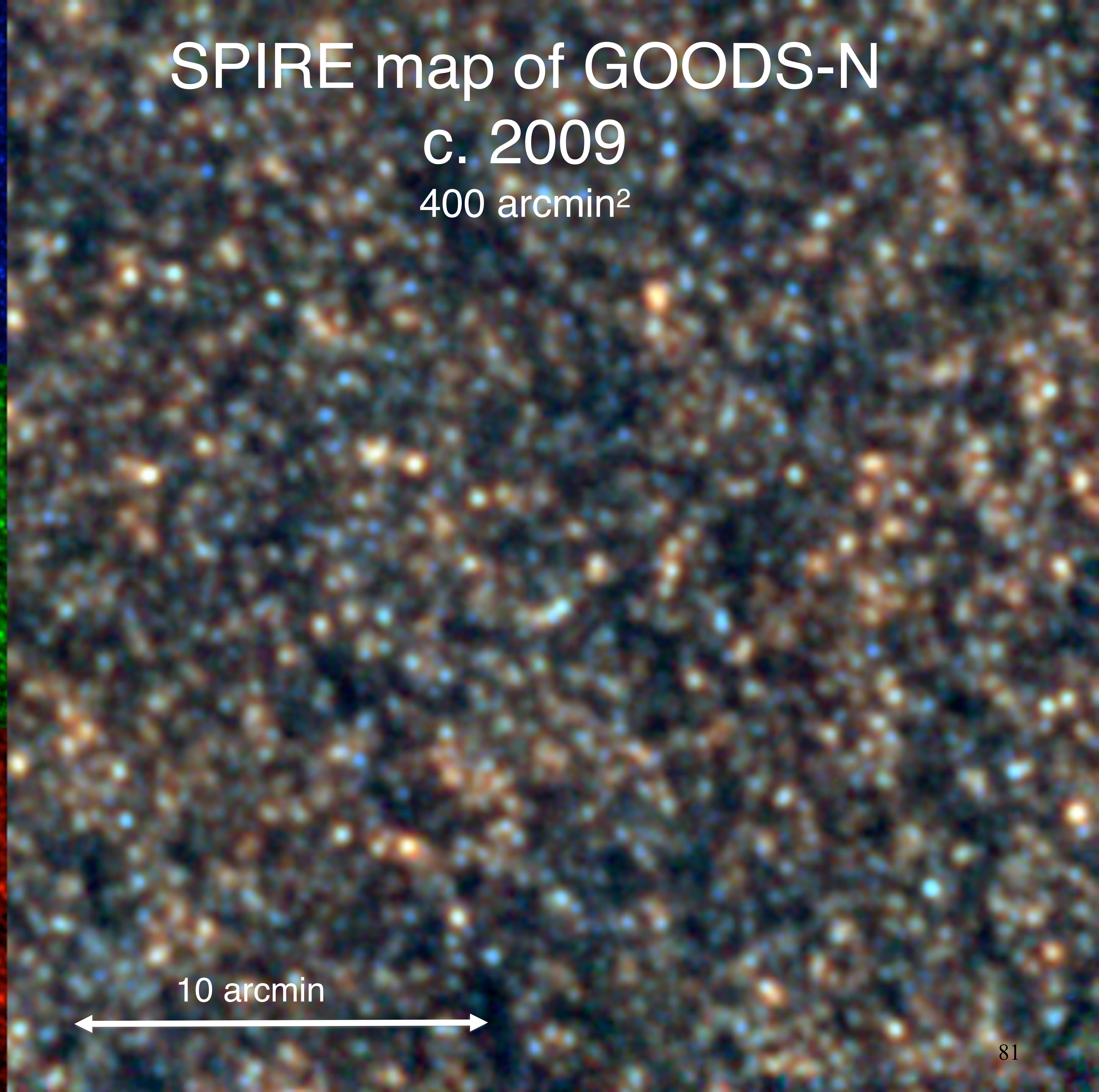


How do you mathematically define:

- Where there’s an object?
- What it’s flux is ?

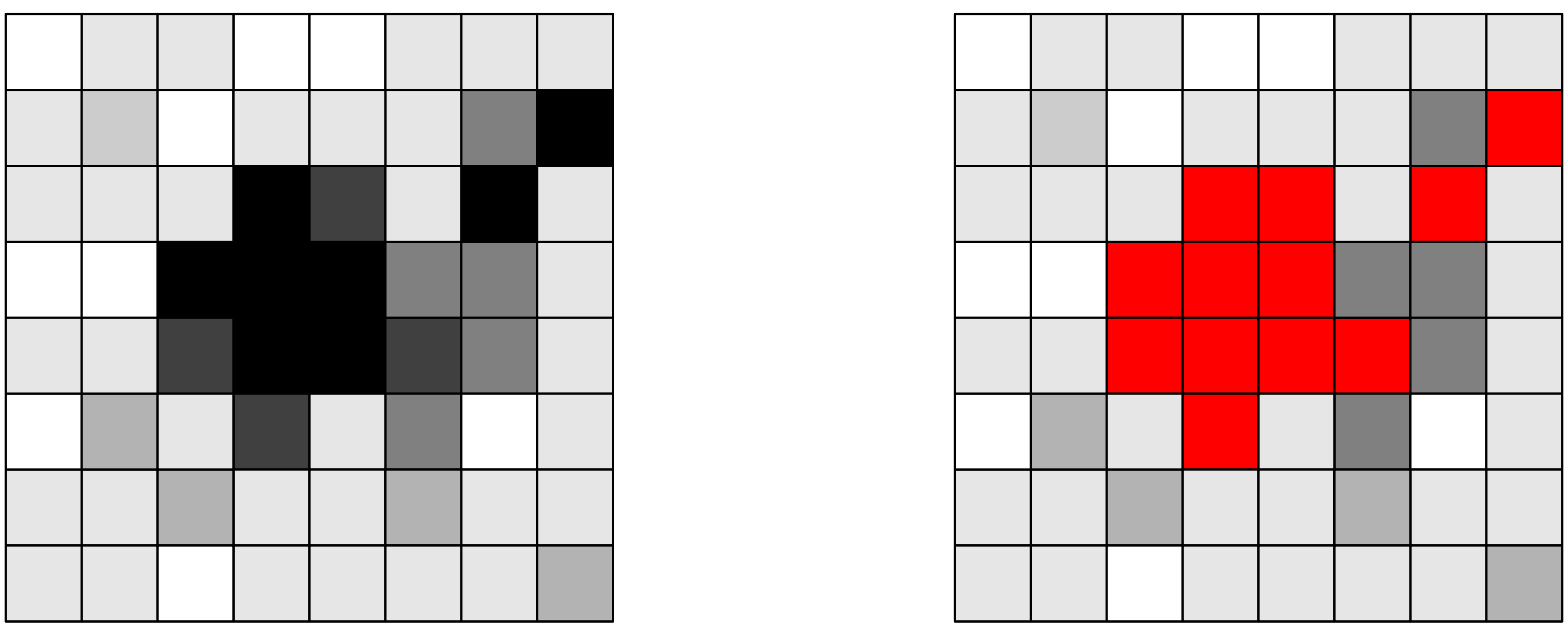
SPIRE map of GOODS-N
c. 2009
400 arcmin^2





I: General Considerations

- Object Detection
  - Define a *detection threshold* and *detection area*. An object is only detected if it has N pixels above the threshold level.
  - One simple example of a detection algorithm:
    - Generate a *segmentation image* that includes only pixels above the threshold.
    - Identify each group of contiguous pixels, and call it an object if there are more than N contiguous pixels



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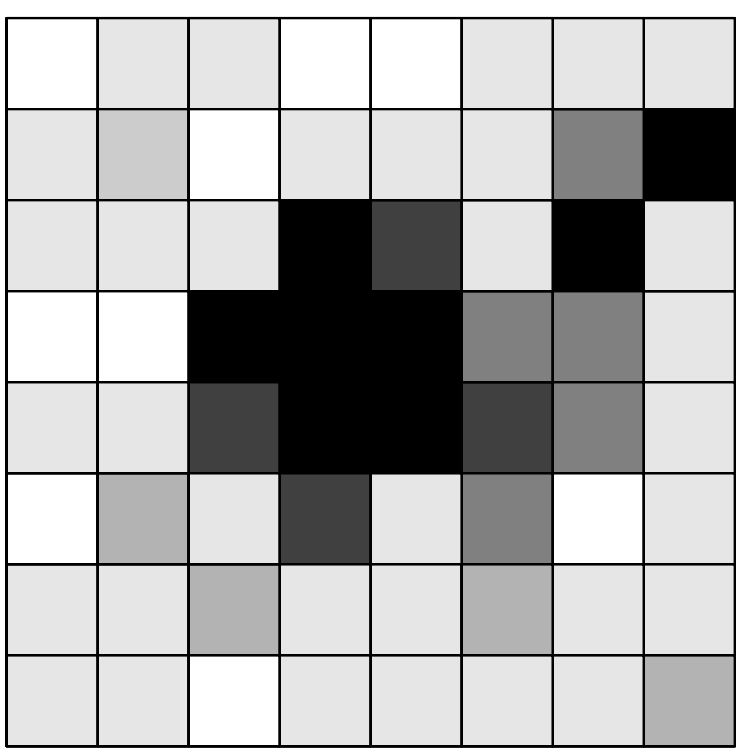
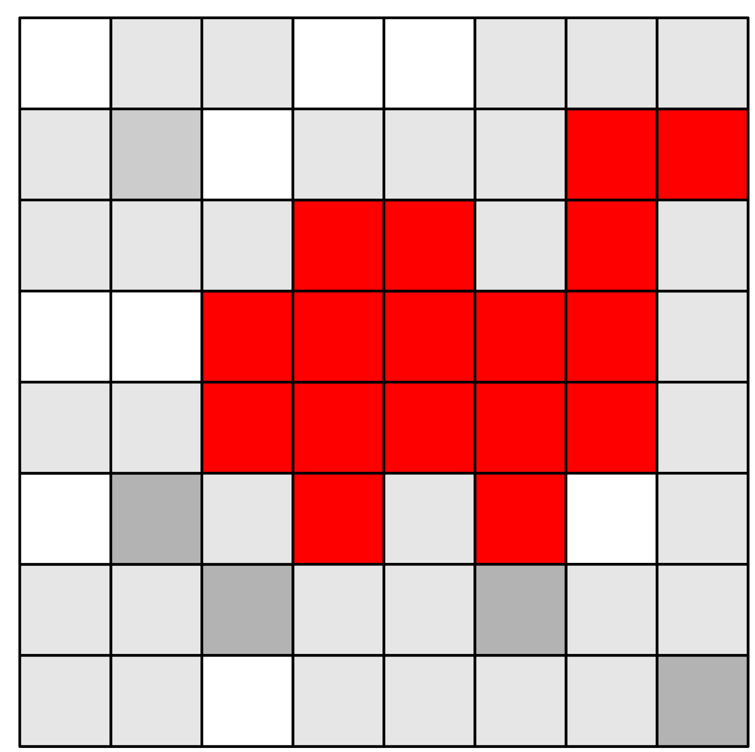
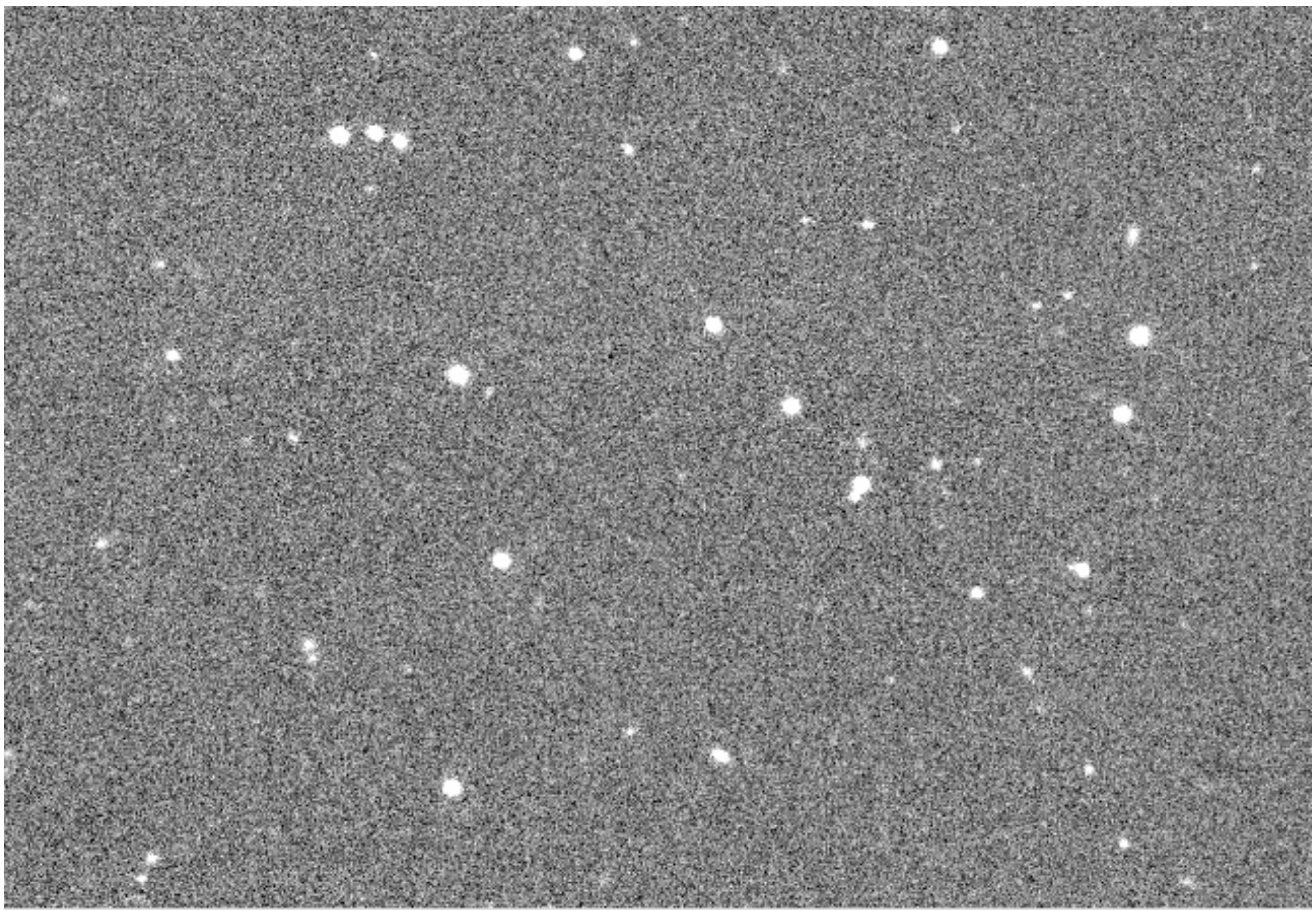
    - Identify each group of contiguous pixels, and call it an object if there are more than N contiguous pixels





I: General Considerations

- Object Detection



I: General Considerations

- Object Detection

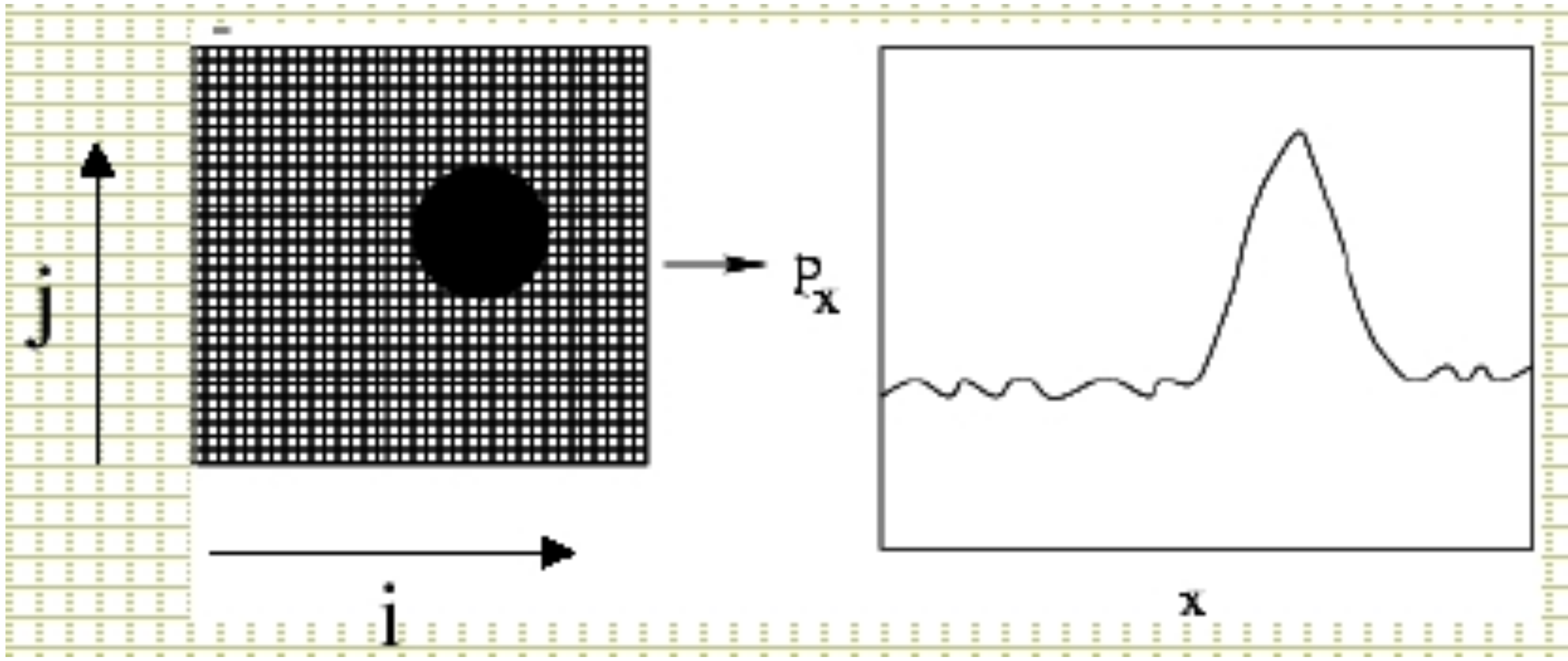
Measuring the Position

- Centroiding

How do you determine the centroid of an object?

Consider an image with flux levels I(i,j) in pixel i,j. The ***marginal distribution*** along a given axis is obtained by extracting a subsection of the image and summing along the row or columns.

Note that this is not the only way to find the centroid.



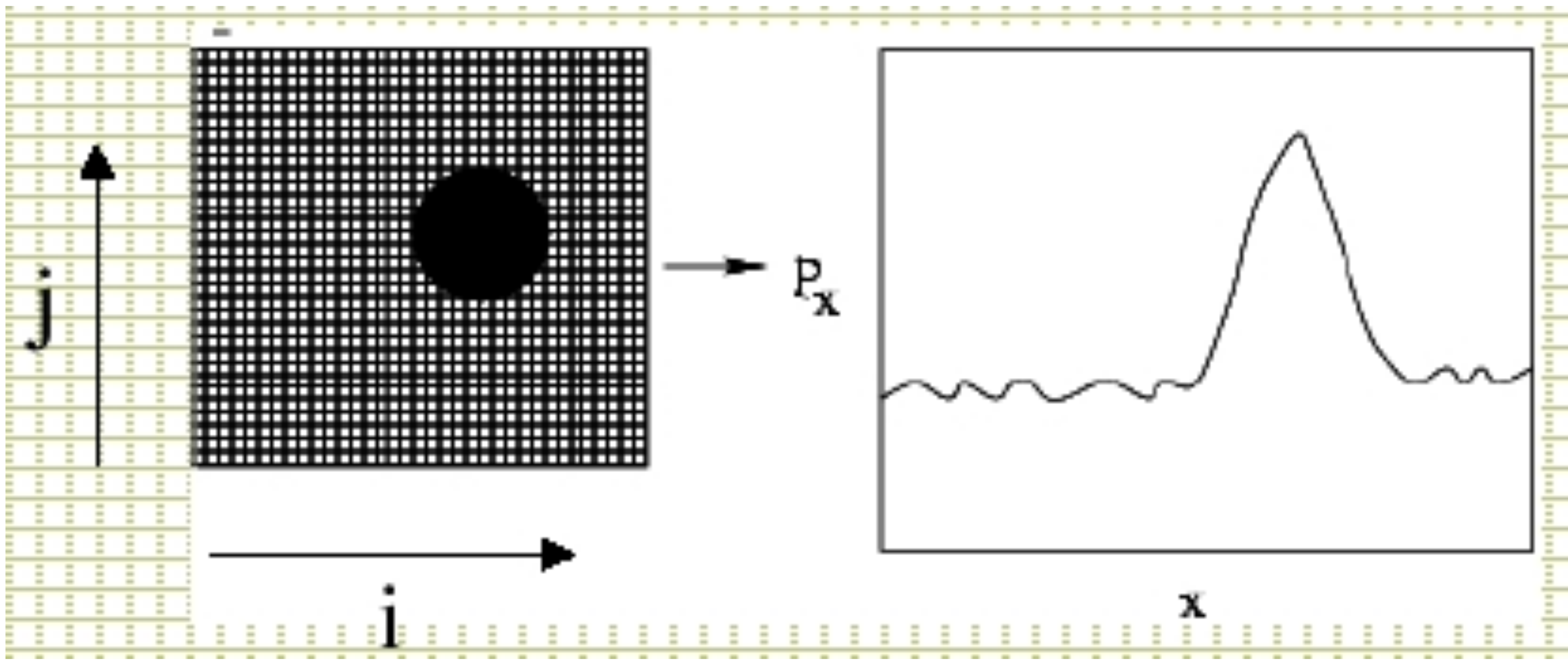
Centroiding

- Centroiding: Marginal Distribution
  - Step 1: Sum the pixel values I_{ij} along the 2N+1 rows and columns around the object.

$$P_{x_i} = \sum_{j=-N}^{N} I_{ij}$$

$$P_{y_j} = \sum_{i=-N}^{N} I_{ij}$$

These are the marginal distributions.



Centroiding

- Centroiding: Marginal Distribution
  - Step 2: Determine an intensity-weighted centroid

$$P_{x_i} = \sum_{j=-N}^{N} I_{ij}$$

$$x_{cen} = \frac{\sum_i x_i \cdot P_{x_i}}{\sum_i P_{x_i}}$$

$$P_{y_j} = \sum_{i=-N}^{N} I_{ij}$$

$$y_{cen} = \frac{\sum_j y_j \cdot P_{y_j}}{\sum_j P_{y_j}}$$

Centroiding

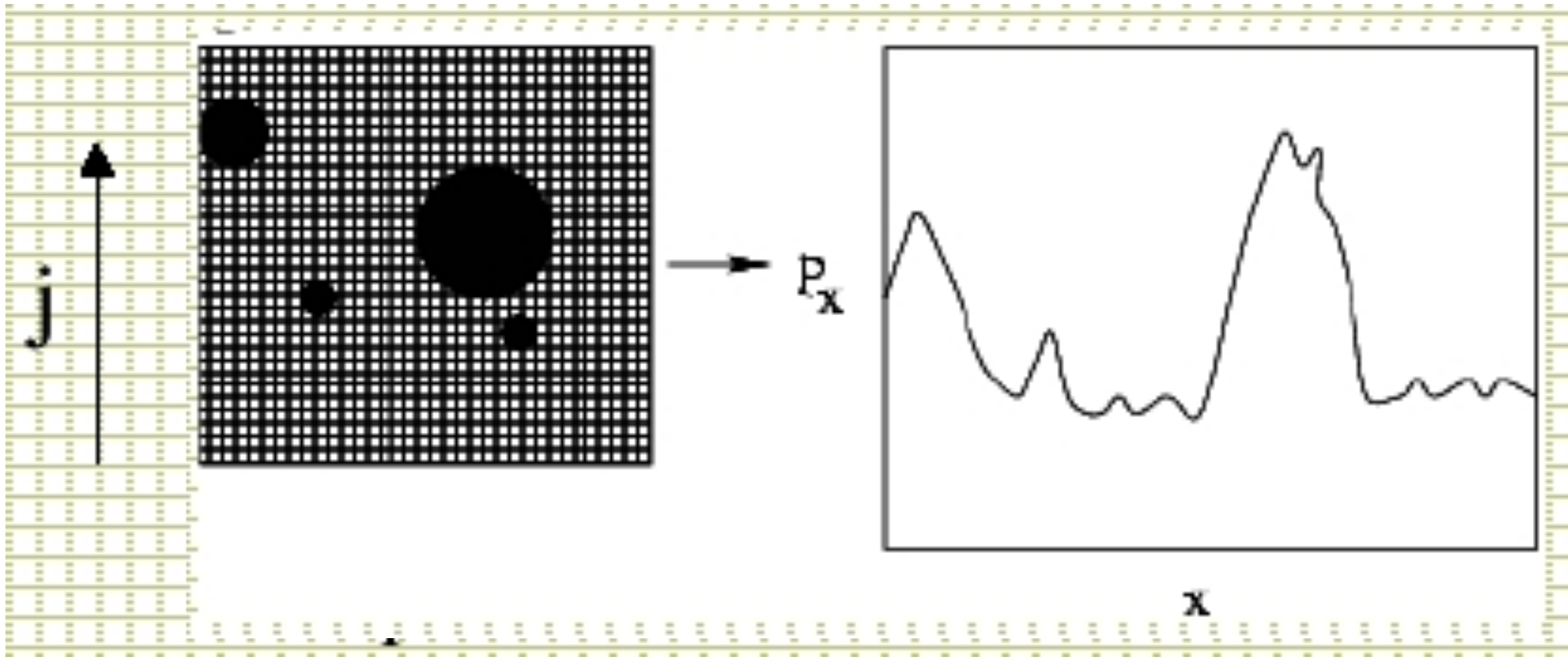
- Centroiding: Marginal Distribution
  - Uncertainties in the centroid locations

$$x_{cen} = \frac{\sum_i x_i \cdot P_{x_i}}{\sum_i P_{x_i}} \qquad \sigma_x^2 = \frac{\sum_i (x_i - x_{cen})^2 \cdot P_{x_i}}{\sum_i P_{x_i}}$$

$$y_{cen} = \frac{\sum_j y_j \cdot P_{y_j}}{\sum_j P_{y_j}} \qquad \sigma_y^2 = \frac{\sum_j (y_j - y_{cen})^2 \cdot P_{y_j}}{\sum_j P_{y_j}}$$

Centroiding

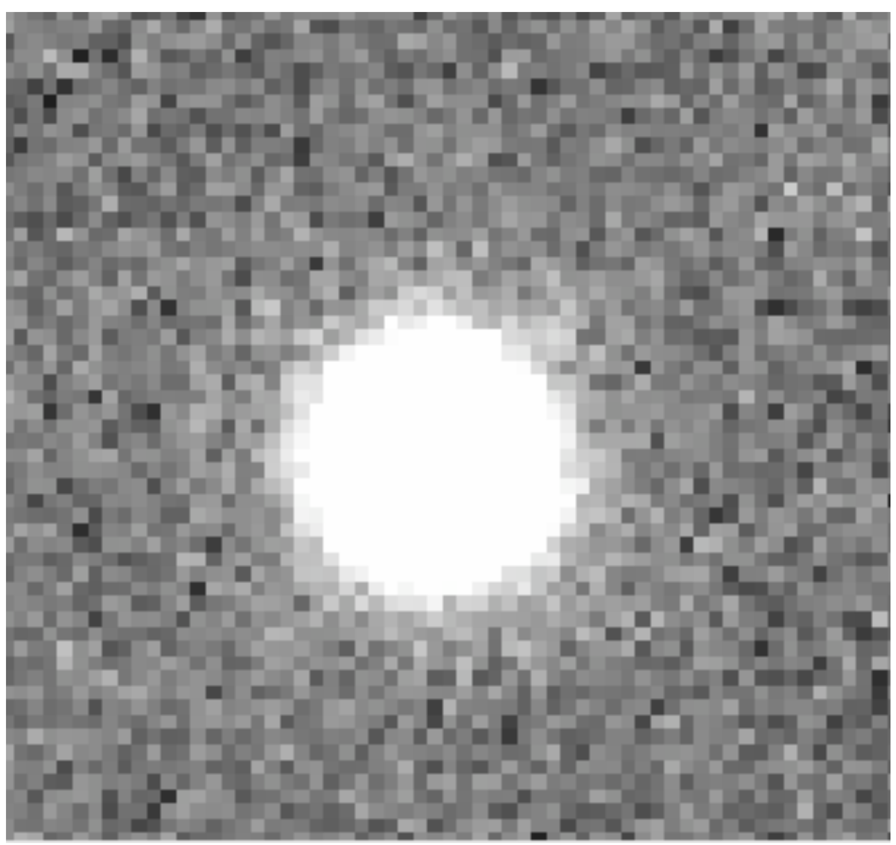
- Complication: Noise and multiple sources in image
  - Must decide what is a source and isolate sources (e.g. segmentation regions).
  - Compute the marginal distributions within isolated subregion.



Measuring Flux in an Image

- How do you measure the flux from an object?
- Within what area do you measure the flux?

The best approach depends on whether you are looking at resolved or unresolved sources.

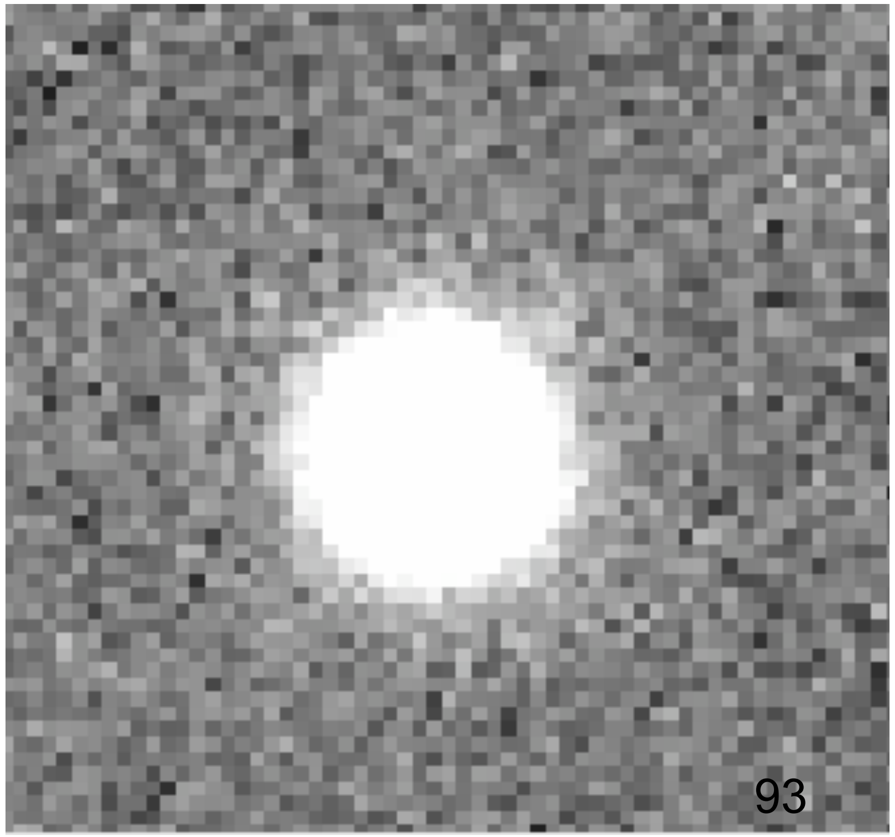


Background (Sky) Flux

- Background
  - The total flux that you measure (F) is the sum of the flux from the object (I) and the sky (S).

$$F = I + S = \sum I_{ij} + n_{pix} \cdot sky / pixel$$

  - Must accurately determine the level of the background to obtain meaningful photometry (We'll return to this a bit later.)



Photometric Errors

Issues impacting the photometric uncertainties:

- Poisson Error
  - Recall that the statistical uncertainty is Poisson in electrons rather than ADU. In ADU, the uncertainty is

$$\boxed{\sigma_{ADU} = \sqrt{ADU / Gain}}$$

- Crowded field contamination
  - Flux from nearby objects can lead to errors in either background or source flux

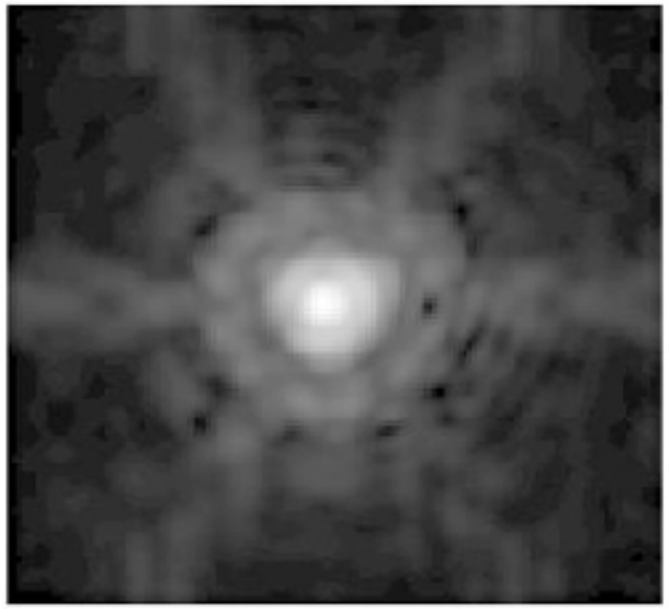
- Gradients in the background sky level

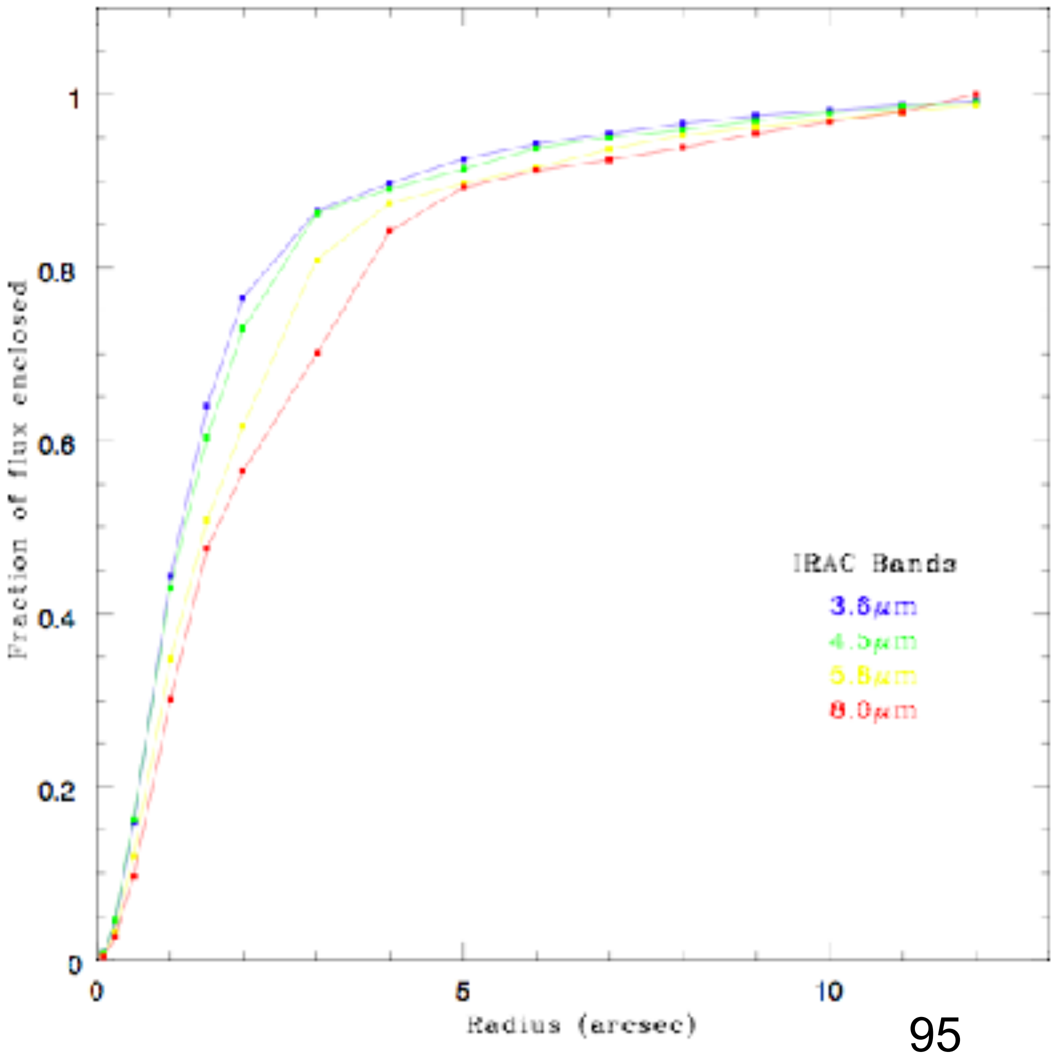
- Correlated pixel statistics
  - Interpolation when combining images leads the uncertainties to be non-Poisson because the pixels are correlated.

II. Stellar Photometry

- Stars are unresolved point sources
  - Distribution of light determined purely by point spread function (PSF)

- "Curve of Growth"
  - Radial profile showing the fraction of total light within a given radius



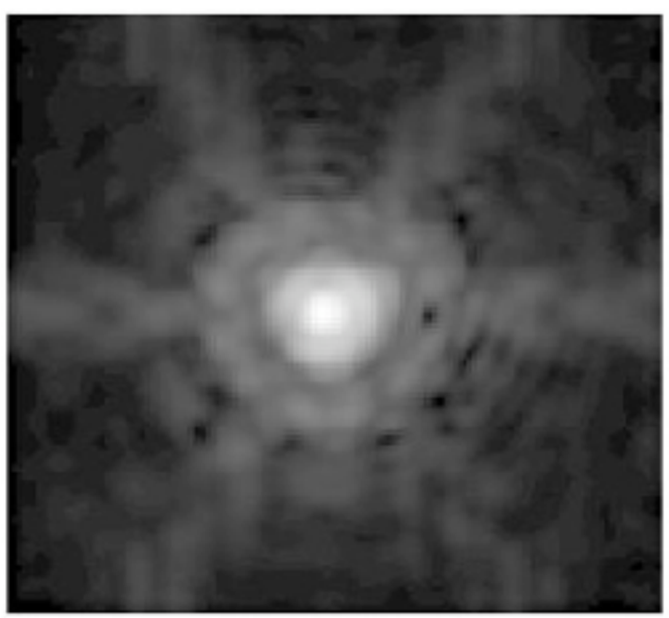


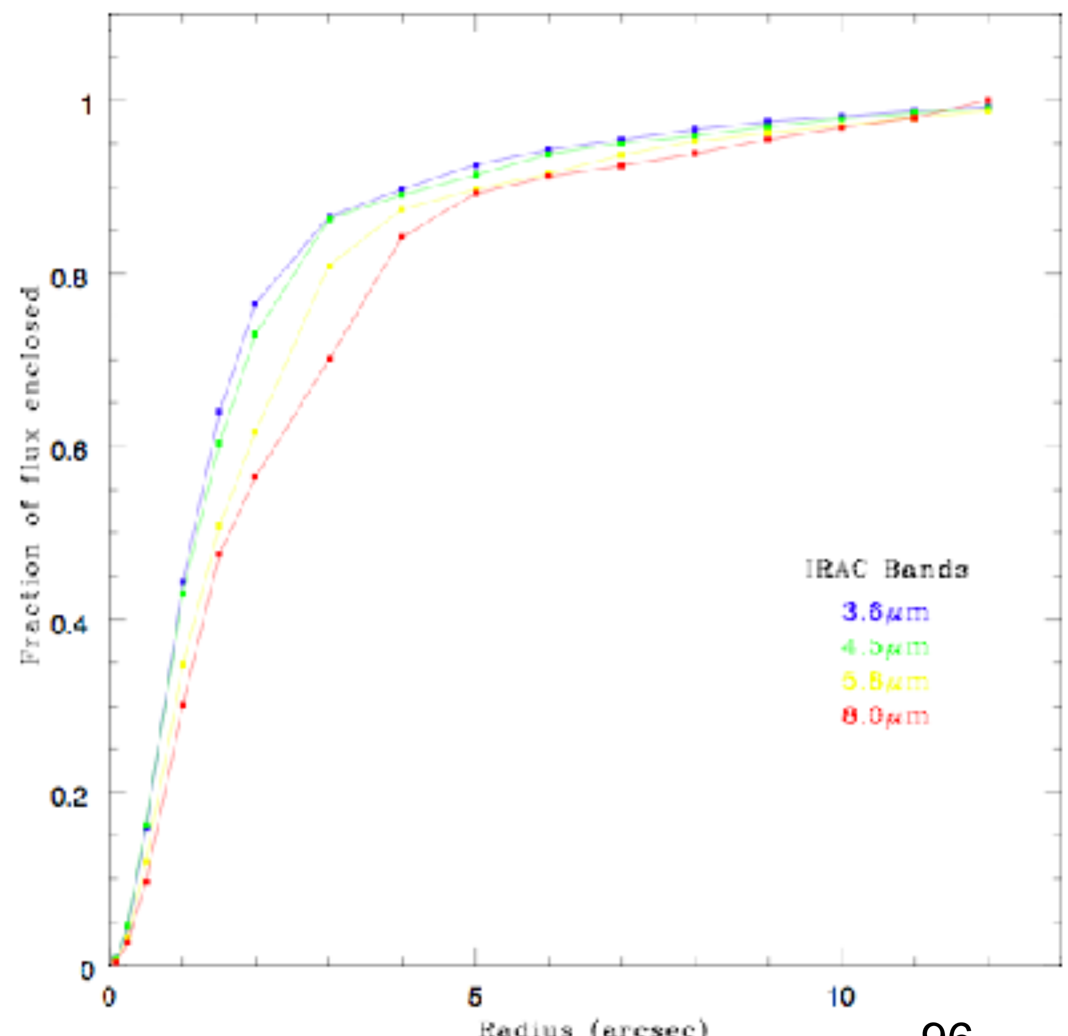
http://ssc.spitzer.caltech.edu/irac/psf.html

II. Stellar Photometry

- Stars are unresolved point sources
  - Distribution of light determined purely by point spread function (PSF)
  - How do you measure the light?

  Options:
  - Aperture photometry
  - PSF fitting

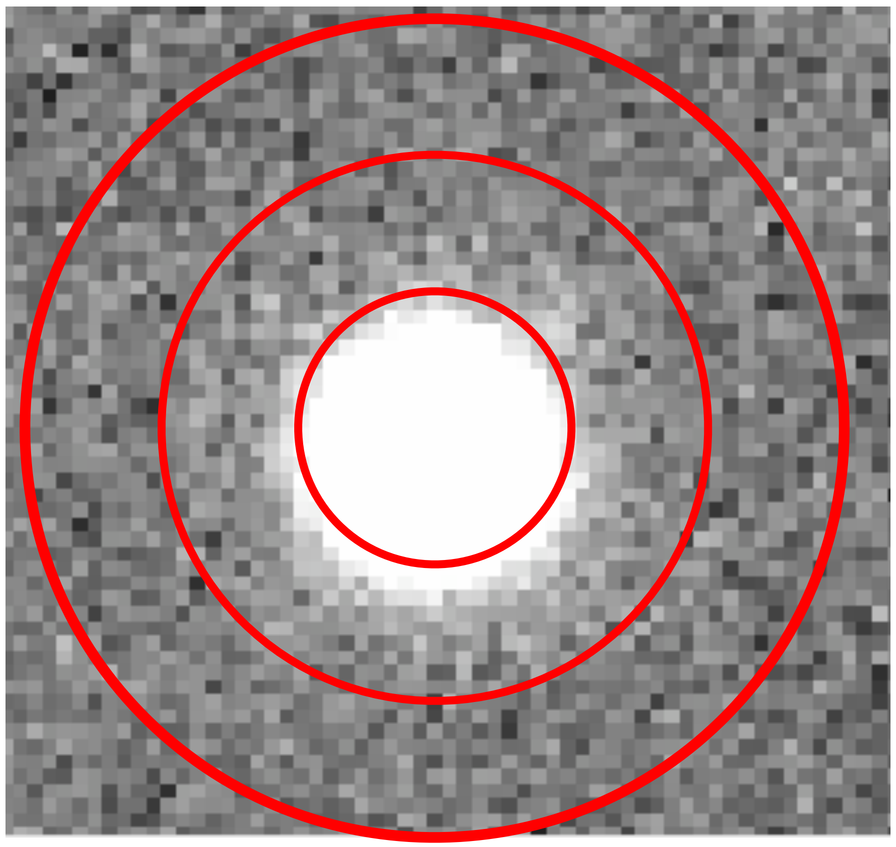


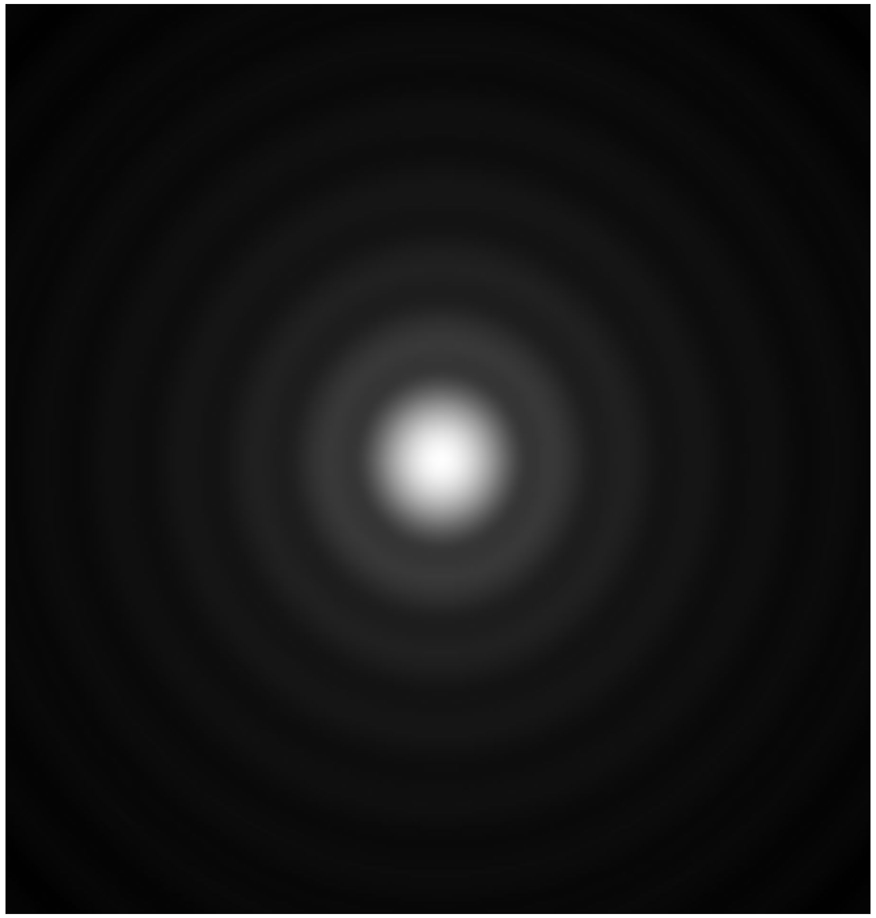


http://ssc.spitzer.caltech.edu/irac/psf.html

II. Stellar Photometry

- Aperture Photometry:
  - Measure the flux within an pre-defined (typically circular) aperture.
  - Can calibrate as long as you use the same aperture for your standard star.
  - Can compute total flux if you know curve of growth.



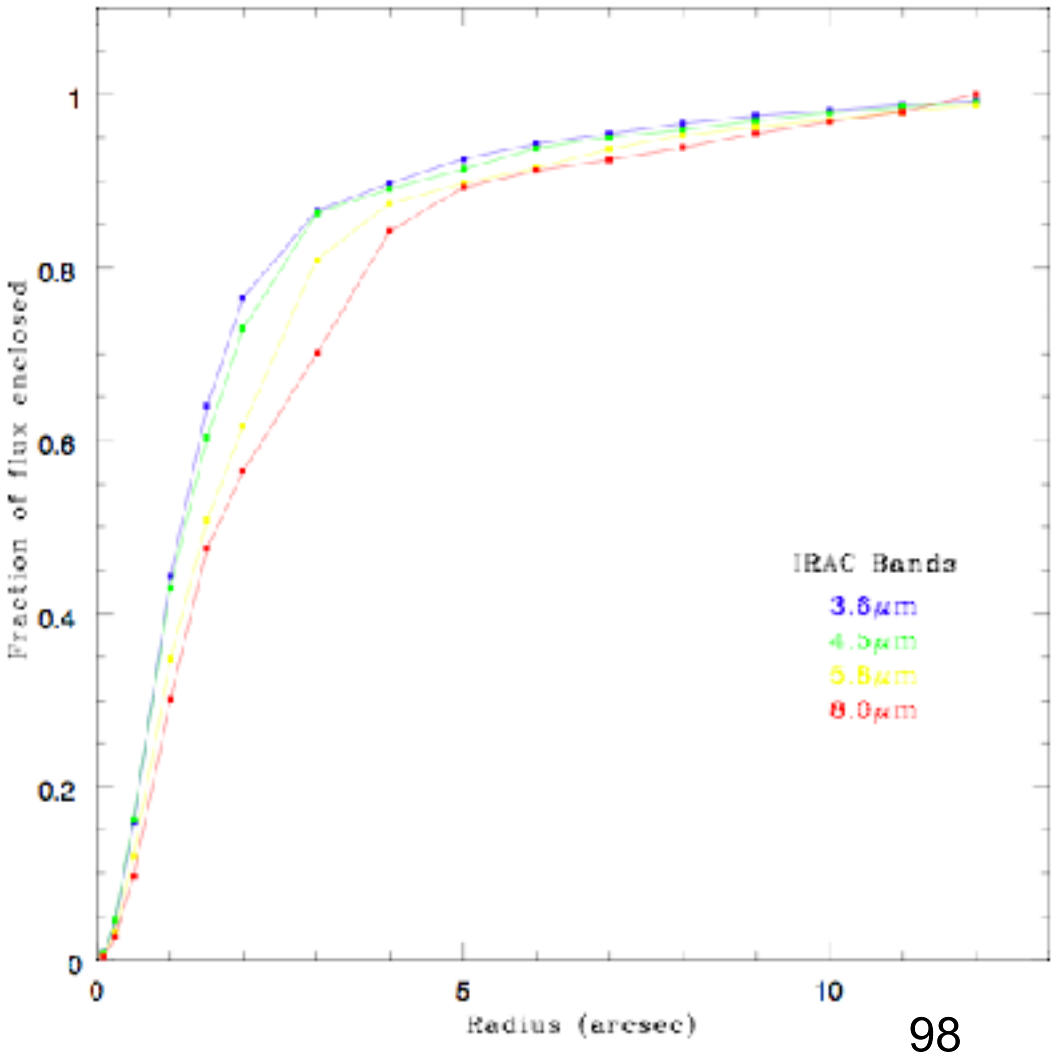


What are the potential drawbacks?

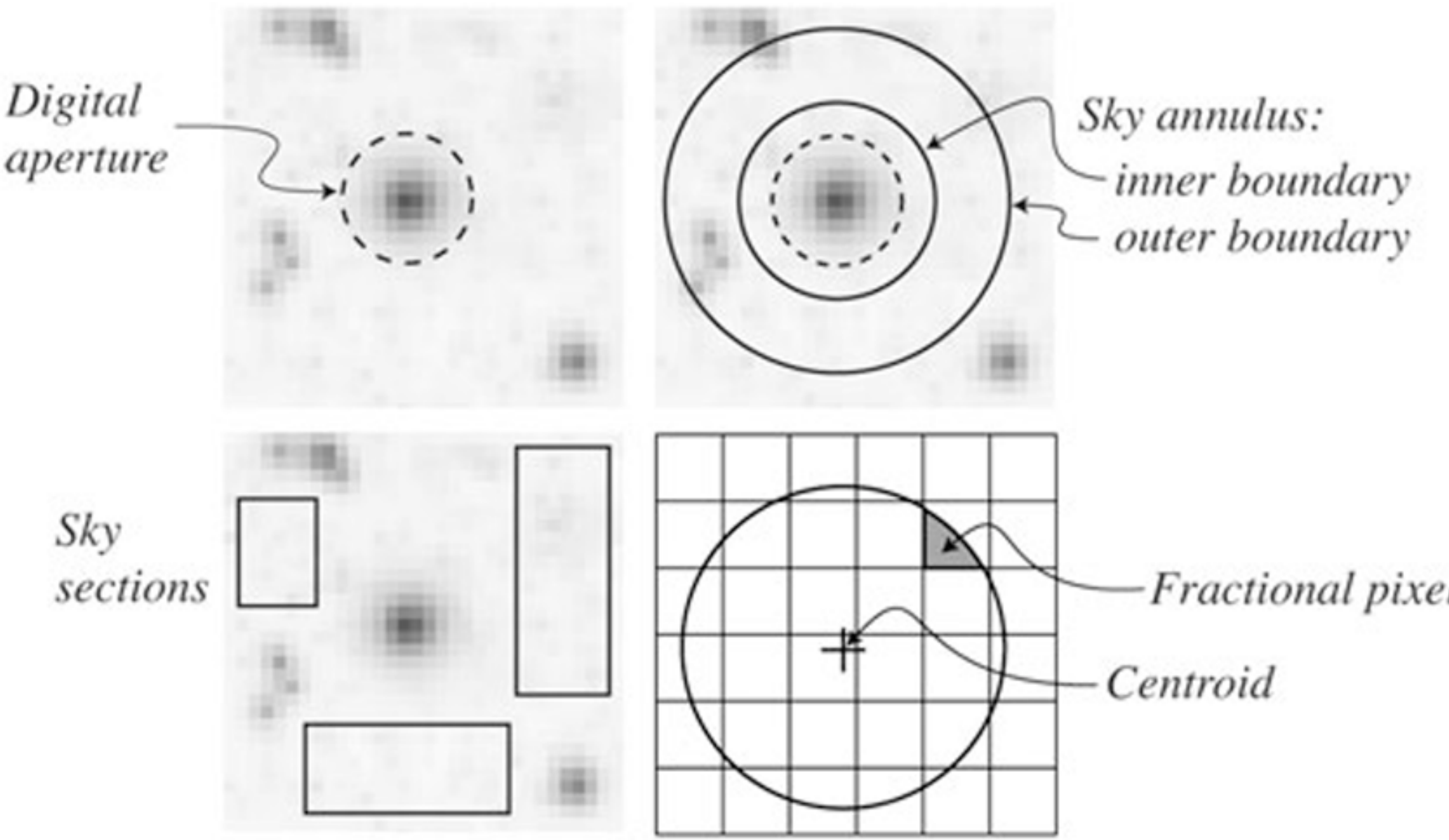
II. Stellar Photometry

- Stars are unresolved point sources
  - Distribution of light determined purely by point spread function (PSF)
  - How do you measure the light?

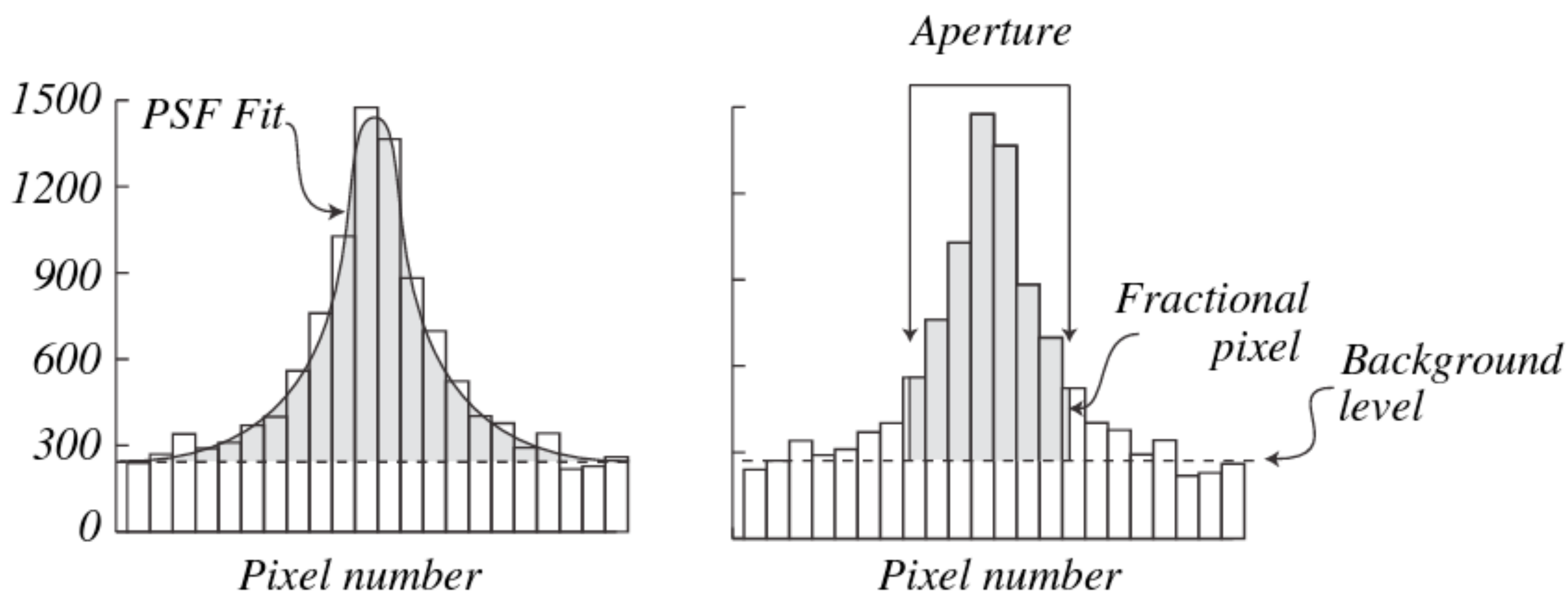
- “Curve of Growth”
  - Radial profile showing the fraction of total light within a given radius



aperture photometry



aperture photometry



II. Stellar Photometry

- PSF fitting:
  - Determine the form of the PSF and then fit the amplitude to all the stars in the image.
  - Can use an empirically constructed PSF or an analytic parameterization
  - Typical parameterizations of PSF
    - Gaussian

      $I(r) = \exp(-0.5 * (r/\sigma)^2)$

      $F(r) = 1 - \exp(-0.5 * (r/\sigma)^2)$

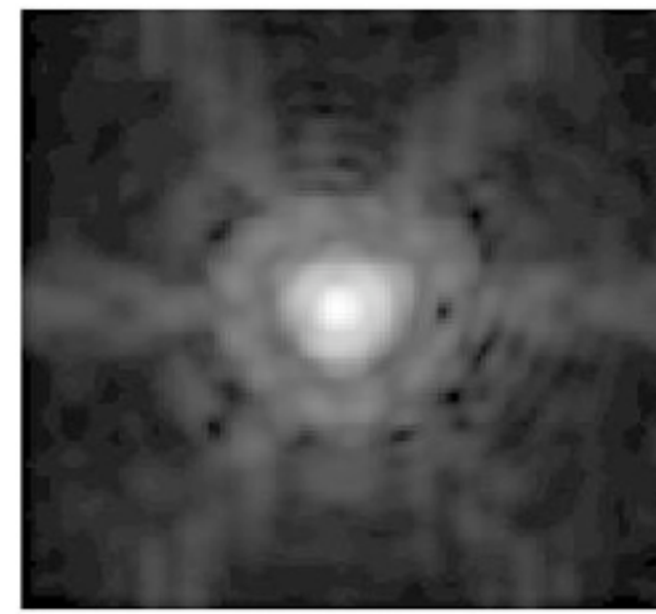
      $FWHM = 2\sigma * \mathrm{sqrt}(2 * \ln(2))$

    - Moffatt

      $I(r) = (1 + (r/\alpha)^2))^{-\beta}$

      $F(r) = 1 - (1 + (r/\alpha)^2))^{(1-\beta)}$

      $FWHM = 2\alpha * \mathrm{sqrt}(2^{1/\beta} - 1)$



where I(r) is the intensity profile and F(r) is the enclosed flux profile. F(r) is typically what is fit to determine the best parameters. The FWHM formulae correspond to what you would see in IRAF using imexam.

PSF-fitting

- **PSF-fitting:** Also called profile fitting. This method relies on modeling the image rather than summing over the image. The stellar image is usually compared to a Gaussian profile:

$$I(r) = I(0)\, e^{-r^2/2\sigma^2}$$

- where e = 2.718 is the base of natural logarithms, *I*(0) is the peak intensity and *r* is the radial distance from the center of the image. The quantity *σ* measures the width of the distribution; 68% of the light lies within ± 1σ and 98.7% of the lies within ±2.5σ.
- Provided the PSF (i.e. σ) is constant across the image, programs like DAOPHOT (Stetson 1987) will identify the bright stars, deduce their Gaussian profiles and subtract those profiles away, thereby revealing fainter stars. Other mathematical functions can also be used. It is often more convenient to describe the point spread function in terms of its Full Width at Half of the Maximum intensity; the FWHM (in pixels or arcseconds) is related to σ of a Gaussian by the simple equation

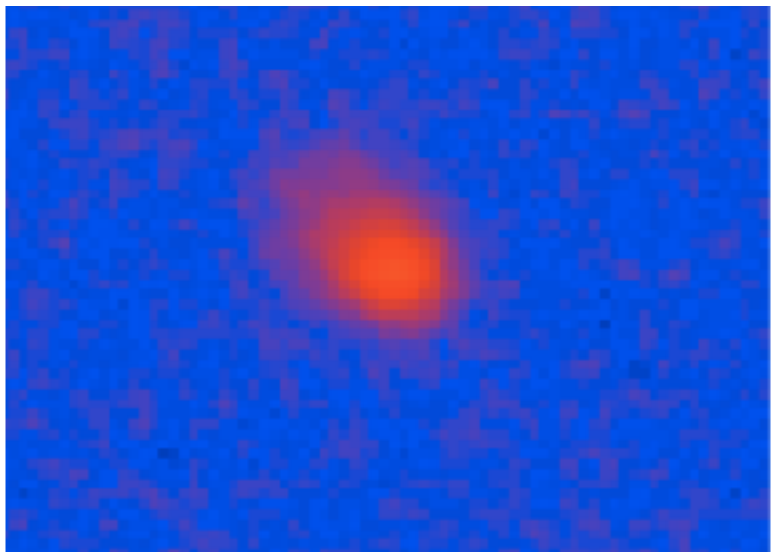
$$FWHM = 2.35\sigma$$

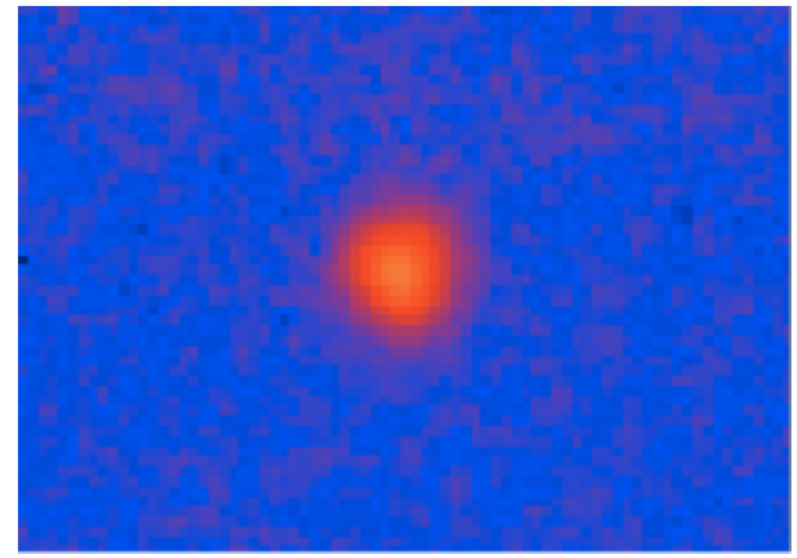
(McLean 2008)

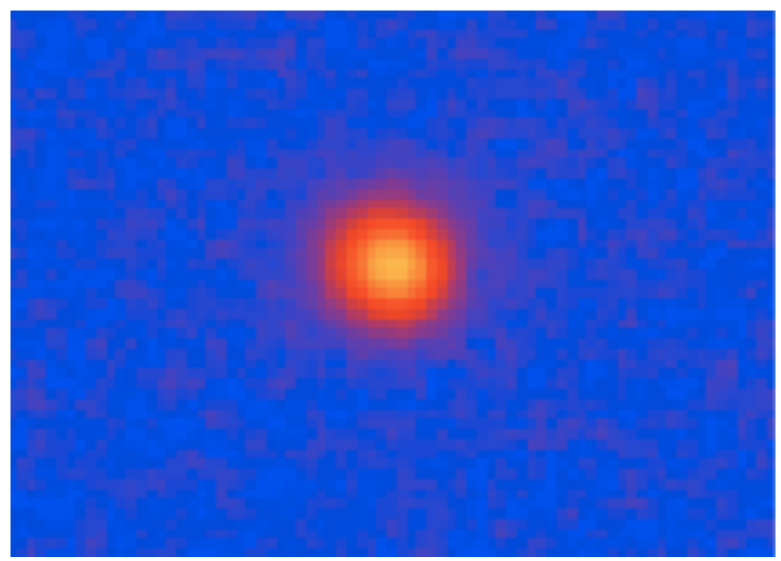
II. Stellar Photometry

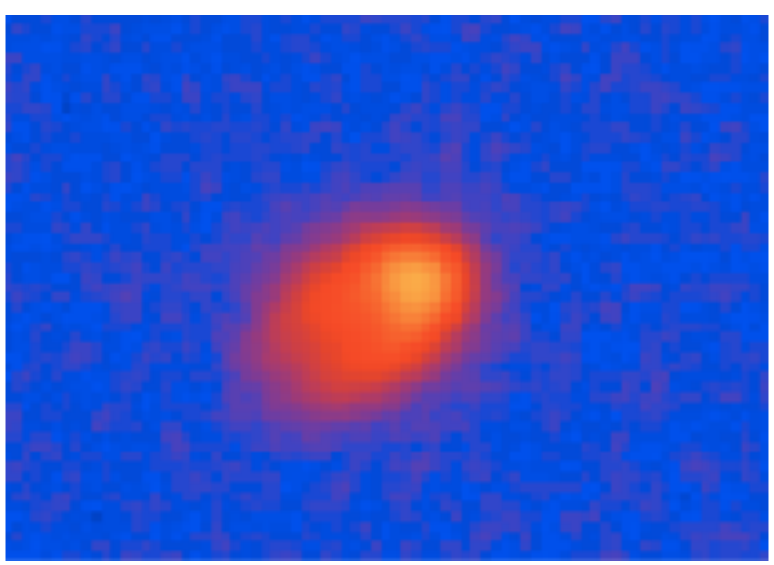
- PSF fitting:
  - Advantages:
    - Still works in crowded fields (can fit the center)
    - Regions with highest S/N have most weight in determining fit
    - Background is included as one additional parameter (constant in the fit)
  - Potential problems:
    - The PSF is not well described by the parametric profiles.
    - The PSF varies across the detector.

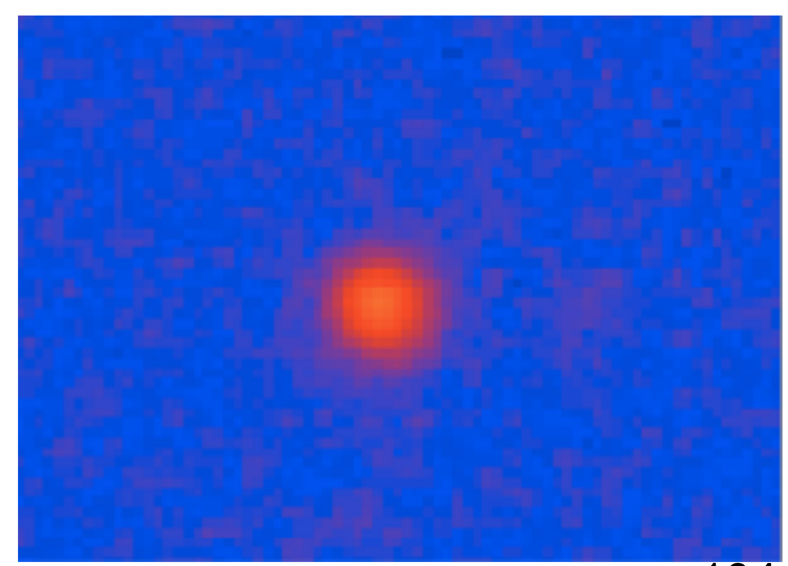
II. Stellar Photometry











II. Stellar Photometry

- Potential problems:
  - The PSF is not well described by the parametric profiles.
  - The PSF varies across the detector.

- Solutions:
  - PSF variations
    - Generate multiple PSF models for different parts of the detector and interpolate between these models

  - If parametric representation bad
    - Empirical PSF or include a non-parametric component in your PSF model
      - Use a very bright star
      - Fit the best psf model
      - In based upon parametric fit, keep a map of the residuals to correct for variations.

II. Stellar Photometry

- Determining Photometric Errors
  - Best approach: Artificial Star Tests
    - Basic idea - Insert a large number of fake stars into image and then obtain photometry for these objects.
    - Provides a direct measure of the scatter between true and observed magnitudes
    - Caveat: Requires that you have a good model for the PSF

Procedure for Stellar Photometry

- **Establishing the photometric system**
  - Set the zero point (Johnson used six A0 stars averaged, but that was quickly forgotten and his work became the "Vega system"
  - Get accurate standard star measurements over the entire sky relative to the zero point defining stars
  - m = apparent magnitude = $-2.5 \log (f_{star}/f_{zero\ pt.})$
  - Then in terms of the network of well-measured standard stars,

    $m=-2.5 \log (f_{star}/f_{standard}) + m_{standard}$

A Happier Solution

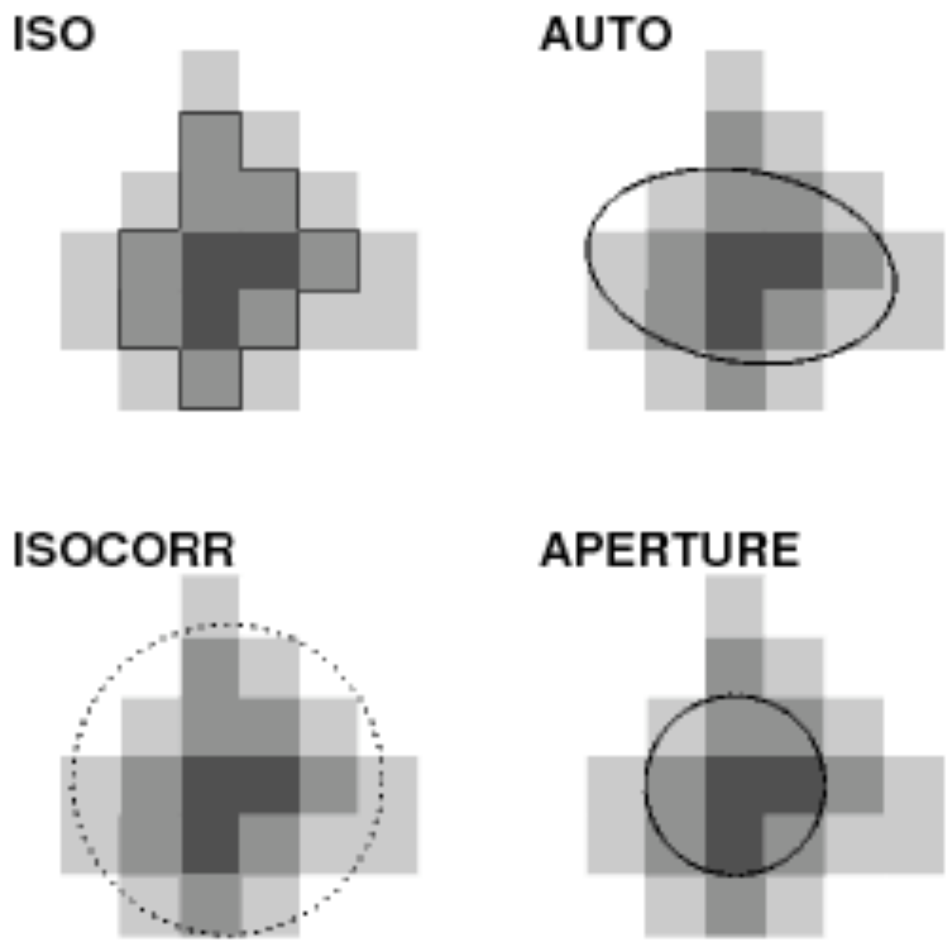
- **Use all-sky surveys**
  - Near infrared: 2MASS, accurate to about 2 – 3%
    - Not quite as good as it seems because of ~ 2% offsets between read 1 and read 2, plus some general calibration drift over the sky (but < 2%)
  - BV: Tycho, accurate to better than 1% (but sometimes gets confused by stars with small separations)
- **These surveys are as uniform over the whole sky as the best previous standard systems were over limited regions and with very small numbers of stars.**

- **It used to be necessary to take photometry of widely spaced standards, including at large air mass, and solve for the instrumental zero point plus the air mass corrections. Now there is an option of tying in directly with these all sky networks.**

III. Extended Source Photometry

5-10% accuracy generally considered decent for galaxies.

- Galaxies, HII regions, and many other astronomical objects are extended
  - Distribution of light determined by convolution of PSF and intrinsic shape
  - How do you measure the light?
  - How far out does the galaxy extend?



- Multiple Methods
  - Non-parametric
    - Aperture magnitudes
    - Isophotal magnitudes
    - Kron magnitudes
    - Petrosian magnitudes
  - Parametric
    - Assume profile for object

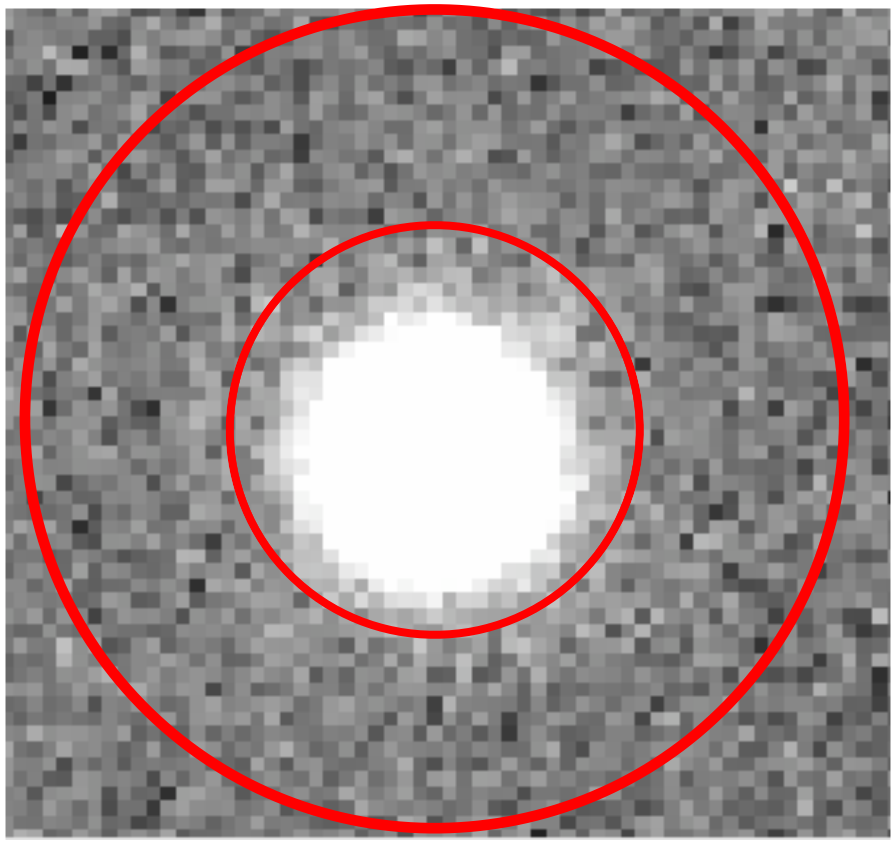
III. Extended Source Photometry

Parametric

1. Assume a parametric model for the object.
   Examples: Exponential disk, Disk+Bulge, Sersic Profile

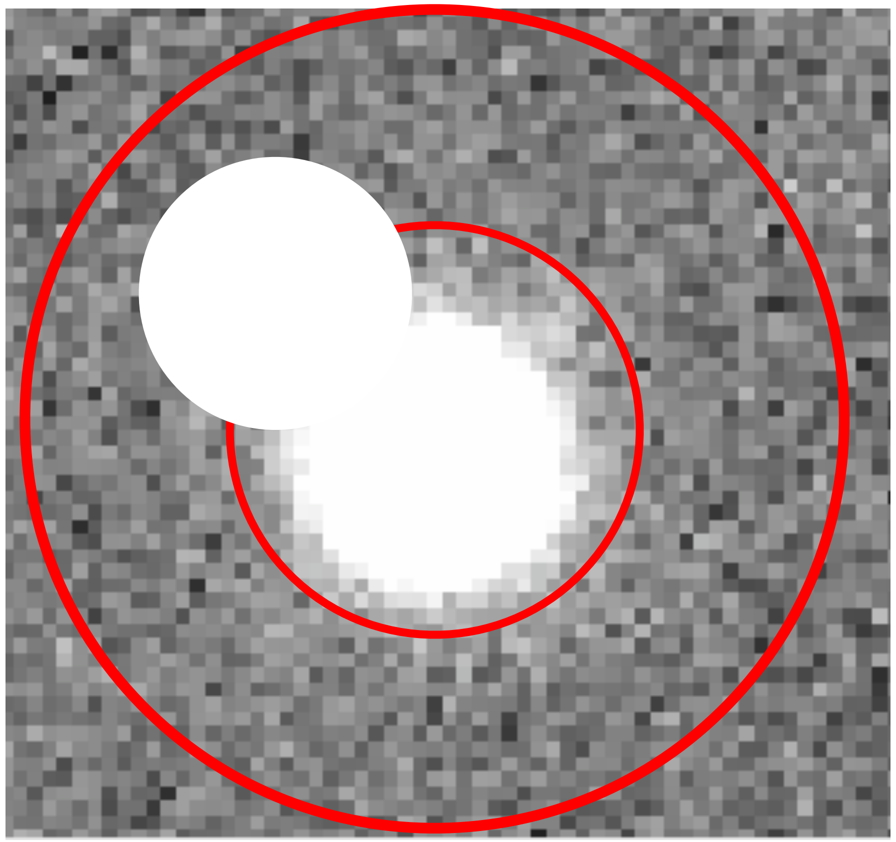
2. Perform a chi-squared minimization to obtain the best fit for the object
   Outputs will be position and model parameters, from which one can derive the total flux.

IV. More General Considerations



- What do you do if objects overlap?
- How/where do you determine the sky level?
  - Global (mean sky for image) or
  - Local (some annular region around object) ?
- How do you determine the uncertainty?
  - Do you have Poisson noise in the image?
    - If sky-subtracted, then you need to know what the original sky level was.
    - If N frames have been averaged, then you need to account for this
    - If pixels are correlated (i.e. smoothed data), then most codes will significantly underestimate the errors.

IV. More General Considerations



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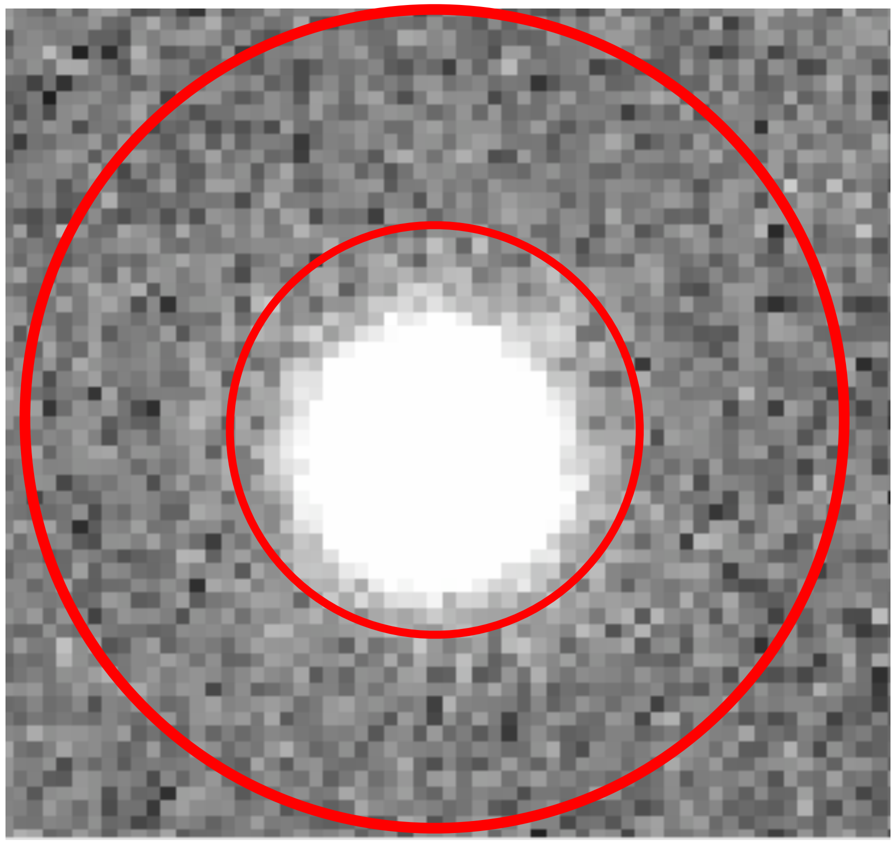

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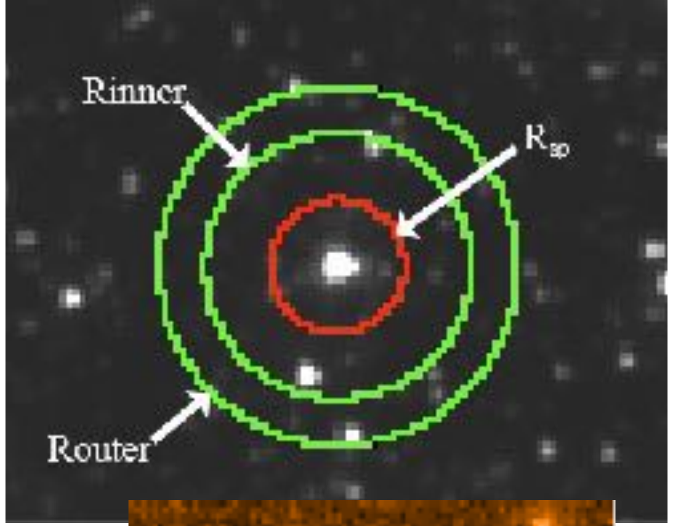
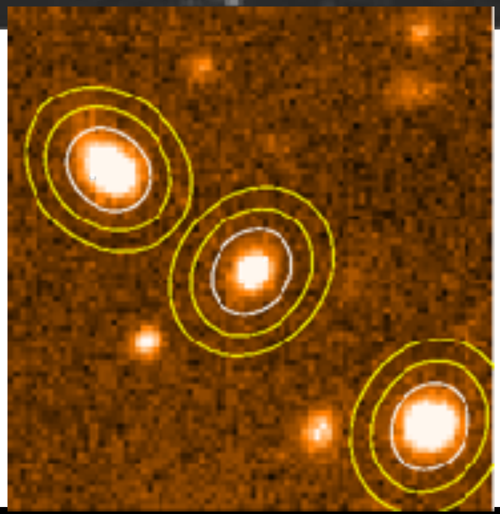
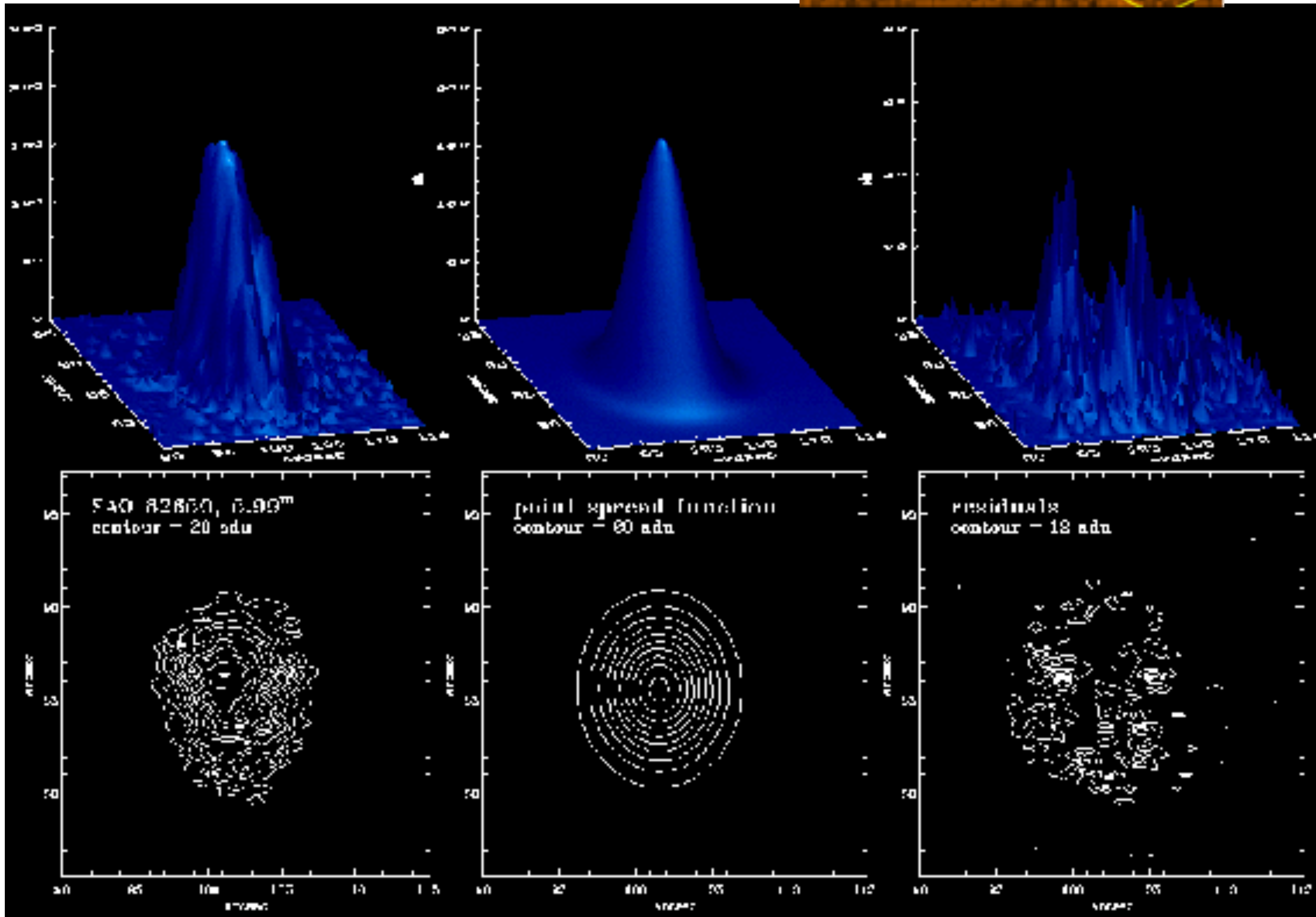
Measurements with Detector Arrays



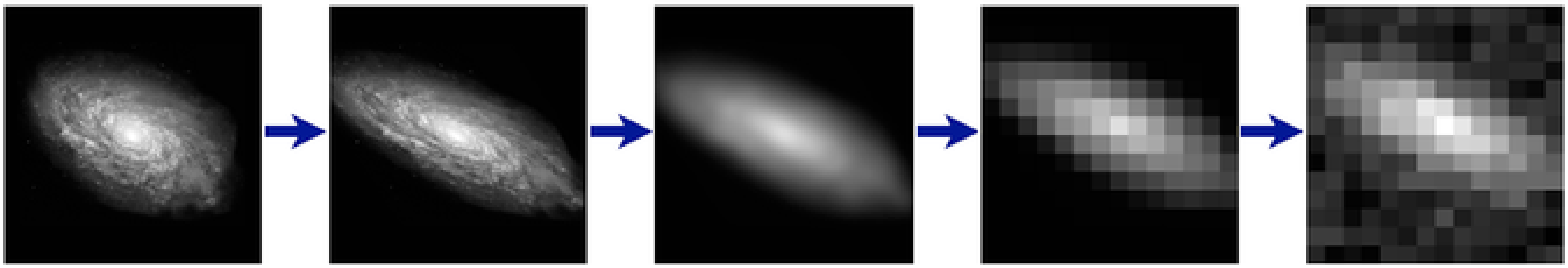


- Start with well-reduced data
- Aperture photometry:
  - Measure signal within an aperture centered on the source and sky/background in an annulus around the source
  - Simplest method, reliable with clean data
  - Subject to errors if have artifacts in the image (they come through unattenuated as signal)
  - Can be adapted to extended sources
- Point spread function (PSF) fitting
  - Fit an idealized image of an unresolved source to the image
  - Best for crowded fields
  - More immune to artifacts (fits tend to de-emphasize them)
  - Determining PSF may be difficult (it may be dependent on where a source is within the FOV, or on time of observation in a night)
  - More difficult to adapt to extended sources

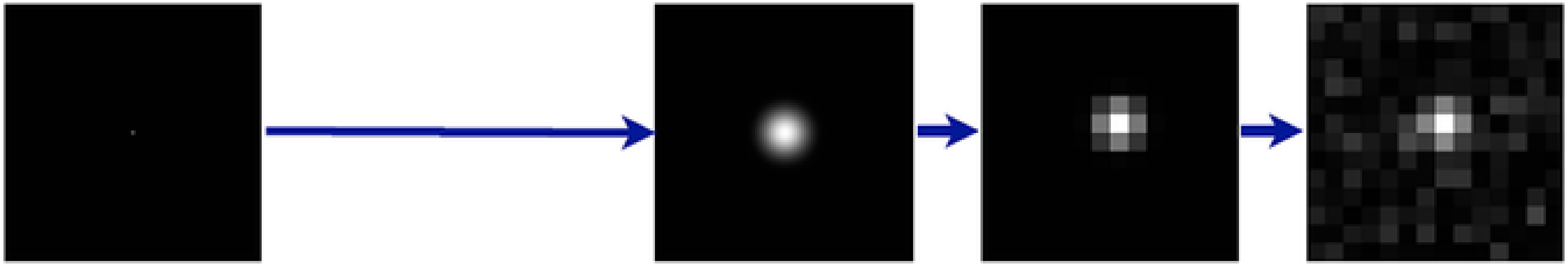


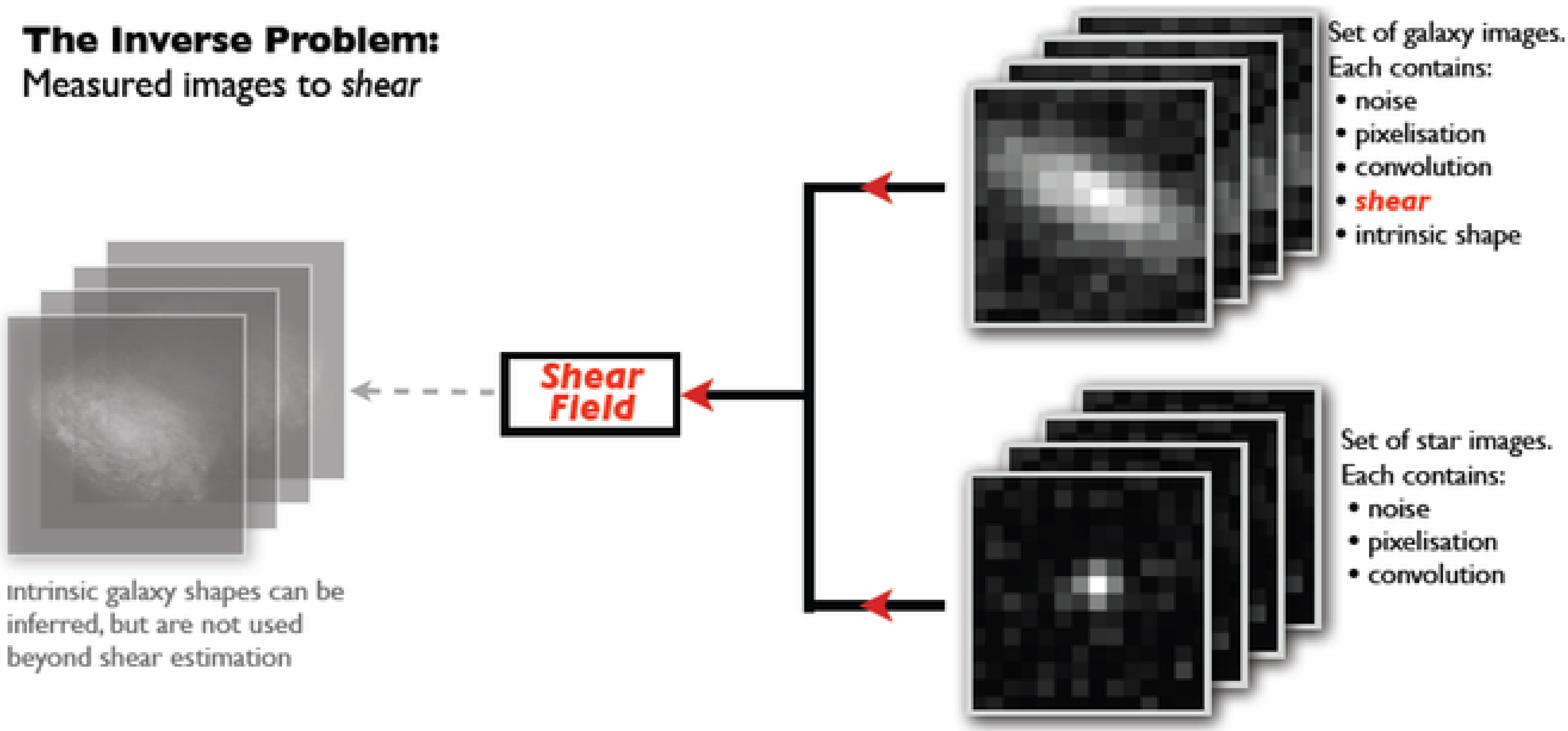
The Forward Process.

Galaxies: Intrinsic galaxy shapes to measured image:



Stars: Point sources to star images:




The Inverse Problem:
Measured images to *shear*
Set of galaxy images.
Each contains:
• noise
• pixelisation
• convolution
• *shear*
• intrinsic shape
Shear Field
Set of star images.
Each contains:
• noise
• pixelisation
• convolution
intrinsic galaxy shapes can be inferred, but are not used beyond shear estimation