How stars form now and in the early universe

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Preliminary Questions

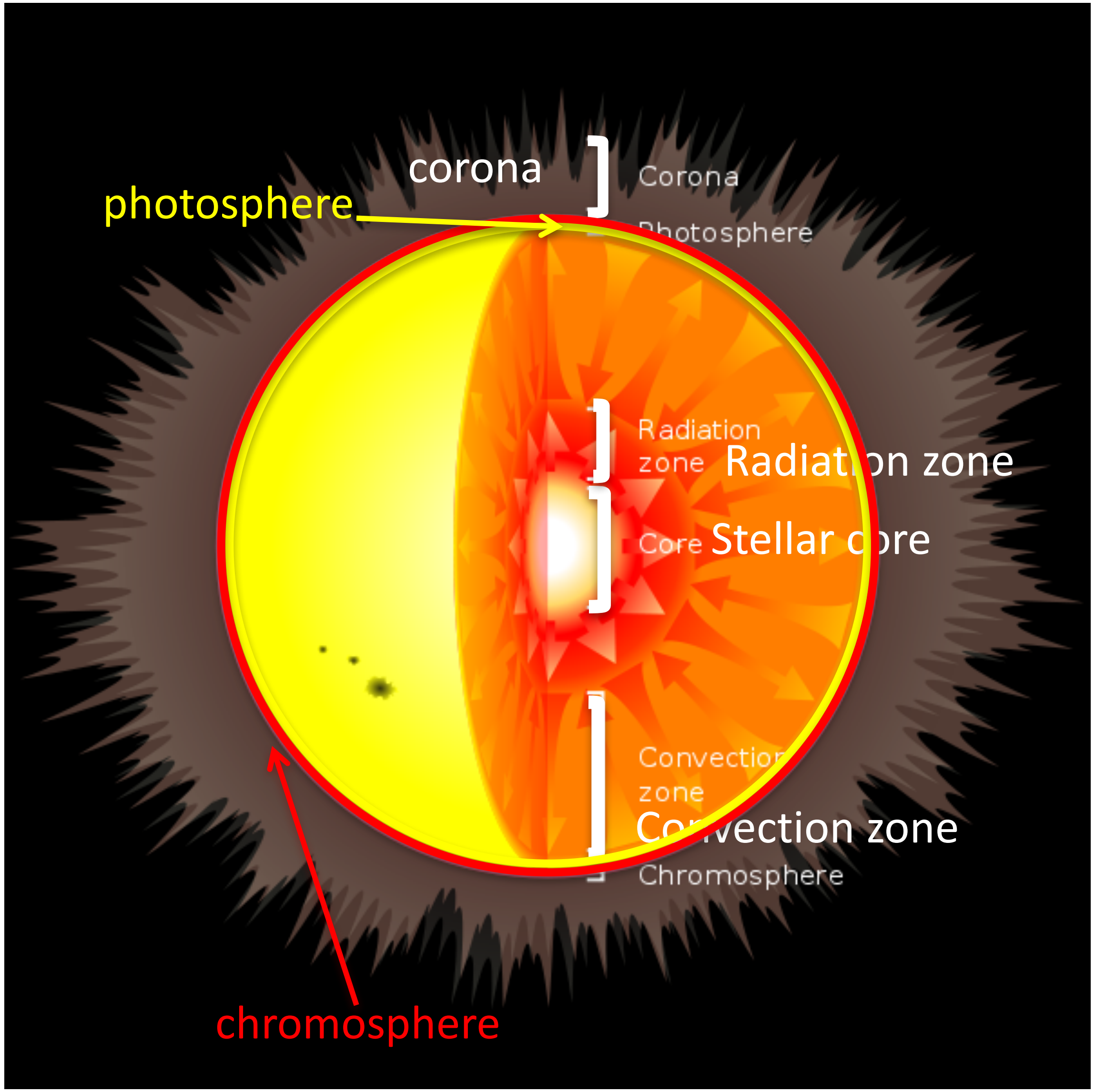
What kind of information is relevant to understand star formation for presently-forming stars and for the first population of stars appearing in the universe?

- The Initial Mass Function
- Cooling and coolants
- The Jeans instability
- Fragmentation of the clouds
- Detection of the first stars

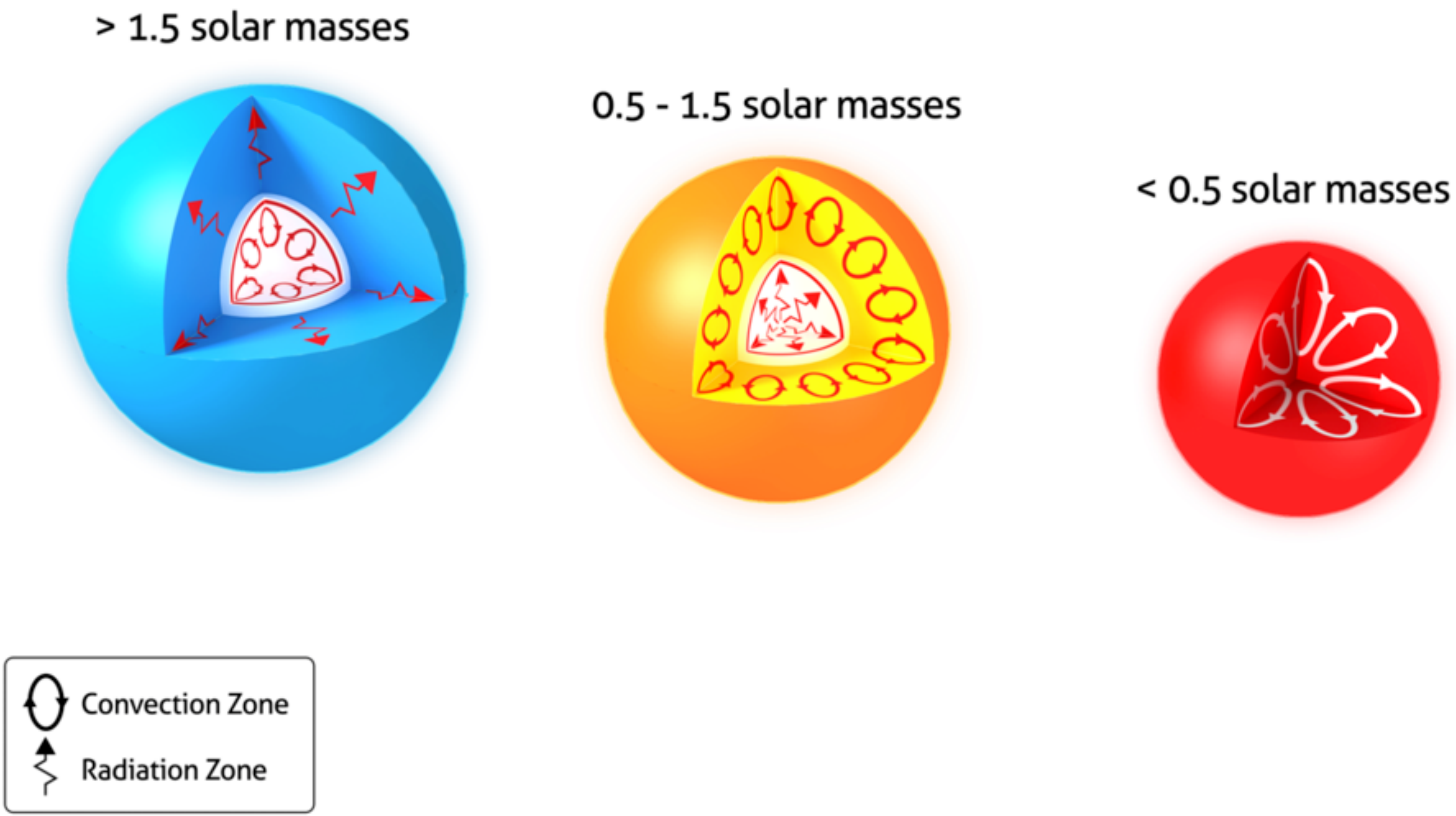
Introduction about stars

The building blocks in the universe: stars

- The **photosphere** is the layer is where the Sun's energy is released as light.
- The next layer is the **chromosphere**. The chromosphere emits a reddish glow as super-heated hydrogen burns off. But the red rim can only be seen during a total solar eclipse.
- The third layer of the Sun's atmosphere is the **corona.** It can only be seen during a total solar eclipse as well. It appears as white streamers or plumes of ionized gas that flow outward into space. Temperatures in the sun's corona can get as high as 2 million °C. As the gases cool, they become the solar wind.



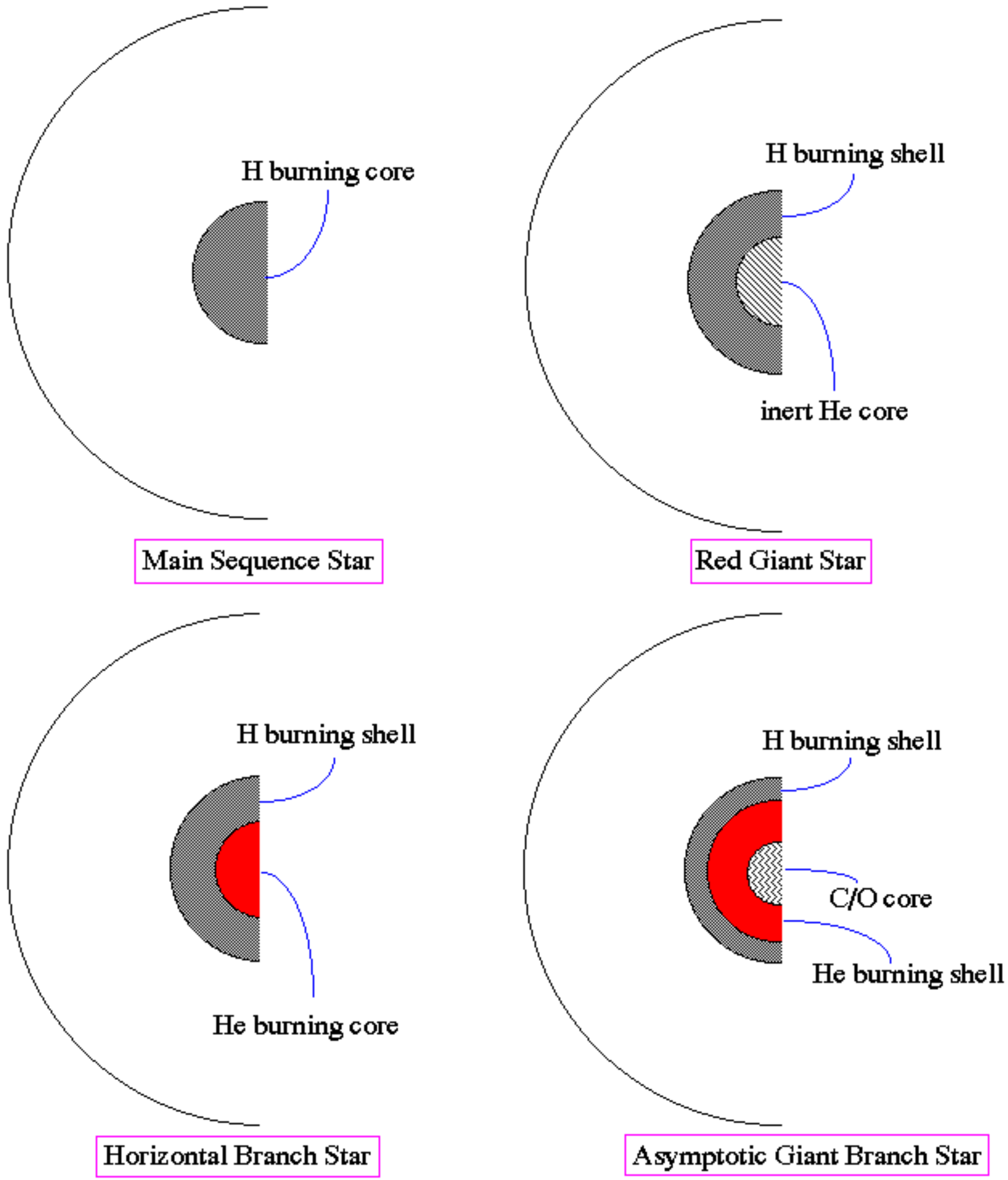
Heat Transfer of Stars



Summary of 1 $M_\odot$ evolution

Approximate typical timescales

Phase	*Time (yrs)*
Main-sequence	9×10^9
Subgiant	3×10^9
Redgiant Branch	1×10^9
Red clump	1×10^8
AGB evolution	$\sim 5 \times 10^6$
PNe	$\sim 1 \times 10^5$
WD cooling	$> 8 \times 10^9$



Star clusters

NGC3603 from Hubble Space Telescope

Star clusters are very useful to understand stars and their evolution:

• Stars are all at same distance

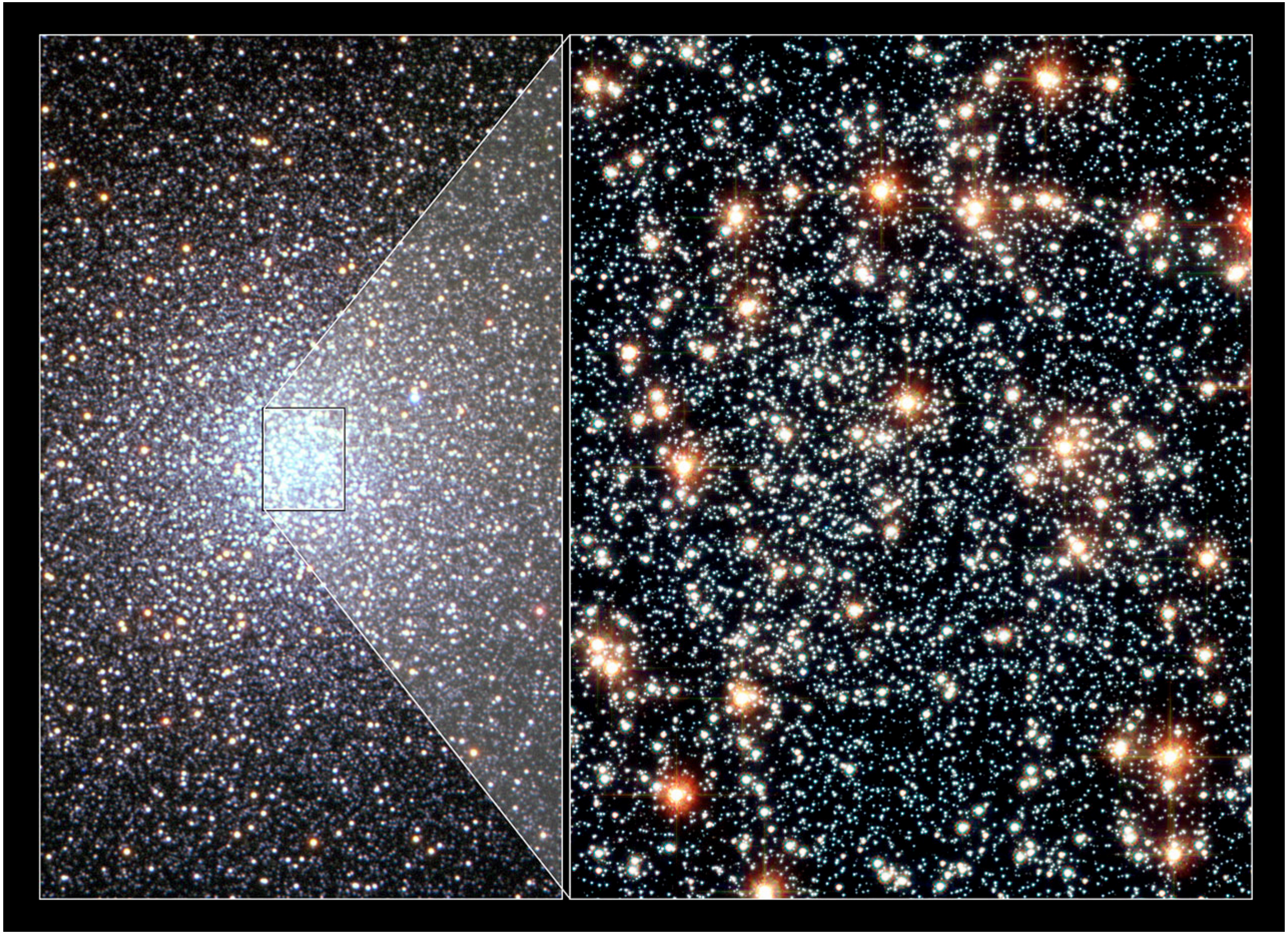
• Star clusters are dynamically bound

• Stars have same age

• Stars have same chemical composition

Can contain $10^3 - 10^6$ stars

Globular clusters are more massive star clusters



Let's open a short parenthesis

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- Galactic Globular Clusters (GCs) had been considered the best example of simple stellar population.
- However, in the last decades strong observational evidence has been gathered to prove that this is only a first approximation. In fact, GCs host multiple populations of stars differing in terms of chemical composition which also reflects in spreads and splits of the evolutionary sequences in CMDs.
- The multiple population phenomenon is very complex and not totally understood yet

Multiple stellar populations in the globular cluster M3 (NGC 5272): a Strömgren perspective

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(this is an example. You can find a few other similar papers on, e.g. Ω Centauri)

- **Blue symbols** represent First Generation (FG) stars
- **Red symbols** show Second Generation (SG) ones.
- The analysis of high-resolution spectra
- quantitatively demonstrates that the two detected sequences correspond to the first (Na-poor) generation and the second (Na-rich) generation.

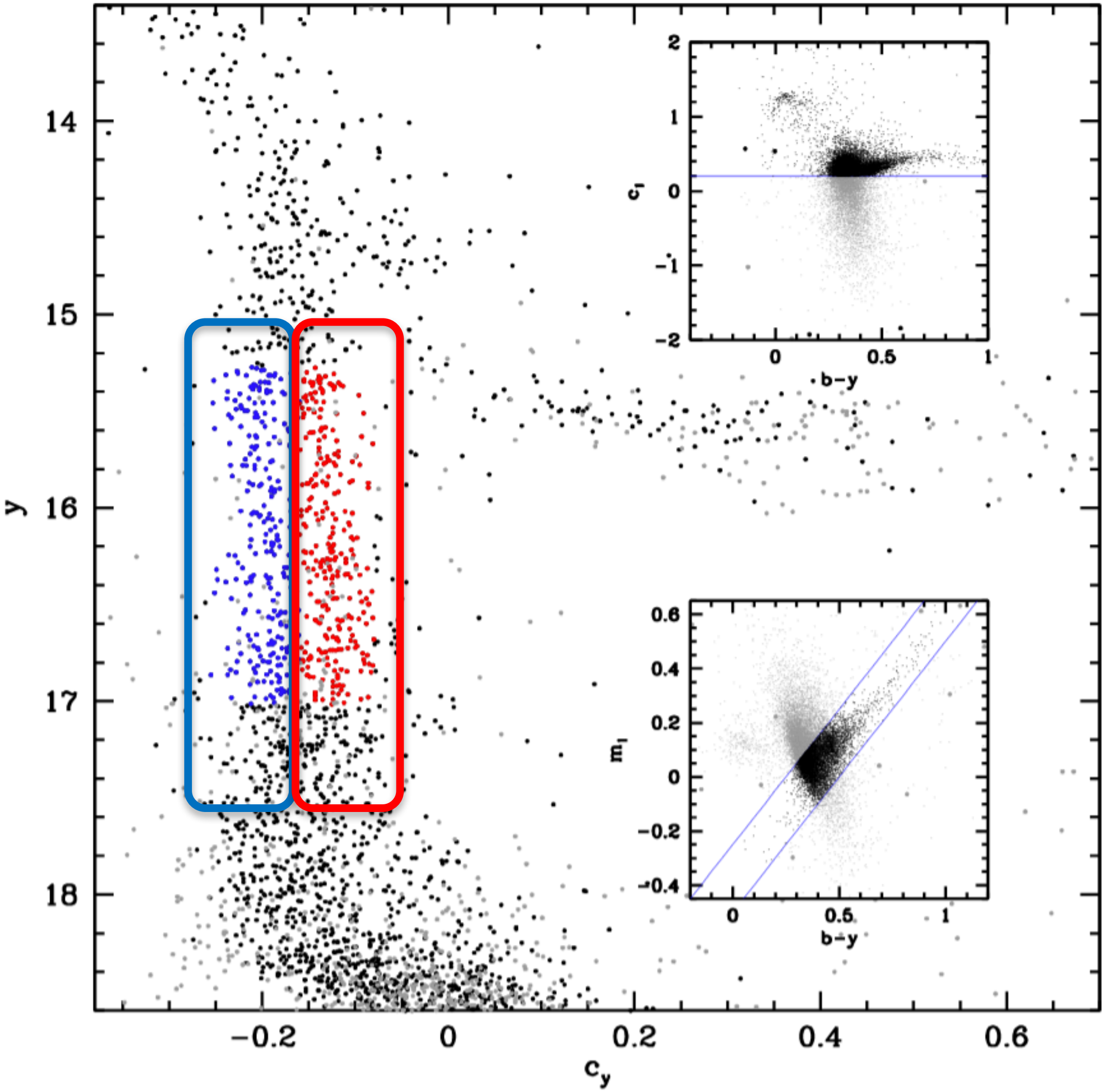

Figure 6. Decontaminated RGB of M3 (black dots) in the c_y, y CMD. The selection criteria are shown in the two inset, where we followed the prescriptions in (Frank et al. 2015) using the $b - y, c_1$ and $b - y, m_1$ planes.

The radial distributions of the two populations turned out to be quite peculiar:

- red and blue sequence being xmixed up to half the cluster half-mass radius (r_{hm})
- the red component staying more centrally concentrated out to about 2.5 r_{hm},
- At 2.5 r_{hm} the two mix again.

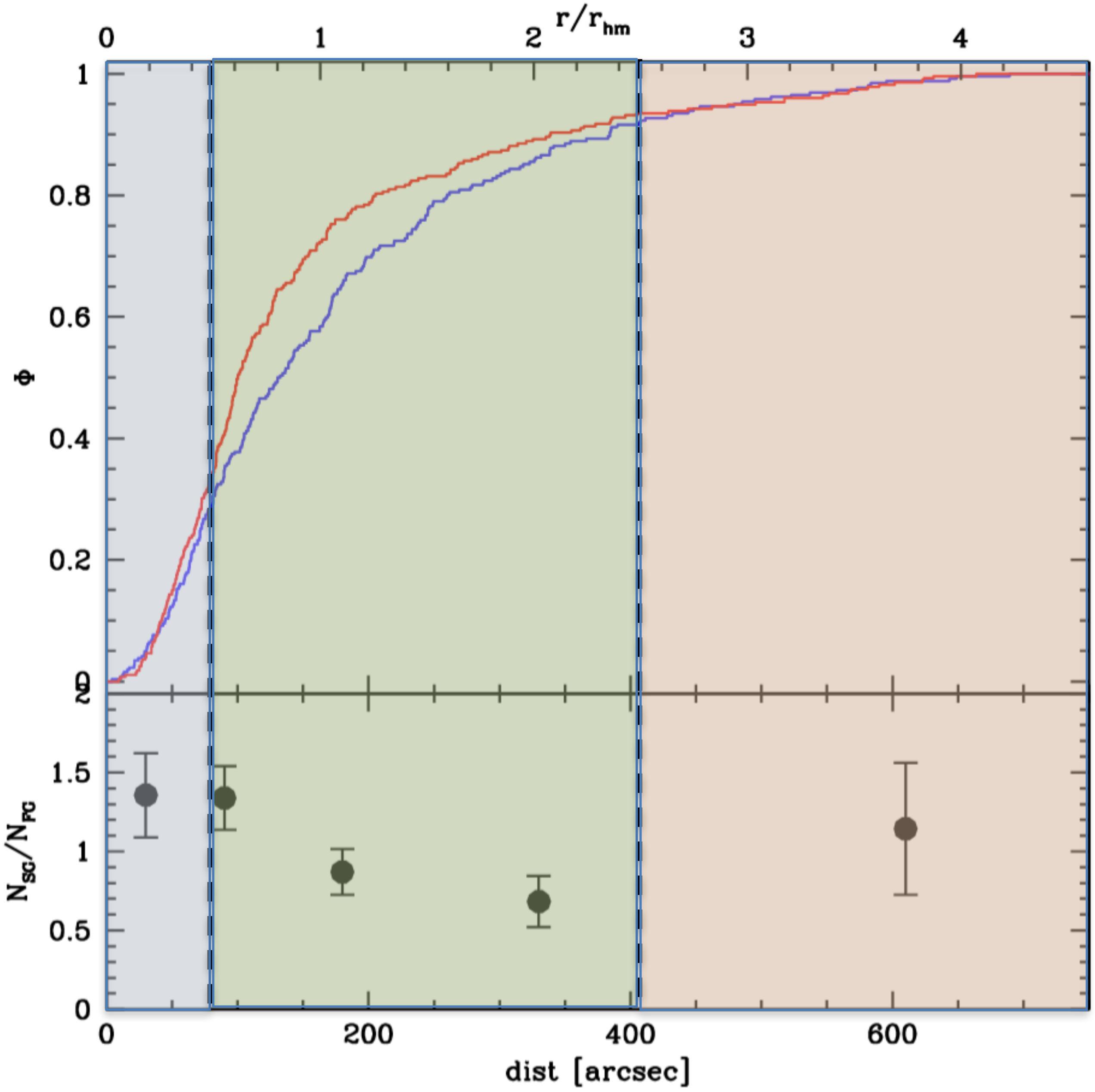

Figure 7. *Upper panel:* radial distributions for FG (blue) and SG (red lines) stars. Lower panel: SG-to-FG ratio as a function of the distance from the cluster centre. Radial bins are taken within 0.2, 0.6, 1, 2 and 3.5 r_{hm}.

With this Word of Caution in mind, let’s close the parenthesis

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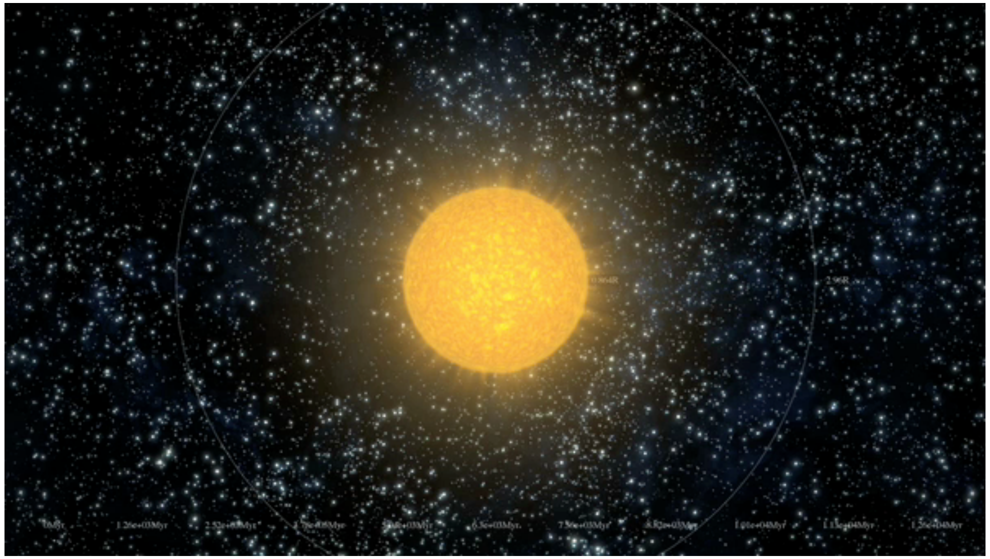
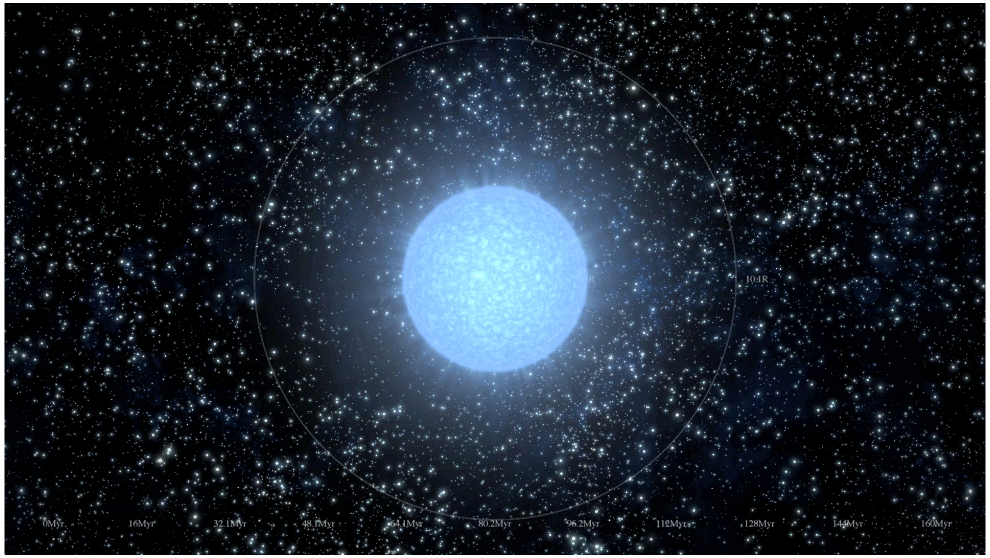
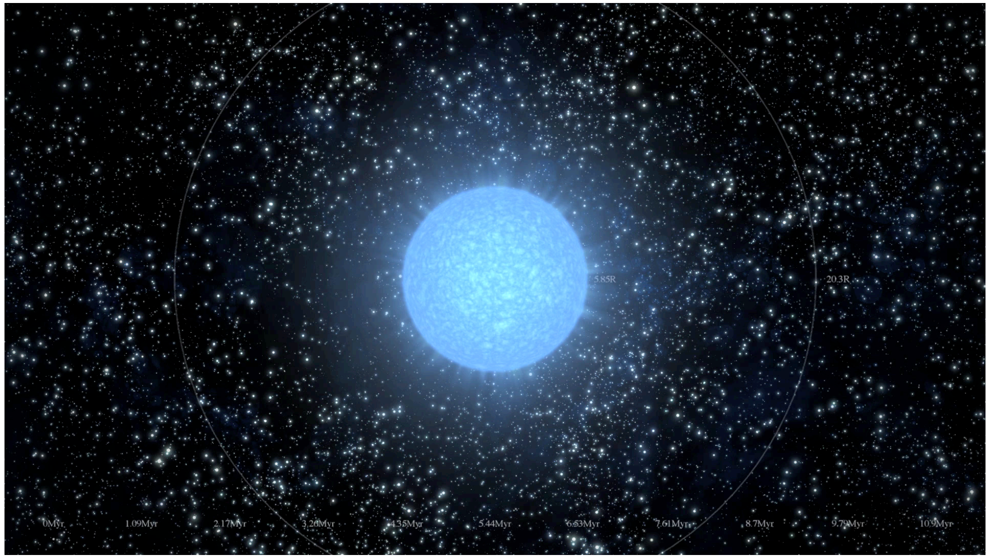
Comparative Evolutions of 1 $M_{\odot}$, 10 $M_{\odot}$ & 20 $M_{\odot}$ star

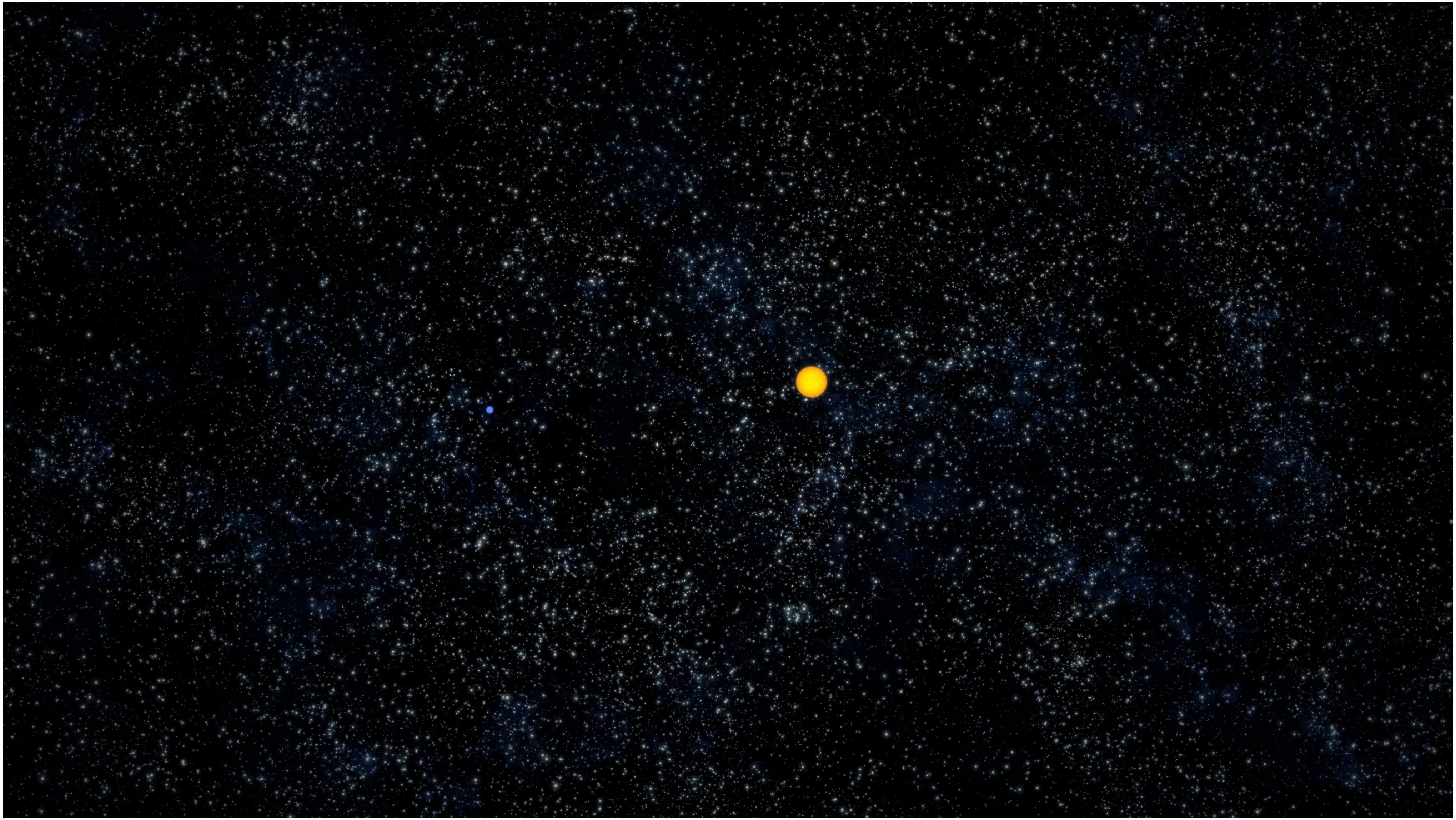






http://ciera.northwestern.edu/Research/stellar_evolution/stellar_evolution.php

Evolution of a binary stellar system (MS + NS)



http://ciera.northwestern.edu/Research/stellar_evolution/stellar_evolution.php

Observable properties of stars

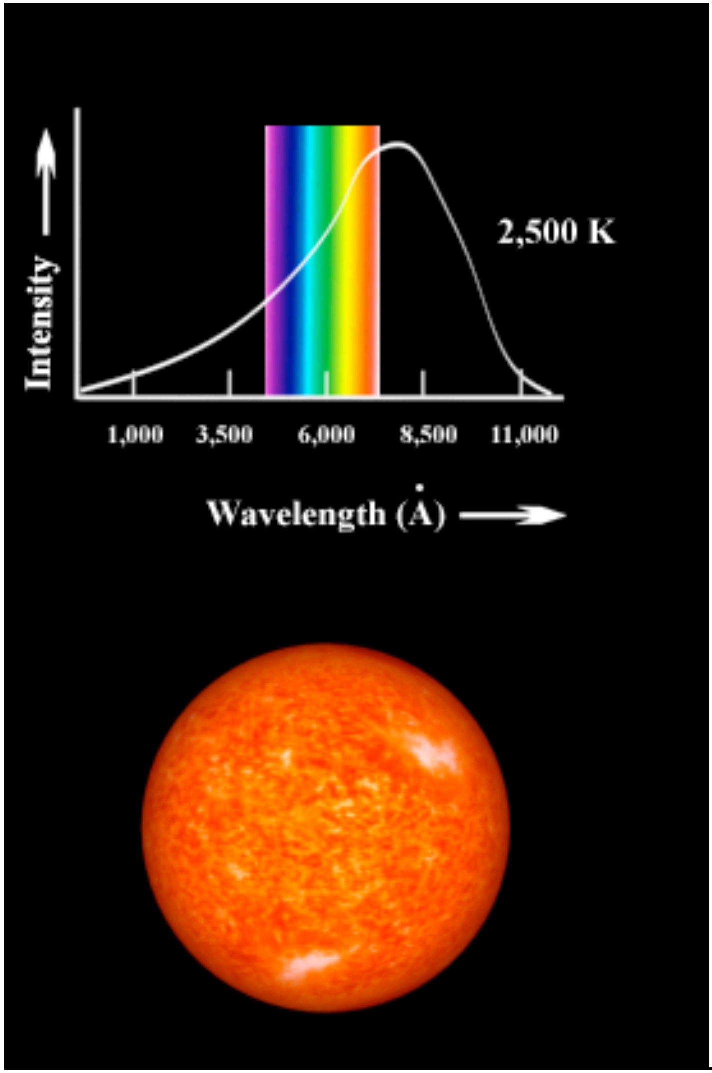
Basic parameters to compare theory and observations:

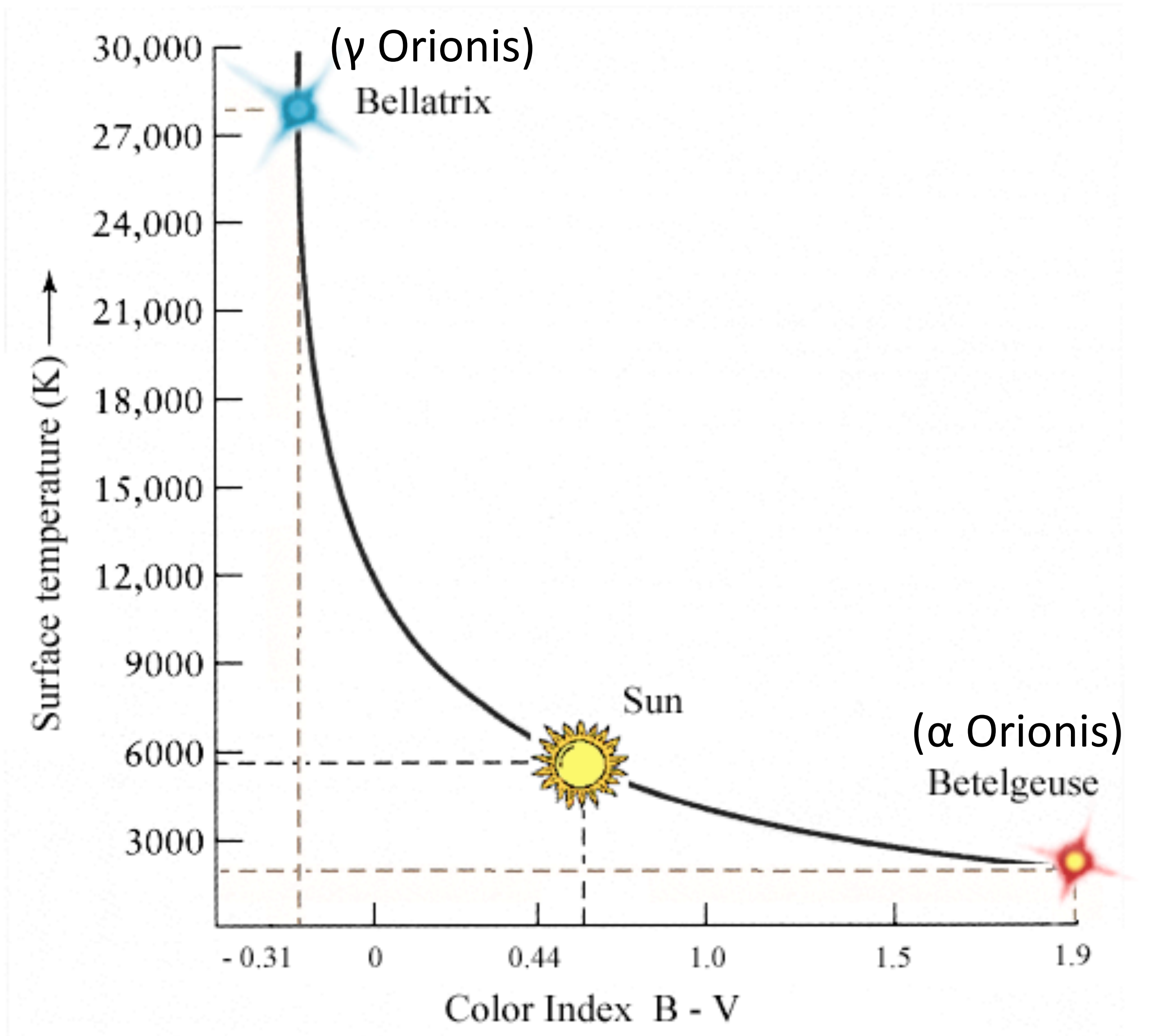
- Stellar mass (M)
- Luminosity (L)
  - The total energy radiated per second i.e. power (in W): $L = \int_0^\infty L_\lambda d\lambda$
  - If F is the flux ($W \cdot m^{-2}$) and L is the luminosity (W) or where F is the flux ($erg \cdot s^{-1} \cdot cm^{-2}$) and L is the luminosity ($erg \cdot s^{-1}$) and D the distance to the star: $\boxed{F = \frac{L}{4\pi D^2}}$
- Radius (R)
- Effective temperature (T_e)
  - The temperature of a black body of the same radius as the star that would radiate the same amount of energy. Thus

    $L = 4\pi R^2 \sigma T_e^4$

  where σ is the Stefan-Boltzmann constant ($5.67 \times 10^{-8}\ Wm^{-2}K^{-4}$)

Magnitudes and Colours





The Hertzsprung-Russell diagram

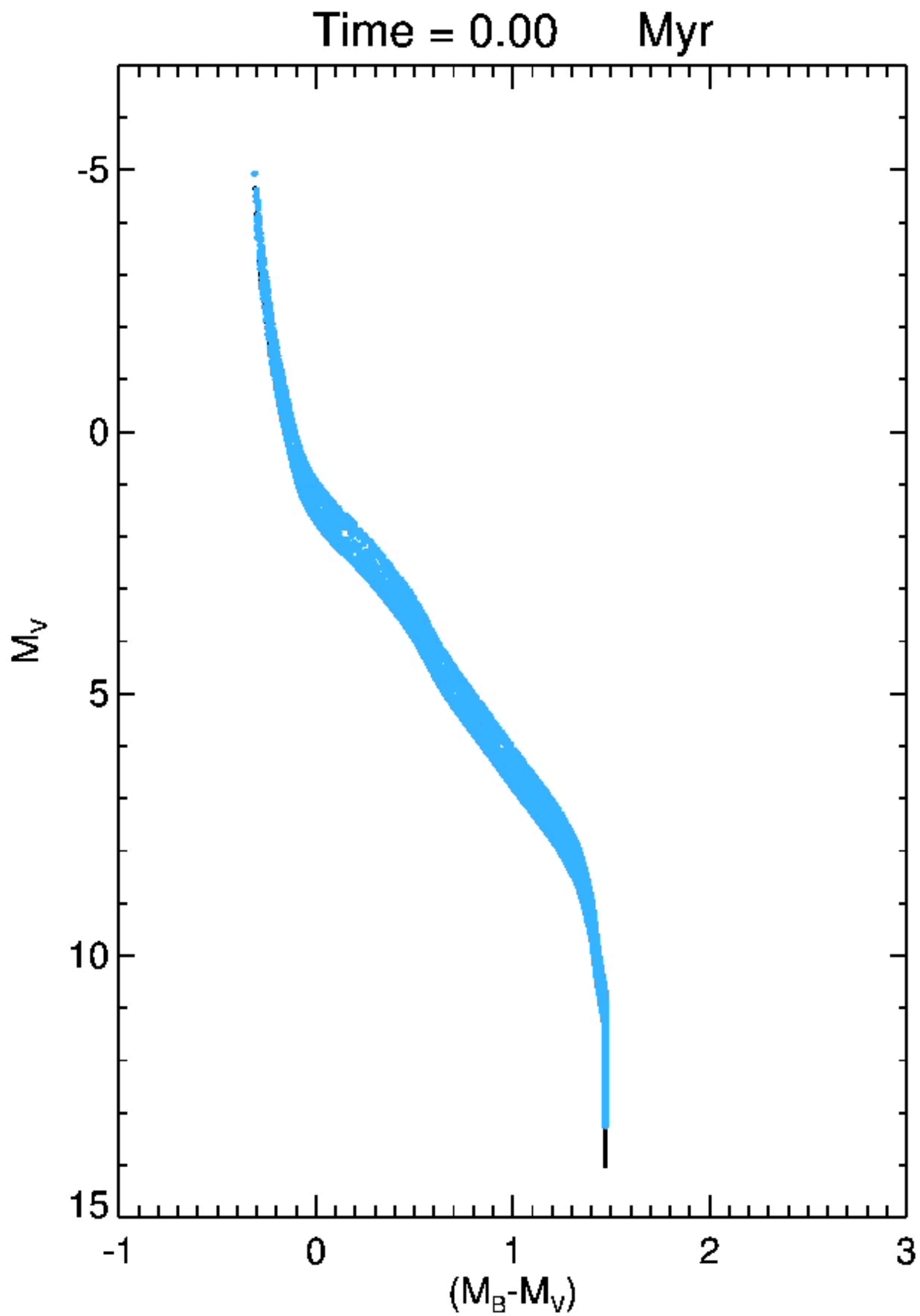
M, R, L and T_e do not vary independently.
Two major relationships:

1. L with T
2. L with M

1. The first is known as the *Hertzsprung-Russell* (HR) diagram or its observed counterpart, the *colour-magnitude* diagram (CMD).

The Hertzsprung-Russell diagram

Evolution of the Color-Magnitude Diagram of a Globular Cluster Assuming a Single Stellar Population



http://ciera.northwestern.edu/Research/stellar_evolution/stellar_evolution.php

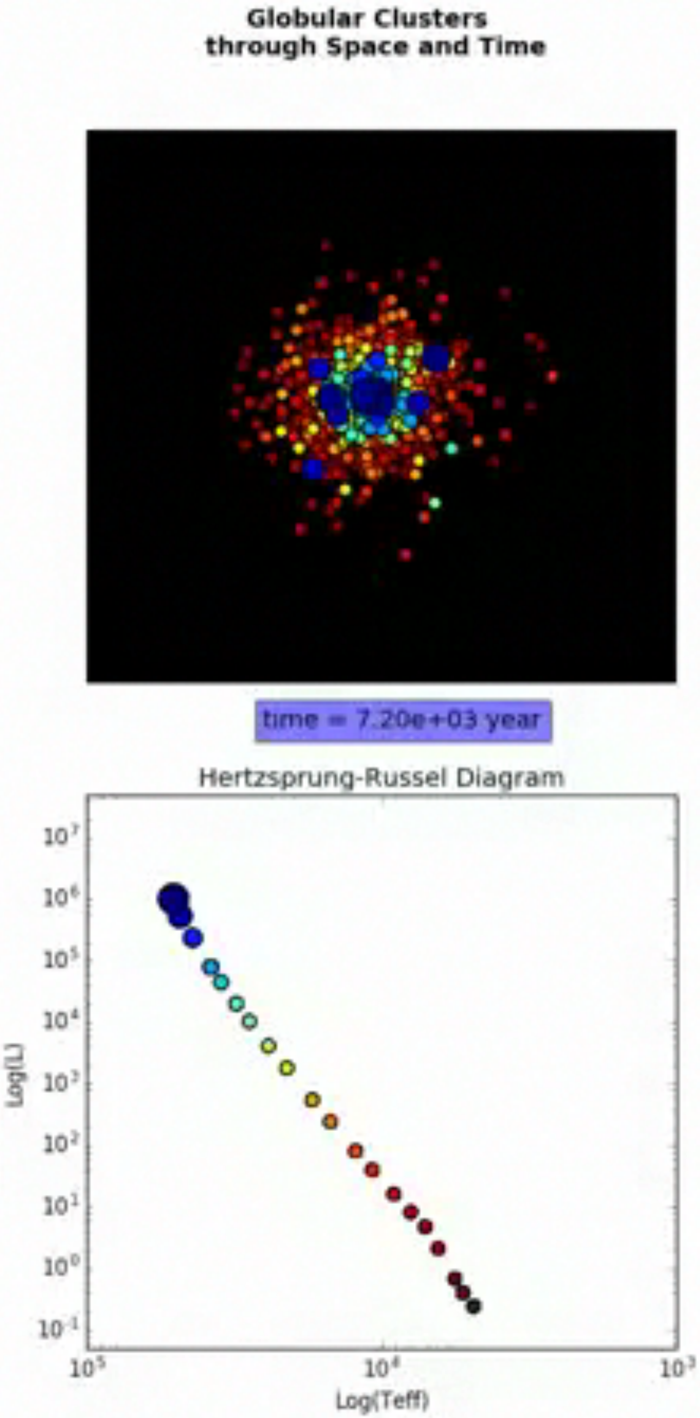
Project carried within the Master SPaCE 2015-2016:

« Globular Clusters through Space & Time »

by

Jon FERNANDEZ OTEGI
Théo LOPEZ
Carla RODRIGUEZ

https://people.lam.fr/burgarella.denis/denis/Master_SPaCE.html




```
# We reset the main parameters of the plot
def update_plot(i, temperature, luminosity,
color, size, elev, azim, dist, scat1, scat2):
    global pause
    # Set colors...
    scat1.set_array(color[i])
    # Set sizes...
    scat1.set_sizes(size[i])
    # Set elevation and azimuth...
    ax1.view_init(elev=elev[i], azim=azim[i])
    # Set distance...
    ax1.dist=dist[i]
    position =
np.column_stack((temperature[i,:],luminosity[i,
:]))
    #print('position', position[i], color[i], size[i])
    scat2.set_offsets(position)
    scat2.set_array(color[i])
    scat2.set_sizes(size[i])

    point.set_data(x, y)
    time_text.set_text(time_template %
(big_array[0,i,0]))

    return scat1, scat2, point, time_text

ani = animation.FuncAnimation(fig,
update_plot, frames=range(numframes),
                              fargs=(temperature_data,
luminosity_data,
                              color_data, size_data,
                              elev_data, azim_data,
dist_data,
                              scat1, scat2), blit=False,
interval=25, repeat=False,
                              init_func=init)

plt.show()
ani.save('Glob_Evol.mp4', fps=24,
extra_args=['-vcodec', 'libx264'])
```

M, *R*, *L* and T_e do not vary independently.
Two major relationships:

1. *L* with *T*
2. *L* with *M*

2. The second is known as the *Mass– Luminosity relation.*

2. The Mass-luminosity relation

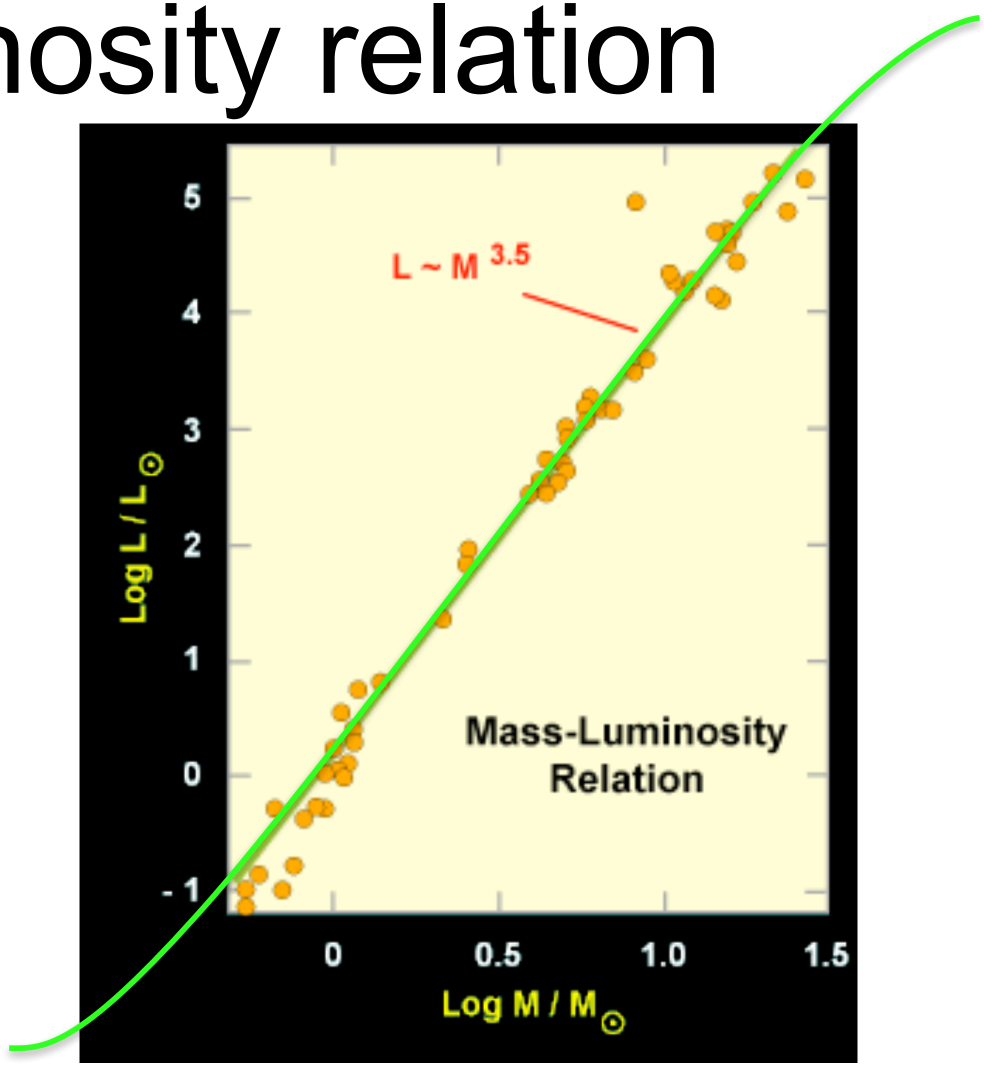
For main-sequence stars, there is a *Mass-luminosity relation.*

$$L \propto M^n$$

Where n ~ 3.5.

Slope changes at extremes, less steep for low and high mass stars.

This implies that the main-sequence (MS) on the HRD is a function of mass i.e. from bottom to top of main-sequence, stars increase in mass



We must understand the *M-L* relation and *L-T_e* relation theoretically. Models must reproduce observations

The Mass-luminosity relation

We assume stars are black body of radius R. The luminosity of a star(σ is Stefan's constant) is obtained as:

$$L = 4\pi R^2 \sigma T_E^4,$$

We also assume the star is in hydrostatic equilibrium:

$$\frac{dP}{dr} = -\frac{Gm(r)\rho(r)}{r^2}.$$

Which gives, by integrating both sides from the core (*r = 0*) to the surface (*r = R*) of the star:

$$\langle P \rangle = -\frac{1}{3}\frac{E_{GR}}{V} \quad \text{(Eq. 1)}$$

If we further assume that the potential energy of a spherical mass distribution is:

$$E_{GR} = -\frac{3}{5}\frac{GM^2}{R}.$$

And substitute V in Eq. 1 , we get:

$$\langle P \rangle \approx \frac{GM^2}{4\pi R^4}.$$

The Mass-luminosity relation

Now, if we use the ideal gas law (*PV* = *nkT*) to get the temperature.

$$\langle P \rangle = \frac{\langle \rho \rangle}{\bar{m}} kT,$$
$$kT = \frac{GM\bar{m}}{3R}.$$

Where $\bar{m}$ is the mean mass of gas particles within the star.

Substituting this into the initial luminosity equation ($L = 4\pi R^2 \sigma T_E^4$)

along with

$$R = \left(\frac{3}{4}\frac{1}{\rho\pi}M\right)^{\frac{1}{3}}$$

Finally yields: .... $L \propto M^{3.33}$.

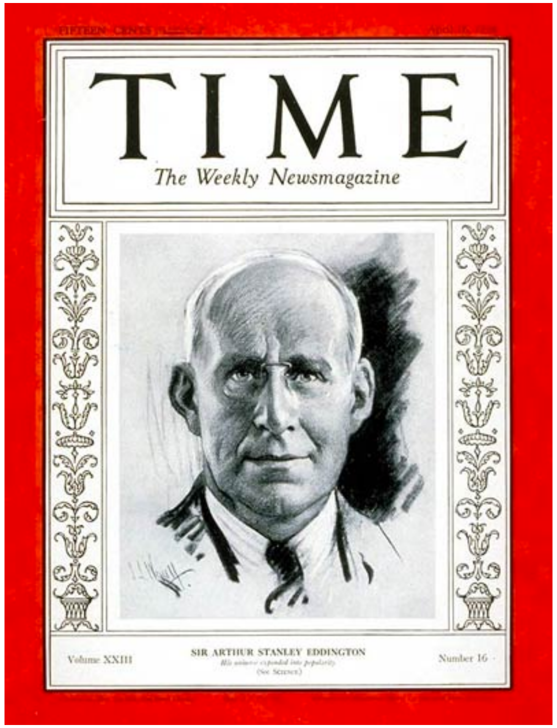
The Mass-luminosity relation

We did not derive a theoretically exact Mass / Luminosity relation here. It requires building a thermodynamical model of the internal structure of a star.

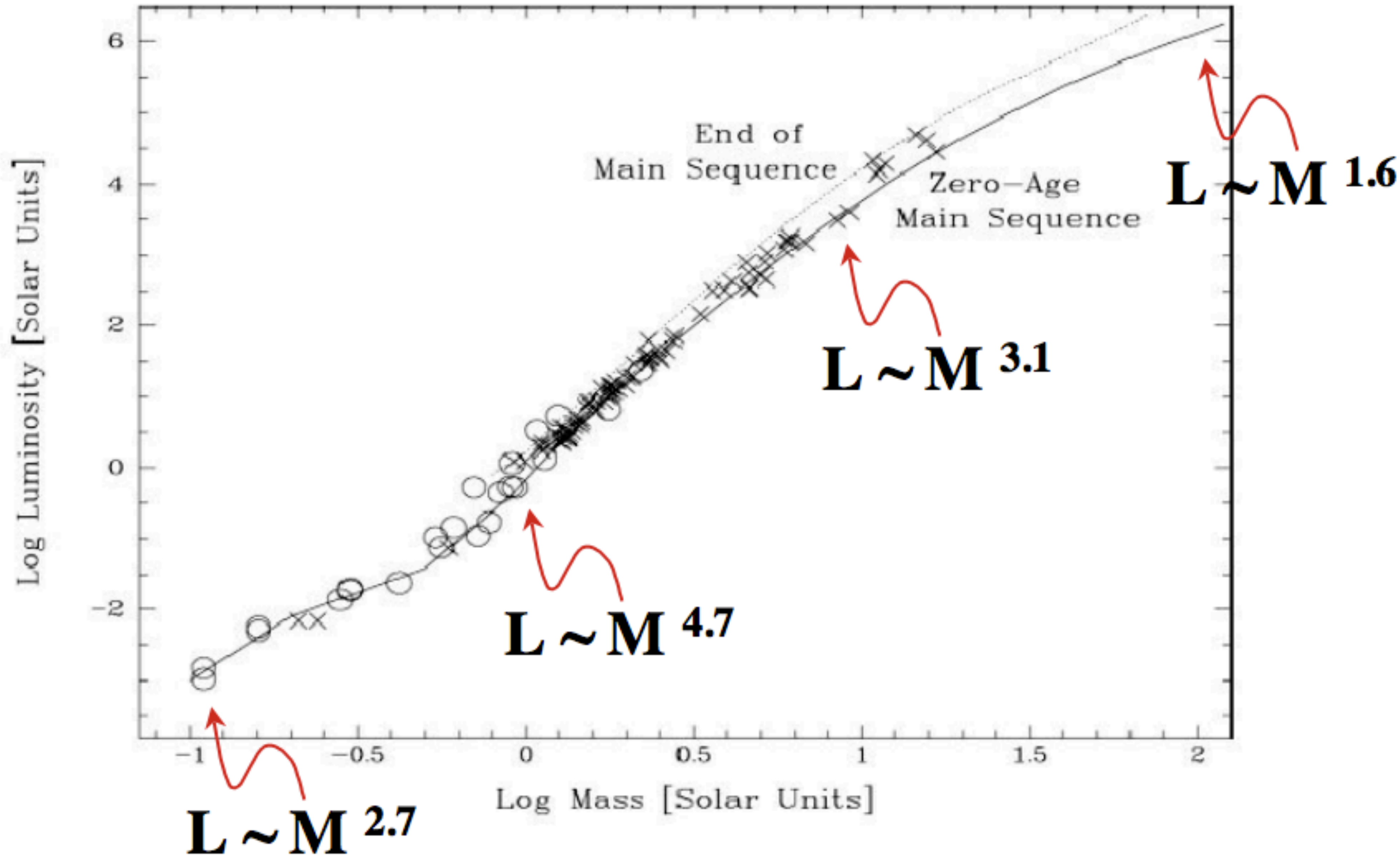
However, using some basics physics, we have reached an honorable representation of the observed relation $L \propto M^3$.

This is about what Arthur Eddington did in 1924.

This shows that stars can be modelled as ideal gases.
It was a new and radical idea at the time.



Over a wide mass range



Mass functions

- The stellar masses are one of the most important factors in determining their evolution, so when studying a stellar population, we are interested in estimating their masses.
- An important information is the number of stars per unit mass which is called a mass function.
- We define the mass function *Φ(M)* such that *Φ(M) dM* is the number of stars with masses between *M* and *M + dM*.
- With this definition, the total number of stars with masses between M_{low} and M_{up} is:

$$N(M_{low}, M_{up}) = \int_{Mlow}^{Mup} \Phi(M) dM$$

- By deriving both sides, we get:

$$\frac{dN}{dM} = \Phi(M)$$

- *Φ* is the derivative of the number of stars with respect to mass, i.e. the number of stars *dN* within some mass interval *dM*.

Total mass of stars between M_{low} and M_{up}

- If we are interested in the total mass of stars between M_{low} and M_{up} in a given system rather than the number of such stars, we must integrate Φ times the mass per star. Thus the total mass of stars with masses between M_{low} and M_{up} is:

$$M_*(M_{low}, M_{up}) = \int_{Mlow}^{Mup} M\Phi(M)dM$$

- or equivalently:

$$\frac{dM_*}{dM} = M\Phi(M) = \xi(M)$$

$y = ln(m)$
$dy/dm = 1/m$
$dy = dm/m$
$d[ln(m)] = dm/m$

- $\xi(M)$ gives the number of star per *ln(M)*, rather than per number in mass.
- **Physical explanation**:
  - Suppose that $\Phi(M)$ = Cte constant: there are as many stars from 1 − 2 $M_\odot$ as there are from 2 − 3 $M_\odot$ as there are from 3 − 4 $M_\odot$, etc.
  - Instead suppose that $\xi(M)$ = Cte: there are equal numbers of stars in intervals that cover an equal range in logarithm, so there would be the same number from 0.1 − 1 $M_\odot$, from 1 − 10 $M_\odot$, from 10 − 100 $M_\odot$, etc.

- The distribution of initial stellar masses (IMF) might be invariant **[*hot topic*]**.

- If we examine two populations, of galaxies of different sizes, then $\Phi(M) = dN/dM$ will be different because they have different numbers of stars. However, they may have the same fraction of their stars in a given mass range.

- So, it is generally common to normalize Φ or ξ so that the integral is unity, i.e. to compute a normalization factor for Φ or ξ such that the integrals are equal to 1.

- When the mass is normalized, $\Phi(M)\, dM$ and $\xi(M)\, dM$ give the fraction of stars (fraction by number for Φ and fraction by mass for ξ) with masses between M and $M + dM$.

$$\int_{m_{\rm low}}^{m_{\rm up}} m\phi(m)dm = \int_{m_{\rm low}}^{m_{\rm up}} \xi(m)dm = 1$$ *with* $m_{low} = 0.1M_\odot$ *and* $m_{up} = 120M_\odot$

In LMC at 165 000 l.y.
•R136a1 : 265 $M_\odot$ (?)
•R136a2 : 195 $M_\odot$ (?)
ηCarinae : 120 $M_\odot$

The Salpeter Initial Mass Function (IMF)

If we ignore the low-mass flattening of the IMF below ~ $1M_\odot$, we might assume that the same slope holds over the range $M_{min} = 0.1M_\odot$ *to* $M_{max} = 120M_\odot$.

This is known as the Salpeter IMF: $\phi(M) = \phi_0\ M^{-2.35}$
The normalization ϕ_0 (in units of $M_\odot$) is evaluated by requiring that the integral equals to 1:

$$1 = \int_{M\,\min}^{M\,\max} \Phi(M)\,dM = \int_{M\,\min}^{M\,\max} AM^{-2.35}\,dM = \frac{A}{-1.35}(M_{\max}^{-1.35} - M_{\min}^{-1.35}) = \frac{A}{1.35}(M_{\min}^{-1.35} - M_{\max}^{-1.35})$$

$$\boxed{A = \frac{1.35}{M_{\min}^{-1.35} - M_{\max}^{-1.35}} = 0.060\ with\ M_{\min} = 0.1M_\Theta\ and\ 120M_\Theta}$$

In a similar way, we have in terms of mass:

$$1 = \int_{M\,\min}^{M\,\max} \xi(M)\,dM = \int_{M\,\min}^{M\,\max} BM^{-1.35}\,dM = \frac{B}{-0.35}(M_{\max}^{-0.35} - M_{\min}^{-0.35}) = \frac{B}{0.35}(M_{\min}^{-0.35} - M_{\max}^{-0.35})$$

$$\boxed{B = \frac{0.35}{M_{\min}^{-0.35} - M_{\max}^{-0.35}} = 0.17\ with\ M_{\min} = 0.1M_\Theta\ and\ 120M_\Theta}$$

Illustration

So, we will have: $\phi(M) = 0.060\ M^{-2.35}$ and $\xi(M) = 0.17\ M^{-1.35}$ for $M = 0.1 - 120M_{\odot}$

We can estimate the fraction of stars by number and by mass in a given mass range. For instance, what is the fraction of stars by number (1) and by mass (2) more massive than the Sun in a new-born population with a Salpeter IMF?

1) $$f_N(>M_\Theta) = \int_1^{120} 0.060\Phi(M)dM = \int_1^{120} 0.060M^{-2.35}dM =$$
$$f_N(>M_\Theta) = \frac{0.060}{1.35}(1^{-1.35} - 120^{-1.35}) \boxed{= 0.045}$$

2) $$f_M(>M_\Theta) = \int_1^{120} 0.17M\Phi(M)dM = \int_1^{120} 0.17M^{-1.35}dM =$$
$$f_M(>M_\Theta) = \frac{0.17}{0.35}(1^{-0.35} - 120^{-0.35}) = \boxed{0.40}$$

4.5% of the stars are more massive that the Sun but,
the mass of these stars amounts to 40% of the new-born stars.

Two remarks

The total stellar mass of a system is computed from the following integral:

$$M_{tot} = \int_{M\,\min}^{M\,\max} M\Phi(M)dM = \int_{M\,\min}^{M\,\max} AMM^{-2.35}dM = \int_{M\,\min}^{M\,\max} AM^{-1.35}dM$$

$$M_{tot} = \frac{A}{-0.35}(M_{\max}^{-0.35} - M_{\min}^{-0.35}) = \frac{A}{0.35}\boxed{(M_{\min}^{-0.35} - M_{\max}^{-0.35})}$$

It shows that most of the stellar mass lies in low-mass stars.

The total luminosity (assuming L = C $M^{3.5}$) is computed as follows:

$$L_{tot} = \int_{M\,\min}^{M\,\max} L(M)\Phi(M)dM = \int_{M\,\min}^{M\,\max} CM^{3.5}AM^{-2.35}dM = \int_{M\,\min}^{M\,\max} ACM^{1.15}dM$$

$$L_{tot} = \frac{AC}{2.15}\boxed{(M_{\min}^{2.15} - M_{\max}^{2.15})}$$

which shows that the total stellar luminosity is dominated by the most massive stars.

Other more realistic IMFs (mainly for faint stars)

IMF Scalo (1998): $\xi(m) = m^{-\alpha}$

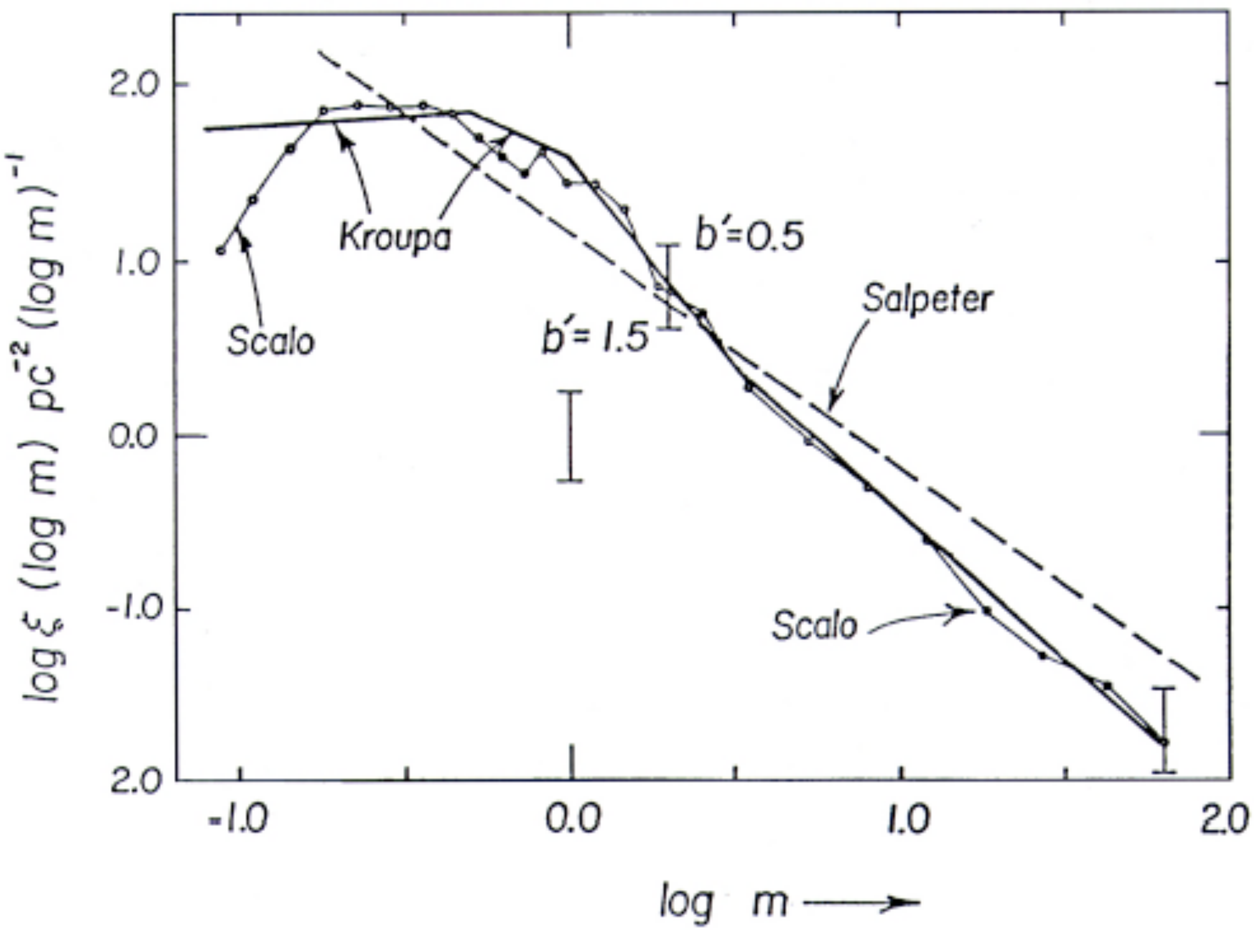
- $\alpha = -0.2 \pm 0.3$ for $0.08 \leq m/M_{\odot} < 1\ M_{\odot}$
- $\alpha = -1.7 \pm 0.5$ for $1 \leq m/M_{\odot} < 10\ M_{\odot}$
- $\alpha = -1.3 \pm 0.5$ for $10 \leq m/M_{\odot} < 100\ M_{\odot}$

IMF Kroupa 2001: $\xi(m) = m^{-\alpha}$

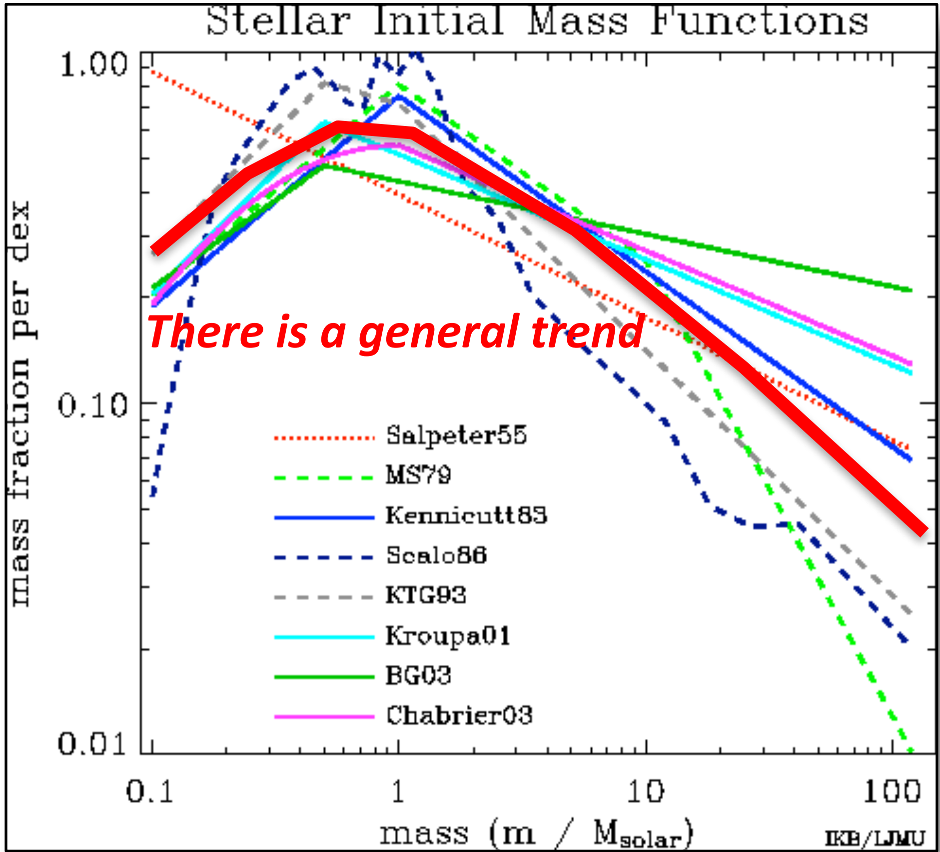
[$\alpha = 0.3$ for $m/M_{\odot} < 0.08$]

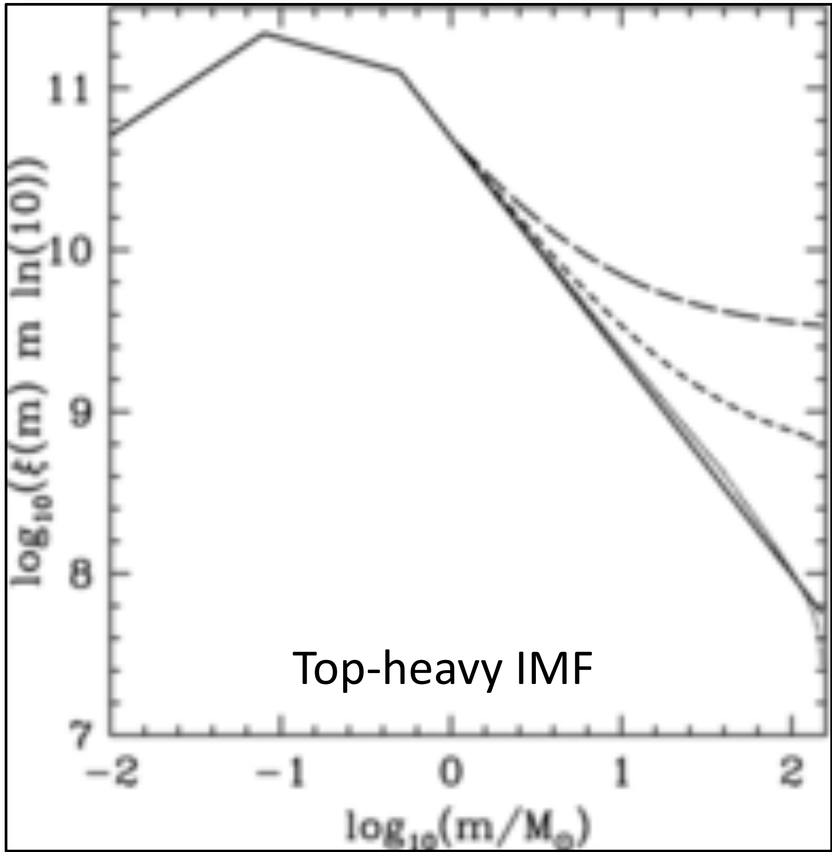
$\alpha = 1.3$ for $0.08 \leq m/M_{\odot} < 0.5$

$\alpha = 2.3$ for $0.5 \leq m/M_{\odot}$

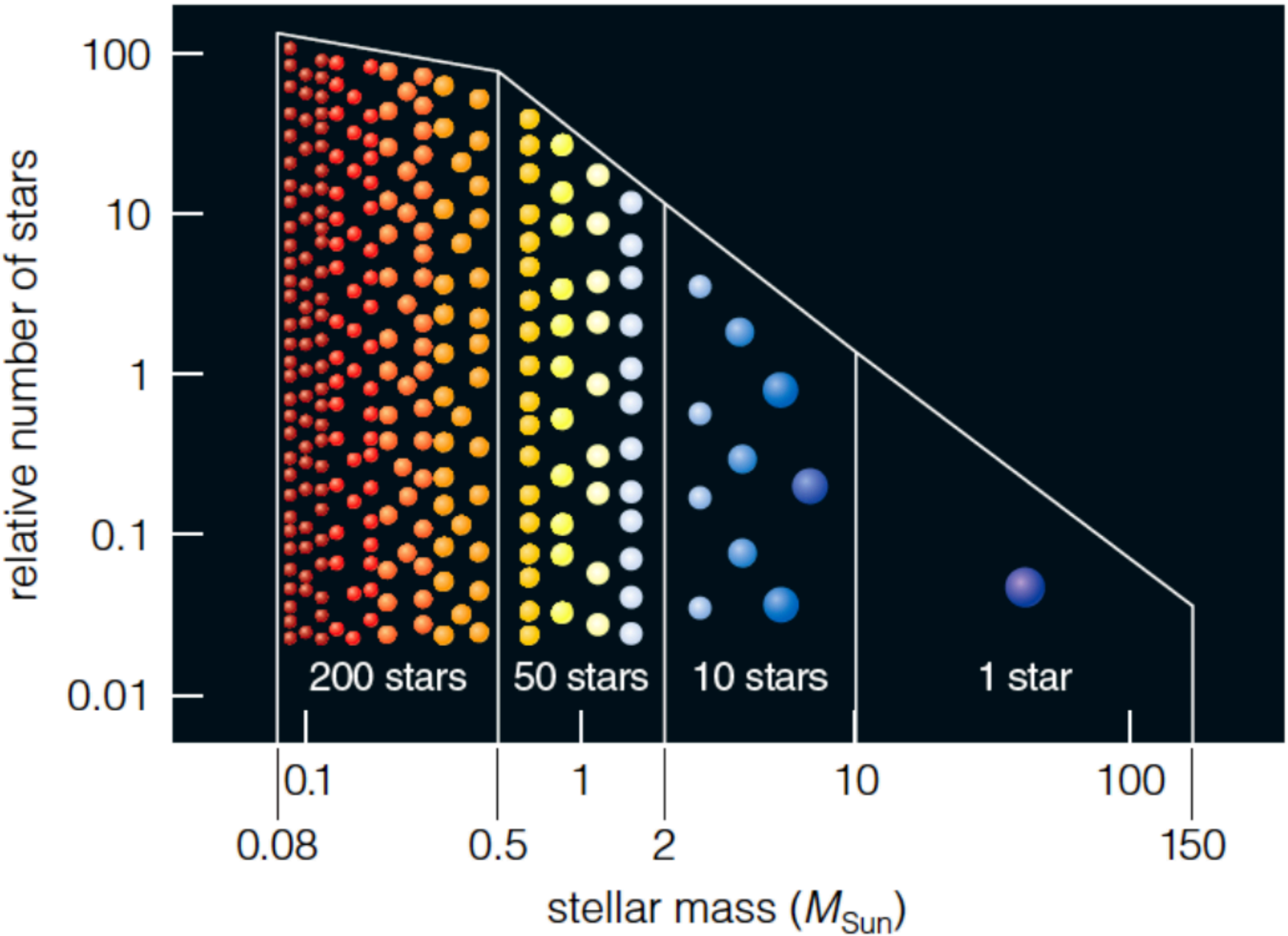


Sample Initial Mass Functions of Stars





$$\Phi_{top-heavy}(M) \propto \begin{cases} \Phi_{Salp}(M) \propto M^{-2.35},\ for\ 0.1 M_{\Theta} < M < 100 M_{\Theta} \\ e.g.\ M^{-1},\ for\ 100 M_{\Theta} < M < 500 M_{\Theta} \end{cases}$$


relative number of stars
100
10
1
0.1
0.01
200 stars
50 stars
10 stars
1 star
0.1
1
10
100
0.08
0.5
2
150
stellar mass (M_{Sun})

Variability of the IMF ?

An important *if not crucial question* is:

- Is the Initial Mass Function universal or does it vary with the environment, the element abundances (metallicity), the redshift, the star formation rate, the Jeans mass, … ?

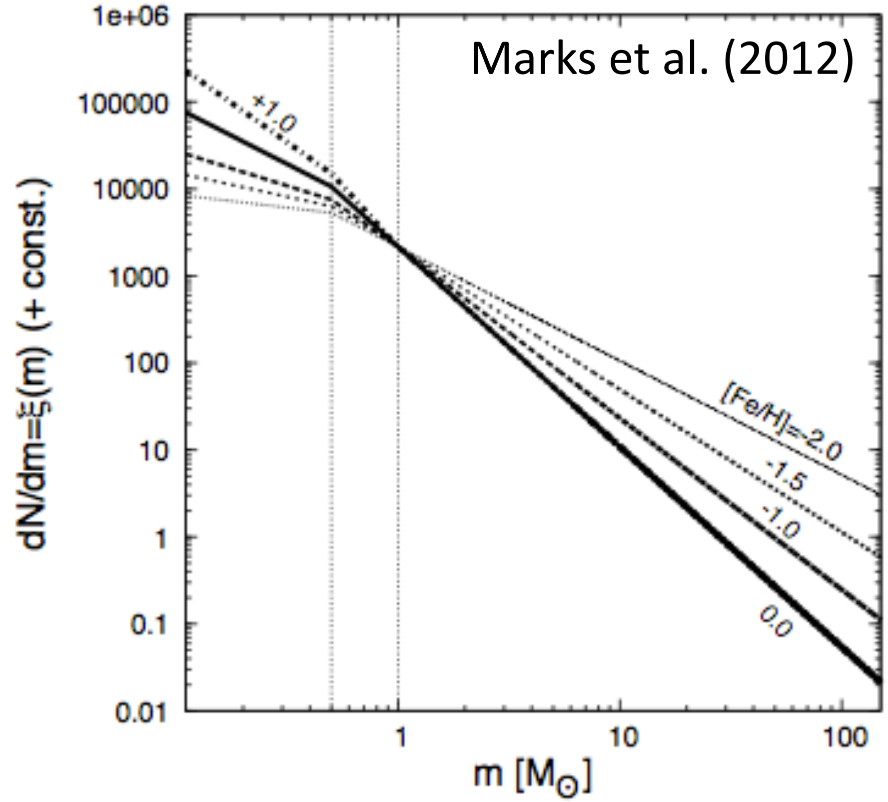


Figure 5. Suggested shape of the stellar IMF for different metallicities, [Fe/H] (not taking into account the density dependence of the IMF). The IMFs are scaled such that their values agree at $m = 1M_\odot$. Above $1M_\odot$ the IMF slope is determined by the present work (Fig. 4, equation 11). Below $1M_\odot$ the parametrisation is by Kroupa (2001, equation 12), whose results suggest tentative evidence that more metal-rich environments produce relatively more low-mass stars. Note that only the metallicity dependence is shown, but not the dependence on mass (Fig. 2) or density (Fig. 3).

Why is the IMF of major importance ?

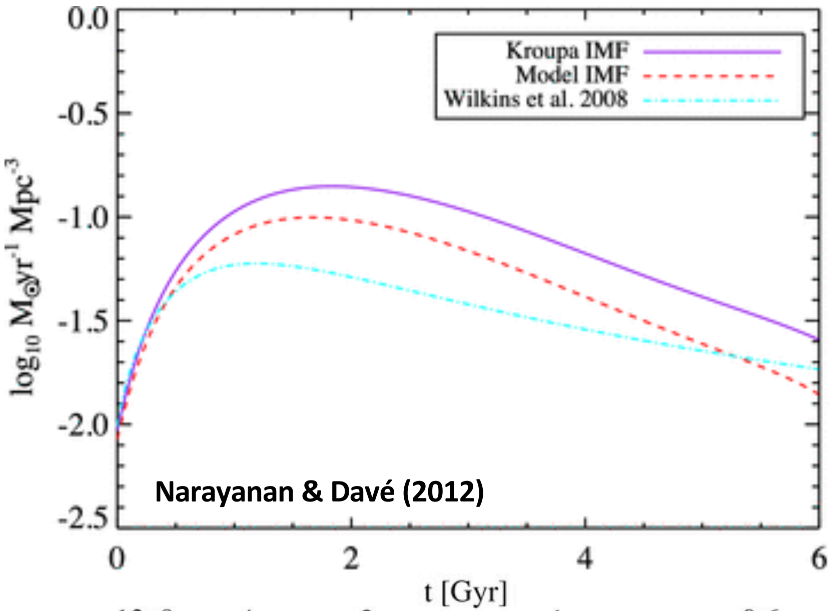
Several reasons that range from a basic better understanding of the star formation process to wider Xtragalactic / cosmological consequences:

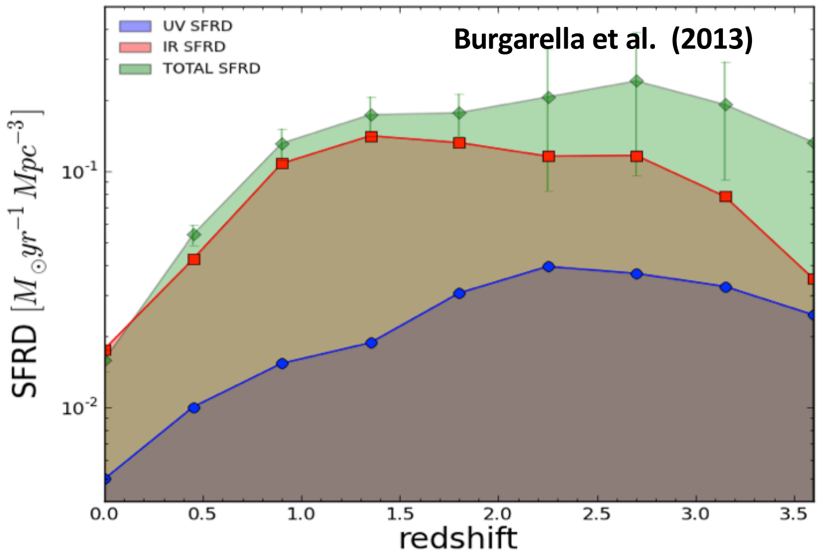
- Changing the IMF => changing the stellar mass (M_*) of the stars in a system (e.g. a galaxy)
- Changing M_* => changing the star formation rate ($SFR = M_* / \Delta t$) and therefore all parameters based on SFR:
    - Cosmic SFR density = density of star formation per unit volume of the universe at a function of redsfhit
    - Specific SFR = SFR / M_* => *very* important parameter that provides an estimate of the star formation activity in galaxies
- Changing the IMF => changing the amount of ionizing UV photons, which is important for reionization.

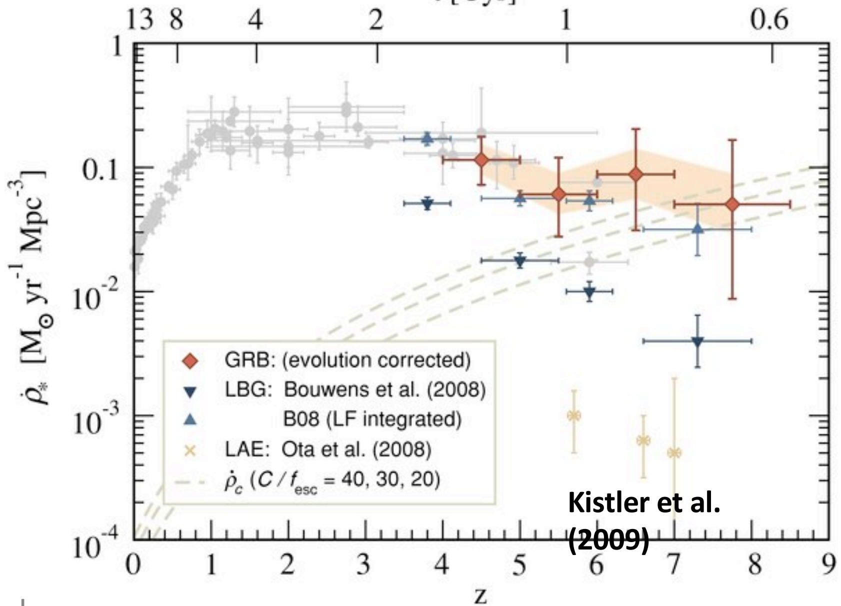
Cosmological implications of a stellar initial mass function that varies with the Jeans mass in galaxies

Evolution of cosmic SFR density.

The purple solid line shows model form from Hopkins et al. (2010), assuming a Kroupa IMF
With a varying IMF model, the SFR density decreases, as is shown by the red dashed line
The blue dash–dotted line shows the Wilkins et al. (2008) SFR density that would be necessary to match the observed present-day stellar mass density

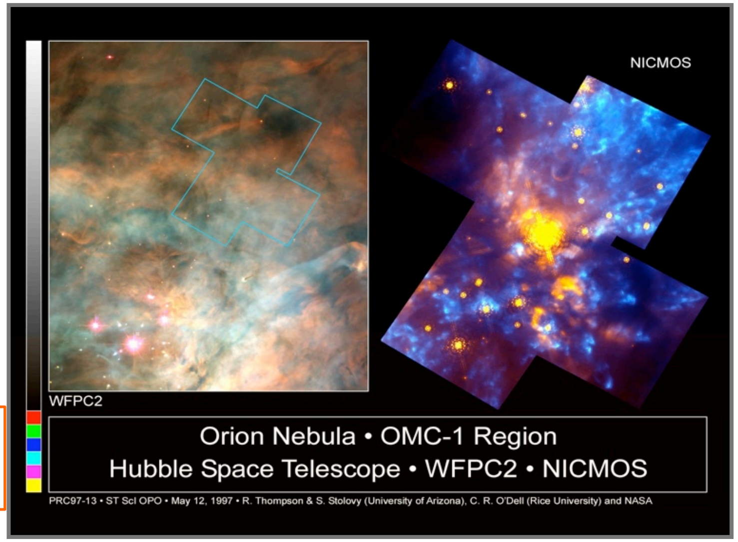






Giant molecular clouds (GMC)

Star formation takes place in cold, dense gas clouds: The molecular clouds. Stars form in groups or clusters. The largest GMC in Orion is about 1000 light years away. Hot young stars (25-50 million year old) ionize their surroundings and are therefore easily visible.

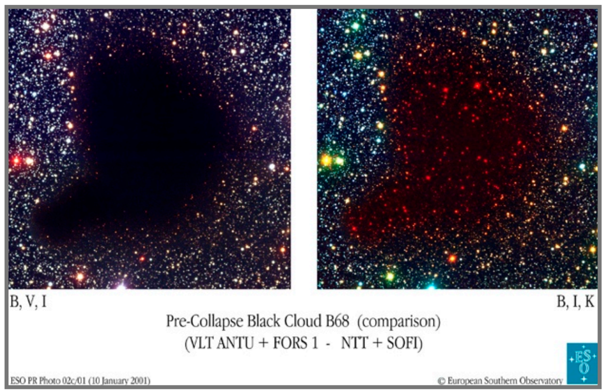


- *main characteristics:*
  - molecular hydrogen
  - ~1% of dust (Si and C)
  - organic and non-organic molecules ammoniac NH_3, formaldehyde H_2CO, acetylene HC_3N
  - ~ 10^5-10^6 solar masses
  - ~ 1000 atoms per cm^3
  - clumpy: 10^3-10^4 solar masses clumps
  - T ≈ 10 K

$\Rightarrow$ *Wien law* $\lambda_{max} = 2.9\times10^{-3} / T \sim 290\ \mu m$
$\Rightarrow$ *Observation in Far-IR (Herschel, ALMA, PdB)*

- *structure*
  - emission of the H_2 molecule but difficult because small dipole moment. Molecules with larger dipole moment (CO) are better but do not represent the mass...
  - dust emission (radio band). Interpretation difficult since many parameter unknown (departure from LTE, opacity, dust temperature, etc.)
  - IR cameras on large telescopes measure the absorption on thousands of background stars → map the dust distribution in the cloud

Bok globule: Barnard 68



Bok globules
Example characteristics:
- 410 light years away
- diameter ~ 12'500 AU
- T ~ 16 K
- ~ 2.1 solar masses
- in gravitational equilibrium with $P_{bound}=12.5 \times 10^{-12}$ Pa

Star Formation Sequence in brief

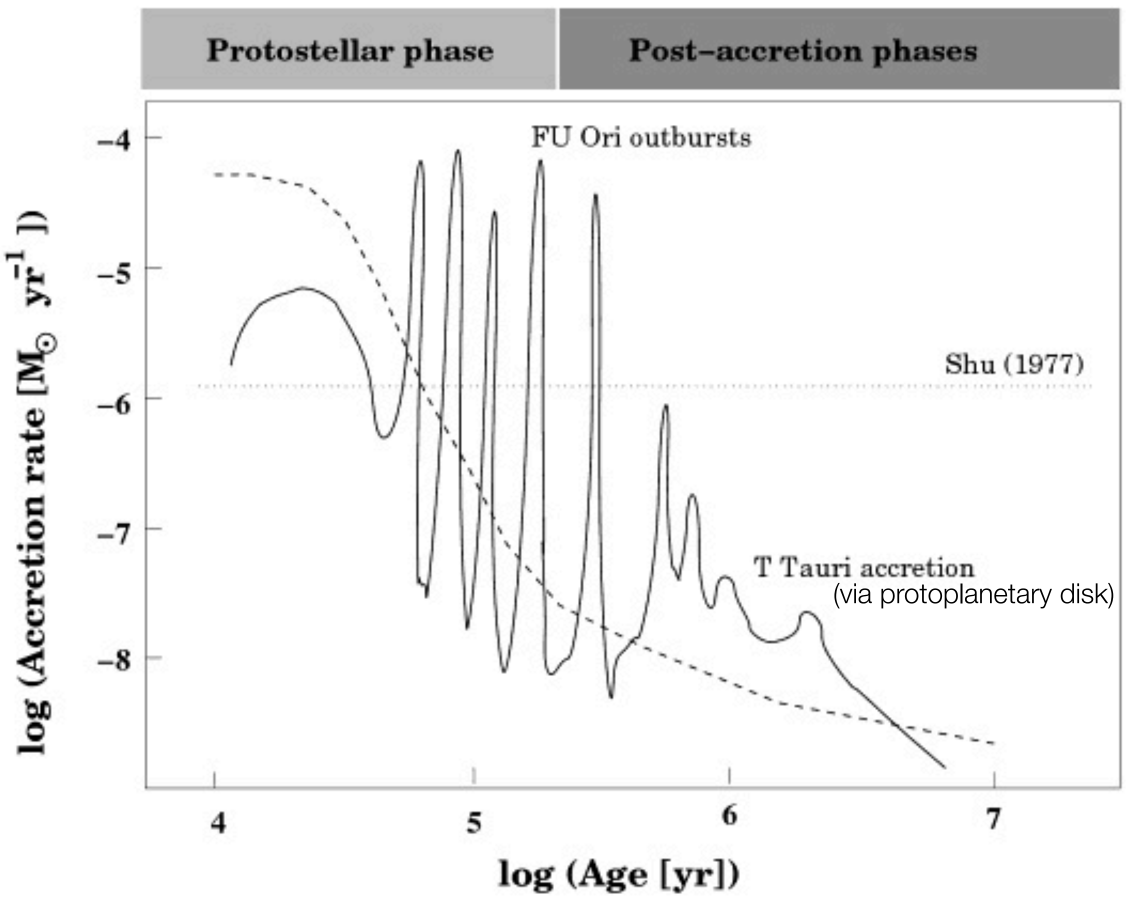
- Jeans instability => gas cloud collapse begins (~ 10^5 M_{sun})
- Isothermal collapse (free-fall time 10^8 years)
- Fragmentation of the cloud of gas
- Center of cloud becomes opticall thick: adiabatic compression (~ 10^{-13} g/cm^3) => thermodynamically isolated i.e., no heat transfered to the surroundings
- First hydrostatic core forms: ~ 170K (first equilibrium phase)
- H_2 dissociation (T ~ 2 000K) and second core collapse
- Ionized H in second core, dynamically stable again (10^{-3} M_{sun}, 20 000K, second equilibrium phase)
- Pre-Main Sequence contraction
- Zero Age Main Sequence (ZAMS): luminosity produced by H fusion, minimum mass 0.08 M_{sun}

Overview star formation sequence: Nomenclature

- Protostar
  - optically thick stellar core
  - forms during the end of the adiabatic contraction phase and grows during the accretion phase.
  - large accretion rates ~10^{-5} M_{sun}/yr

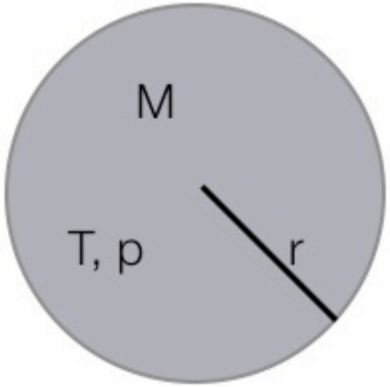
- Pre Main Sequence Star
  - visible in the optical
  - small accretion rates ~10^{-7} M_{sun}/yr
  - energy generation mainly via contraction

- ZAMS: Zero age main sequence
  - PMS contraction => center heats up
  - ~3 10^6 K: H burning ignites
  - contraction stops, energy production mainly via fusion



Jeans instability

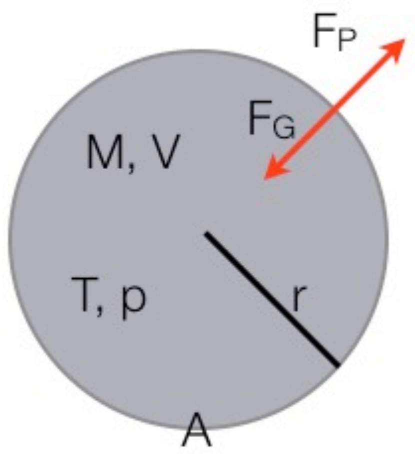
Now we want to understand under which conditions small perturbations of a gas cloud grow exponentially, leading to the collapse of the cloud (and, in the end star formation). The so-called Jeans instability describes the gravitational instability of a self-gravitating gas cloud. There are several ways to derive such a criterion, with increasing complexity.



The initially stable, static cloud can get initially compressed a bit (perturbed) by a shock wave due to a nearby supernova, passing spiral arms of the galaxy, ..

Jeans Instability from force balance

The simplest, order of magnitude estimation can be obtained by force balance arguments.



Collapse occurs if the inwards directed gravitational force is bigger than the outwards directed pressure force.

Consider mean unit forces per volume.

$$\frac{F_P}{V} \sim \frac{F_P}{Ar} \sim \frac{p}{r} \qquad \frac{F_G}{V} \sim \frac{GM^2}{r^2 V} \sim \frac{GM\rho}{r^2} \sim G\rho^2 r$$

Force balance $\frac{p}{r} = G\rho^2 r$ so $r = \sqrt{\frac{p}{G\rho^2}}$ critical maximal radius that allows stability.

We note: $r \propto \sqrt{p}$ higher pressure=>more stability

$r \propto \frac{1}{\rho}$ higher density=>less stability

with $p = \frac{k}{\mu m_H}\rho T \; = c^2\rho$

$$\boxed{r = \frac{c}{\sqrt{G\rho}}}$$

dense, cold clouds are unstable

Jeans Instability (from the more precise Virial Theorem Method)

The Jeans length is from this analysis

$$\boxed{\lambda_J = \sqrt{\frac{45}{16\pi}}\frac{c}{\sqrt{G\rho}}} \propto \sqrt{\frac{1}{\rho}}$$

The numerical constant is very close to unity (ca. 0.95), so we find a similar result as before.

The corresponding Jeans mass is

$$\boxed{M_J = \frac{45}{16}\sqrt{\frac{5}{\pi\rho}}\left(\frac{kT}{Gm_H\mu}\right)^{3/2}} \approx 4.89 \times 10^{21} \left(\frac{T}{\mu}\right)^{3/2} \rho^{-1/2} \text{ [kg]} \propto \sqrt{\frac{T^3}{\rho}}$$

Using typical values $\mu = 2.35$ $T = 10\mathrm{K}$ $\rho = 10^{-19}$ kg / m^3

we find $M_J \approx 69 M_\odot$ and $\lambda_J \approx$ 3.9 pc ≈ 800 000 AU

The collapse thus begins with large masses.

This is more than the typical mass of a single star (less than 1 M_{sun}). This indicates that during the collapse, only part of the gas ends up in stars, and that the cloud fragments during collapse. Thus many stars form out of one collapsing cloud, which means that young stars get born in clusters.

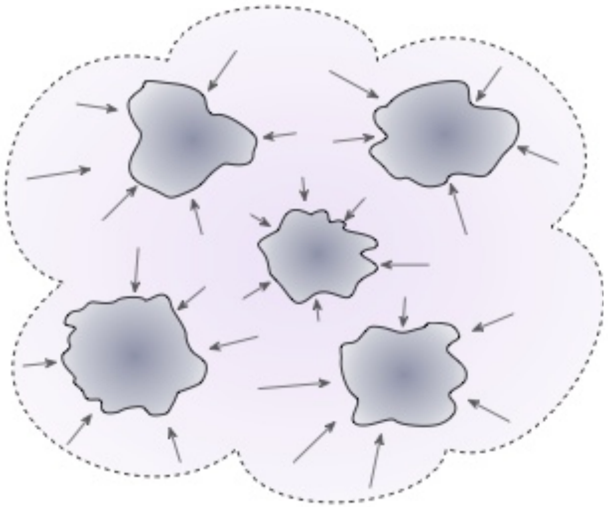
G = 6.67 x 10^{-11} m^3 kg^{-1} s^{-2}, k_B = 1.38 x 10^{-23} J K^{-1} and m_H = 1.67 x 10^{-27} kg

Fragmentation

During the collapse, ρ increases. As long as the density still remains adequately low for the cloud to be transparent, the released thermal energy is radiated into the universe and the temperature remains approximately constant. As

$$\searrow\ M_J \propto \sqrt{\frac{T^3}{\rho}}\ \begin{matrix}\rightarrow\\ \nearrow\end{matrix}$$

suggests, this leads to a decrease of the Jeans mass. In particular, sub-sections of the cloud suddenly surpass their own Jeans limit and start collapsing on their own. As also t_{ff} is smaller for higher densities, these sub-collapses proceed faster. This clearly leads to fragmentation.



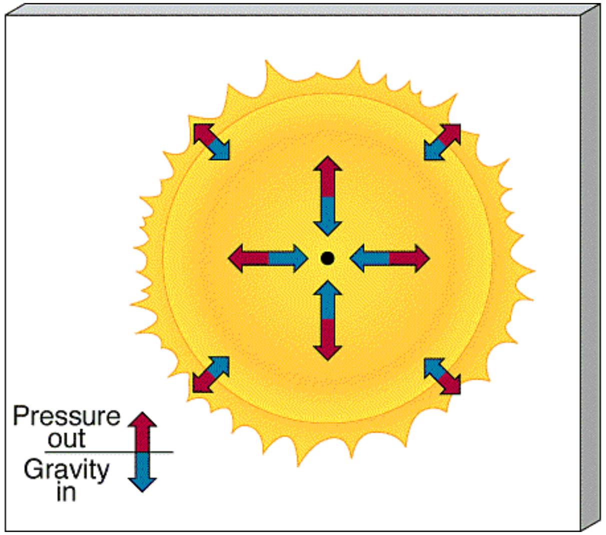
Such places might be the origin of later individual star formation, as they decouple.

• Star formation is governed by two dominant influences:

  ○ gravity, the universal force that causes all matter to attract

  ○ heat.

• Gravity's pull overcomes the random gas motions within an interstellar cloud, initiating a contraction phase that will last approximately 100,000 years and culminate in the formation of a star.



• During this collapse, the gas density increases. Collisions between atoms and molecules become more frequent and the gas temperature rises.

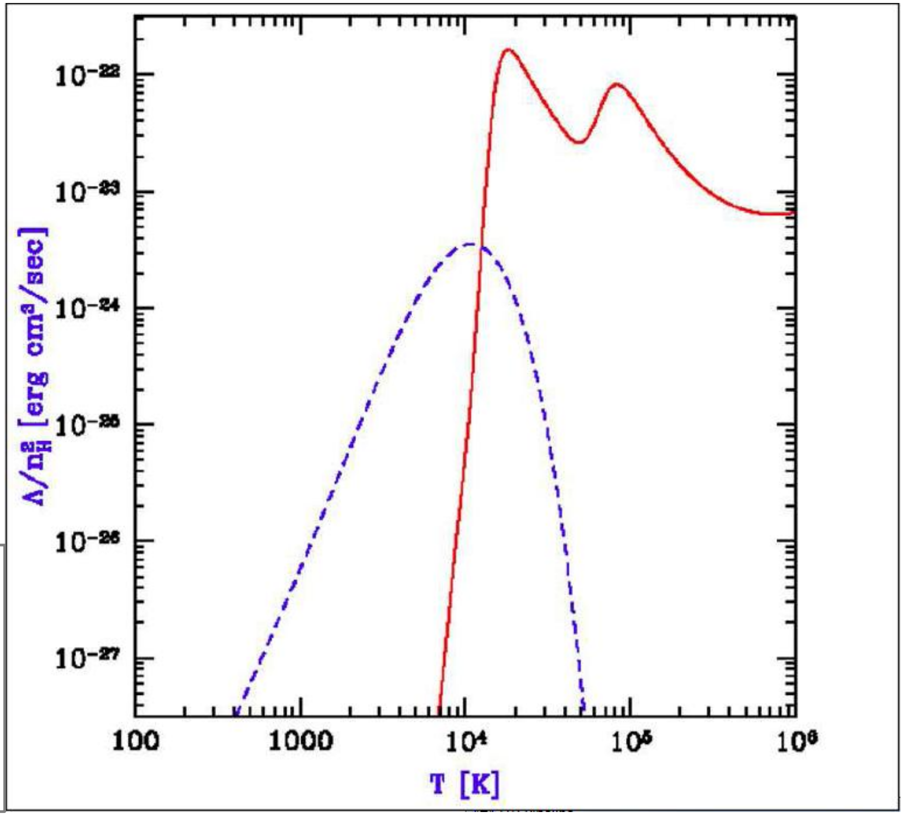
• The heating of the collapsing cloud poses a significant problem: a heated gas wants to expand, the cloud collapse could be halted or even reversed unless heat is effectively and continuously removed from the cloud.

- One process which provides significant cooling involves collisions between molecules.
- When two molecules collide, they convert some of their thermal (kinetic) energy into a form of potential energy. The energy can be stored in the molecule either by simple rotation or by internal vibration or even by lifting one or more electrons into a "higher" less bound orbit around the atoms in the molecule.
- This energy can be later released by the emission of a photon of a particular energy that is characteristic of these molecular species. Photons that escape the cloud carry this energy with them, thus helping to cool the cloud.
- Atoms and molecules are considered to be good coolants if
    - they quickly emit photons following a collision
    - they are present in large enough quantities that a significant number of photons are emitted.
- In this way the collapse of an interstellar cloud is tied to the chemical composition of that cloud.

• Hydrogen and helium are, by far, the most abundant elements in interstellar clouds.

• However, H and He are very poor coolants because they cannot be collisionally induced to emit photons at the low gas temperatures characteristic of molecular clouds.

• A large fraction of the total cooling is produced by a few other atoms and molecules, notably gaseous water (HO), carbon monoxide (CO), molecular oxygen (O), and atomic carbon (C).



Cooling rate of primordial gas as a function of temperature. The solid line represents the contribution from atomic hydrogen and helium and the dashed line represents the contribution from molecular hydrogen. At temperatures below 10^4 K cooling is provided by H2, which is a poor coolant, but at T> 10^4 K more efficient atomic hydrogen line cooling comes to play. Courtesy: Bromm (2012)

From: http://www.cfa.harvard.edu/swas/science1.html

Cooling

- Collisional excitation followed by the emission of an IR photon

Due to their thermal velocity, molecules collide all the time, which can lead to an excitation of an electron. At disexcitation, a photon is radiated, taking away the energy. (We assume here that the cloud is optically thin).

$$A + B \rightarrow A^{*} + B$$
$$A^{*} \rightarrow A + h\nu$$

-Frequent collisions (abundant partners)
-Excitation energy comparable to or less than thermal kinetic energy
-High probability of excitation during collision
-Photon emission before the next collision
-No re-absorption of the photon (low optical depth of gas to line emission)

Heating and cooling III

Cooling is in general described by a cooling function $\Lambda(T)$. Its exact value depends on the detailed chemical composition, but an order of magnitude can be estimated from the reaction rate $<\sigma v>$ multiplied by the amount of energy lost in one collision. Taking $\sigma \approx 10^{-16}$ cm^2, $v \approx 1$ km/s, $\Delta E \approx 0.1$ eV $\approx 10^{-13}$ erg, we obtain a typical value of the cooling function: $\Lambda(T) \approx 10^{-24}$ erg / (cm^3 s).

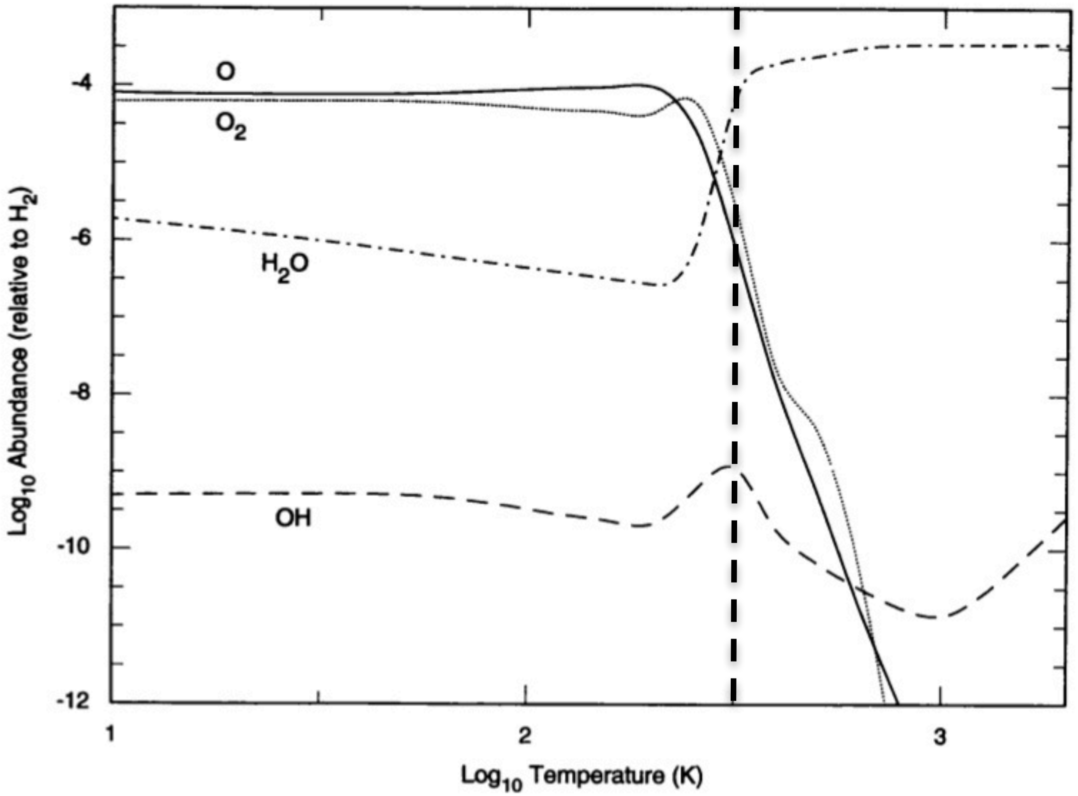
The cooling rate is then obtained by multiplying the cooling function with the abundance of booth cooling species.

$$\boxed{\Gamma_{CE} = \Lambda(T) n_1 n_2 =< \sigma v > \Delta E \; n_1 n_2 = \sigma \sqrt{\frac{8k\bar{T}}{\pi \bar{m}}} \Delta E \; \frac{\rho_1}{m_1} \frac{\rho_2}{m_2}}$$

The cooling time can be estimated (for identical species)as $\tau_{cool} = \frac{3/2nkT}{n^2 \Lambda}$

Coolants

To calculate the specific cooling rate, one must know the chemical composition of the gas.



Neufeld et al. 1995

Chemical networks

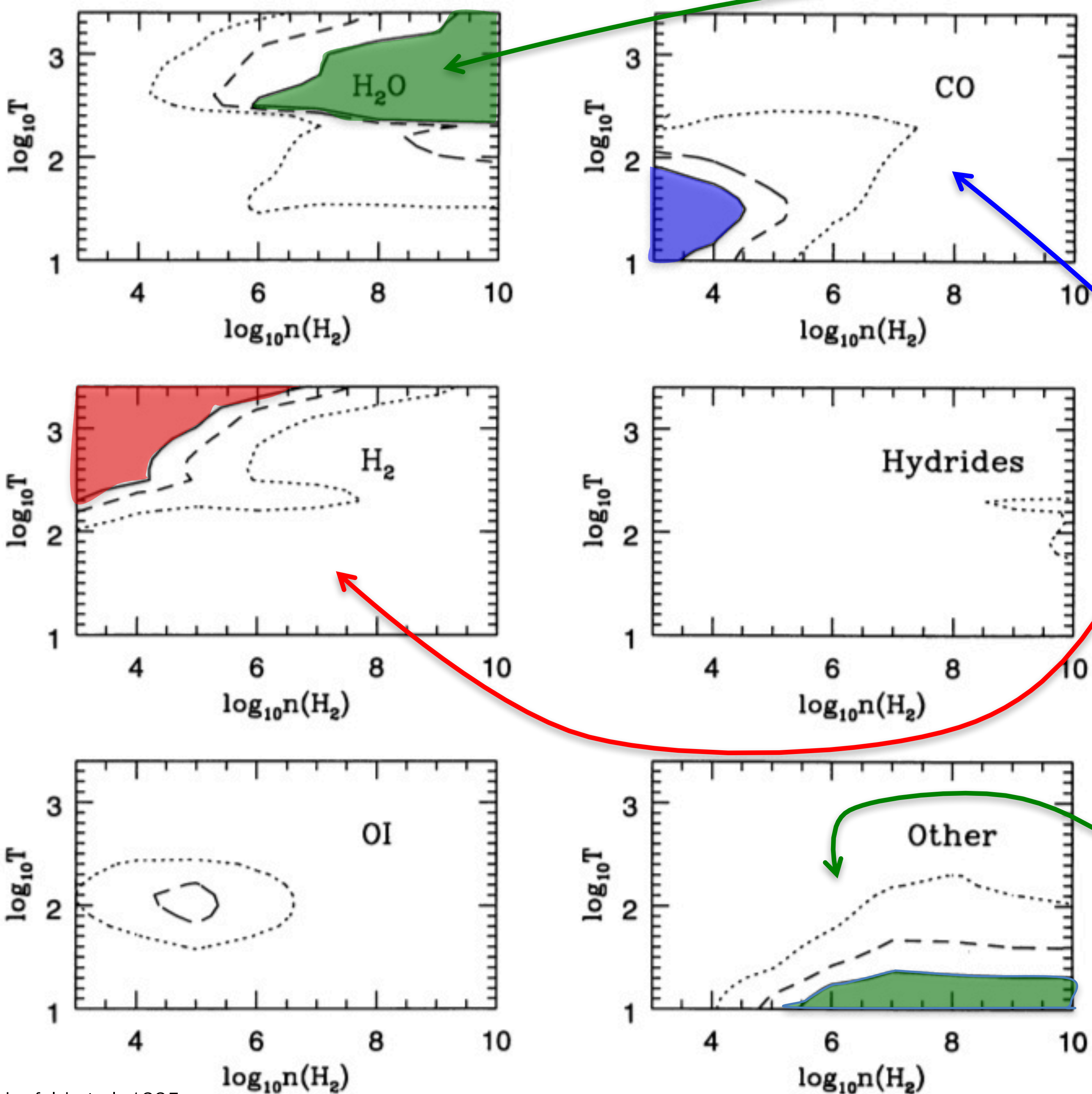
The figure shows the relative abundances of a molecule M, $x(M)=n(M)/n(H_2)$, for $n(H_2)=10^6$ cm^{-3}.

For T>500 K, all the oxygen not locked in CO, is in the form of water.

$$O + H_2 \rightarrow OH + H$$

$$OH + H_2 \rightarrow H_2O + H.$$

Contribution of coolants



Neufeld et al. 1995

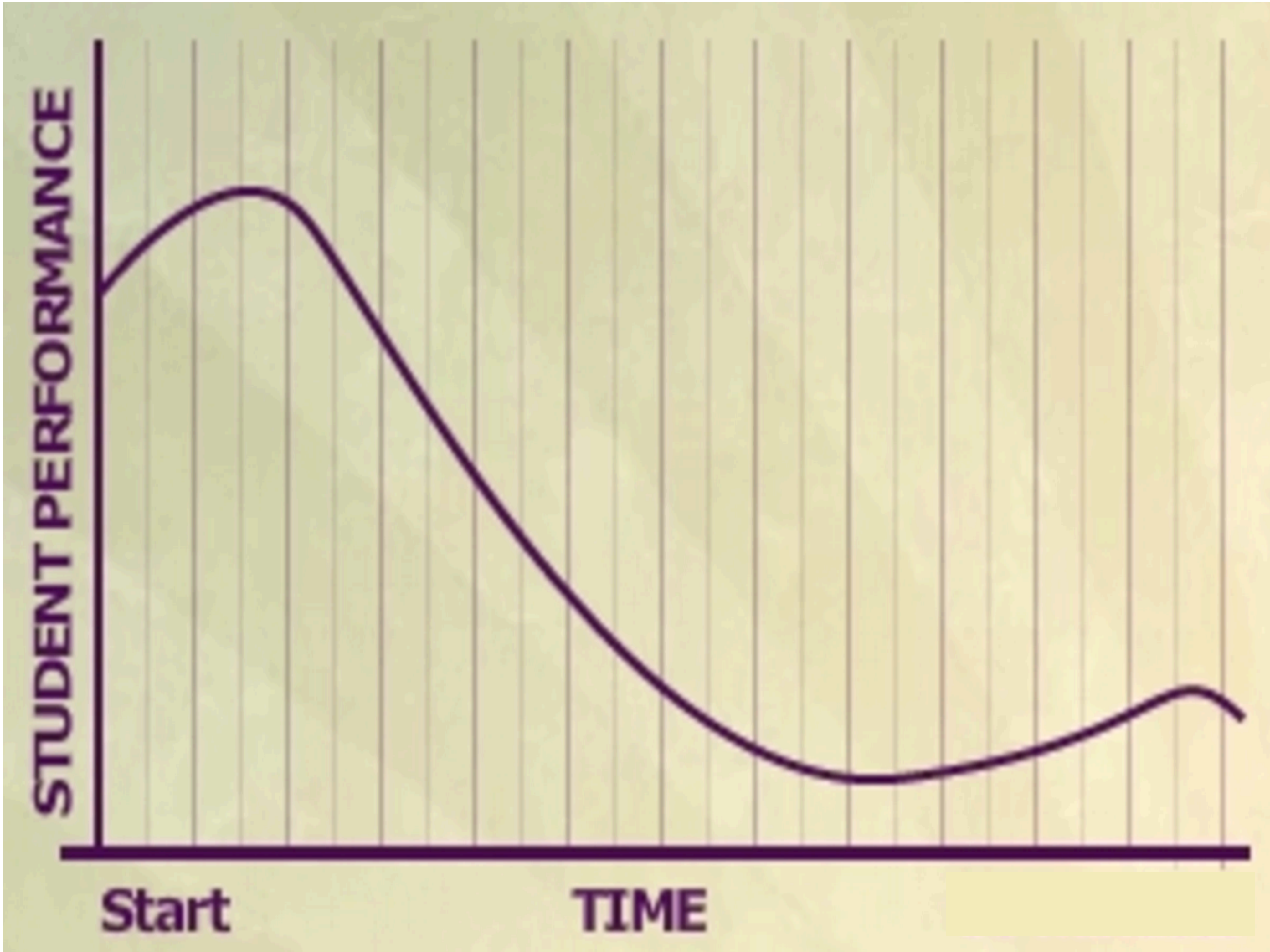
Fractions of the total cooling rate accounted for by emission of various coolants.

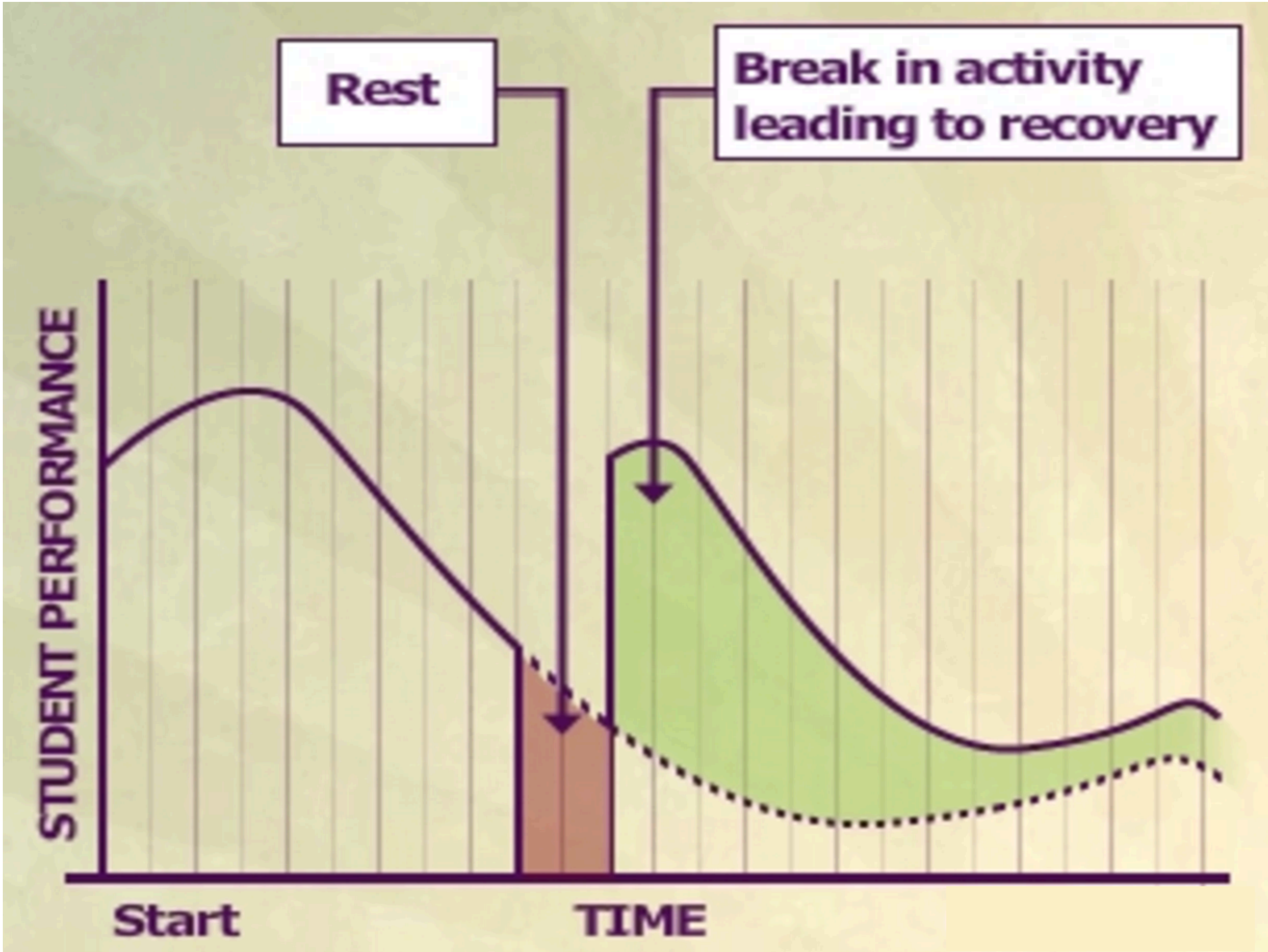
Contours correspond to
20% dotted
50% dashed
70% solid

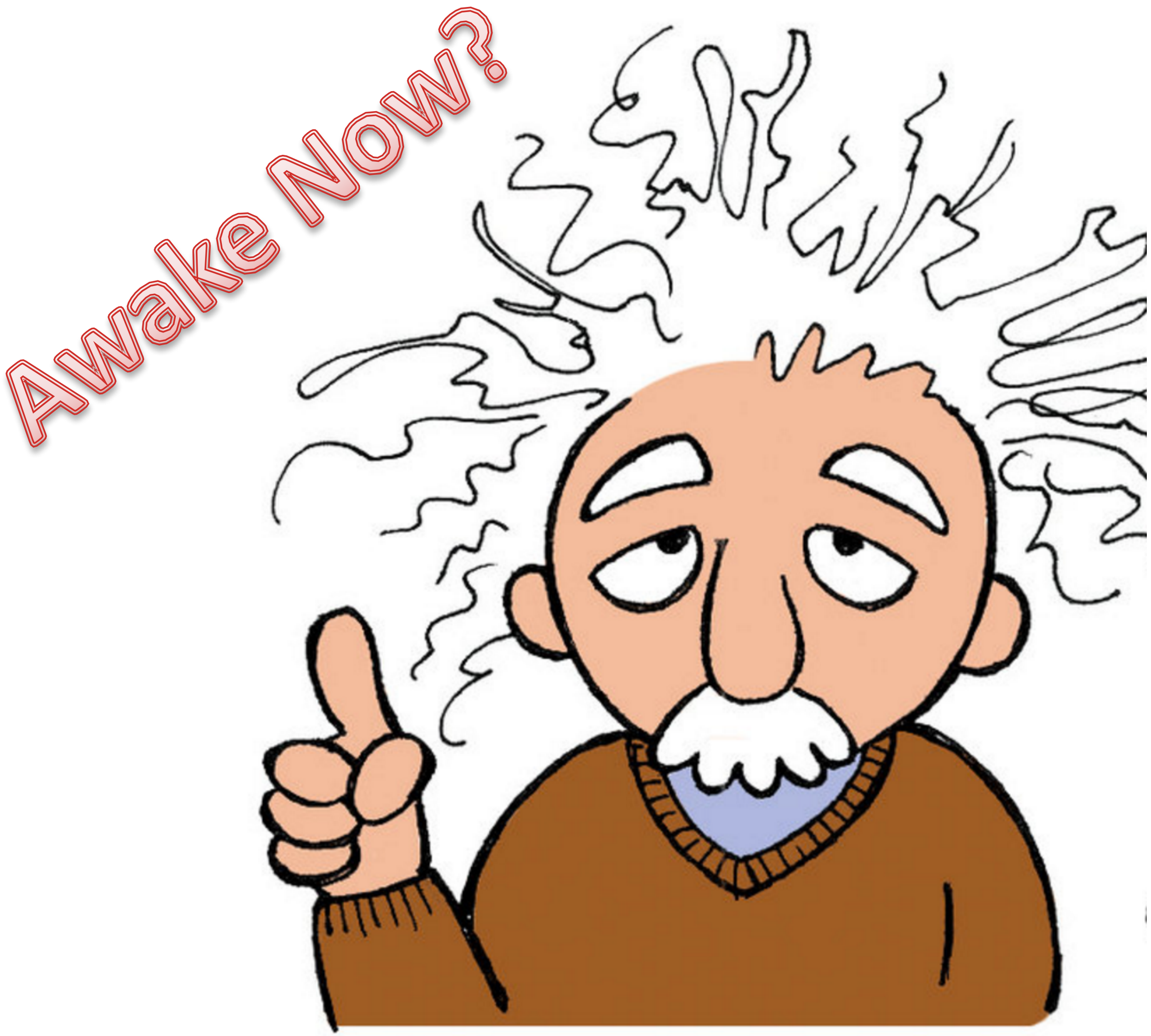
Notes

- H (and He) cannot be collisionally excited at low T. It is there a poor coolant.
- At lower densities and temperatures, CO and O are dominant.
- At high densities, water (high T) and a lot of other molecules (low T) become dominant.

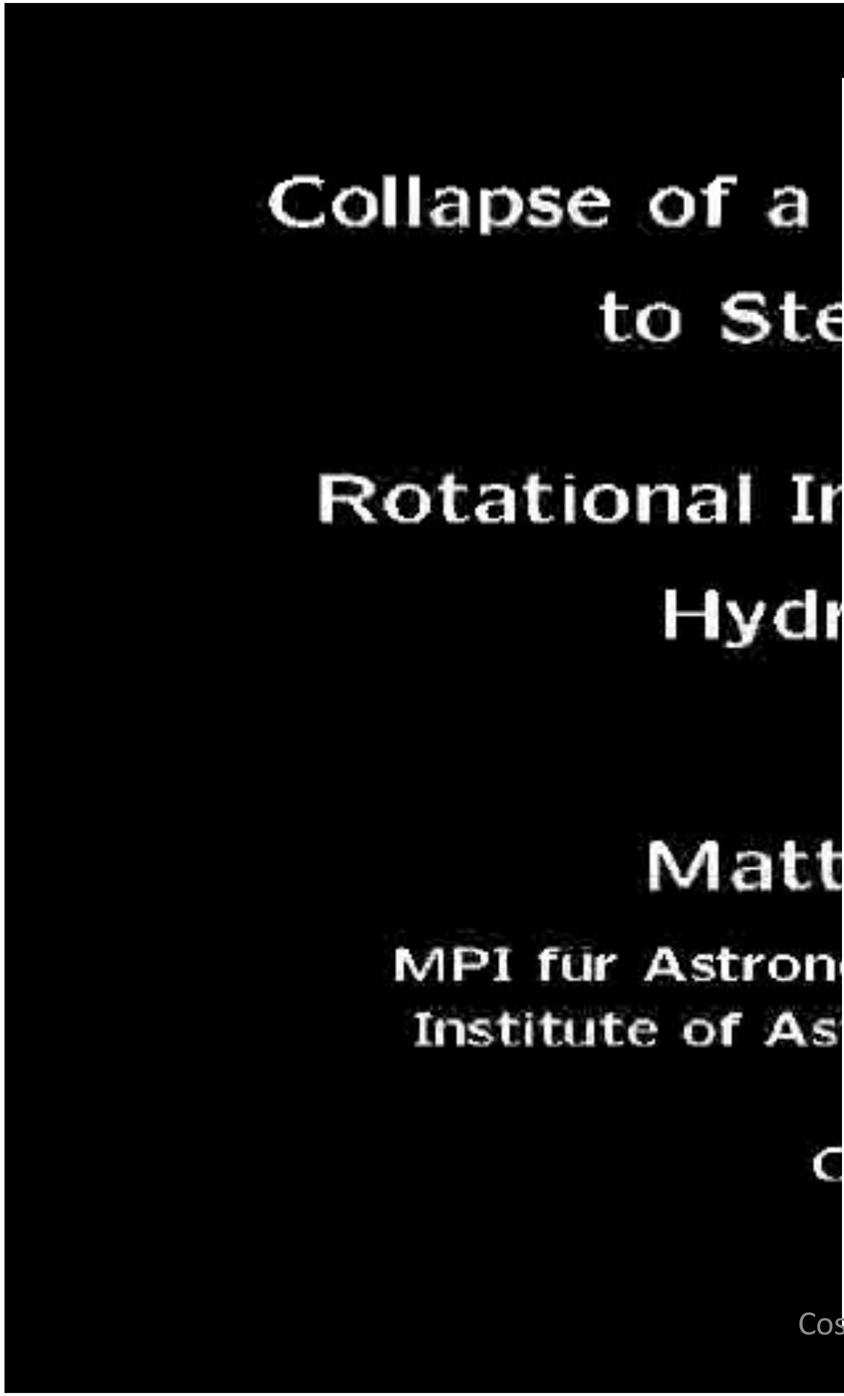
Numerical Simulations of Star Formation


STUDENT PERFORMANCE
Start
TIME


Rest
Break in activity leading to recovery
STUDENT PERFORMANCE
Start
TIME


Awake Now?

Numerical Simulations II



- The Formation of Stars and Brown Dwarfs and the Truncation of Protoplanetary Discs in a Star Cluster (Bate, Bonnell, & Bromm)
- The calculation models the collapse and fragmentation of a molecular cloud with a mass 50 times that of our Sun. The cloud is initially 1.2 light-years (9.5 million million km) in diameter, with a temperature of 10 K.
- The cloud collapses under its own weight and very soon stars start to form.
- Surrounding some of these stars are swirling discs of gas which may go on later to form planetary systems like our own Solar System.
- The calculation took approximately 10^5 CPU hours running on up to 64 processors on the UKAFF supercomputer. In terms of arithmetic operations, the calculation required about 10^16 FLOPS (i.e. 10 million billion arithmetic ops).

Numerical simulations V

Example 2:

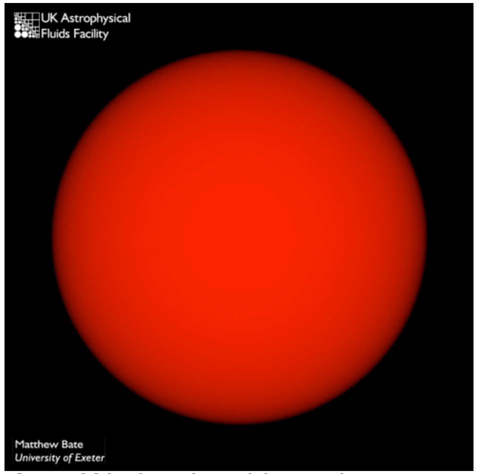
Bate 2009 : Hydrodynamic 3D simulation of a stellar cluster

Models the collapse and fragmentation of a 500 solar mass cloud. The calculation produces a cluster containing more than 1250 stars and brown dwarfs to allow comparison with clusters such as the Orion Trapezium Cluster.

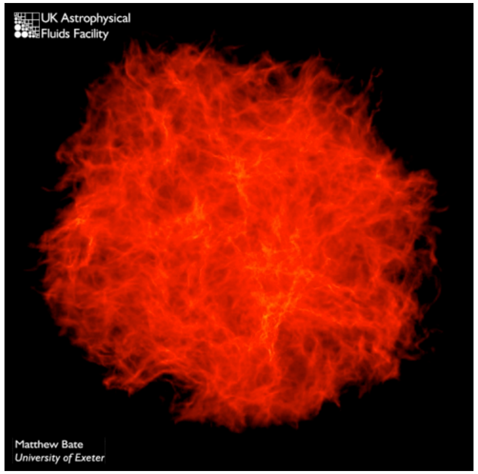
Parameter	Initial condition
R_0	0.8 pc
M_0	500 M_{sun}
v	supersonic turbulent
T	10 K
μ	2.46
T_{ff}	190 000 yrs
Stop	285 000 yrs
Radiation	not included
Magnetic fields	not included
Numerical method	SPH
Stars and BD	sink particles
Accretion radius	5 AU
Min. binary sep.	1 AU
EOS	piecewise polytropic
N particles	35 Mio

Numerical simulations VII

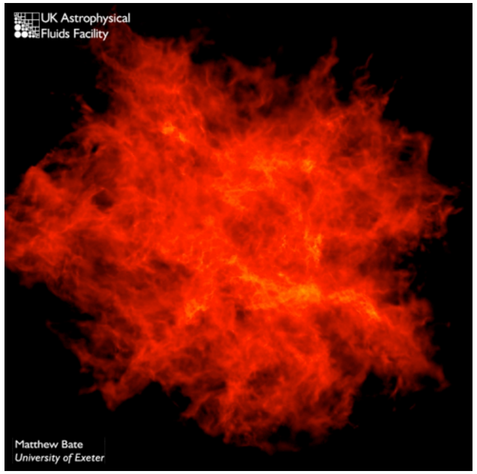
Bate 2009



0 yr: We begin with such a gas cloud, 2.6 light-years across, and containing 500 times the mass of the Sun. The images measure 1 pc (3.2 lightyears across).



38,000 yr: Clouds of interstellar gas are seen to be very turbulent with supersonic motions.

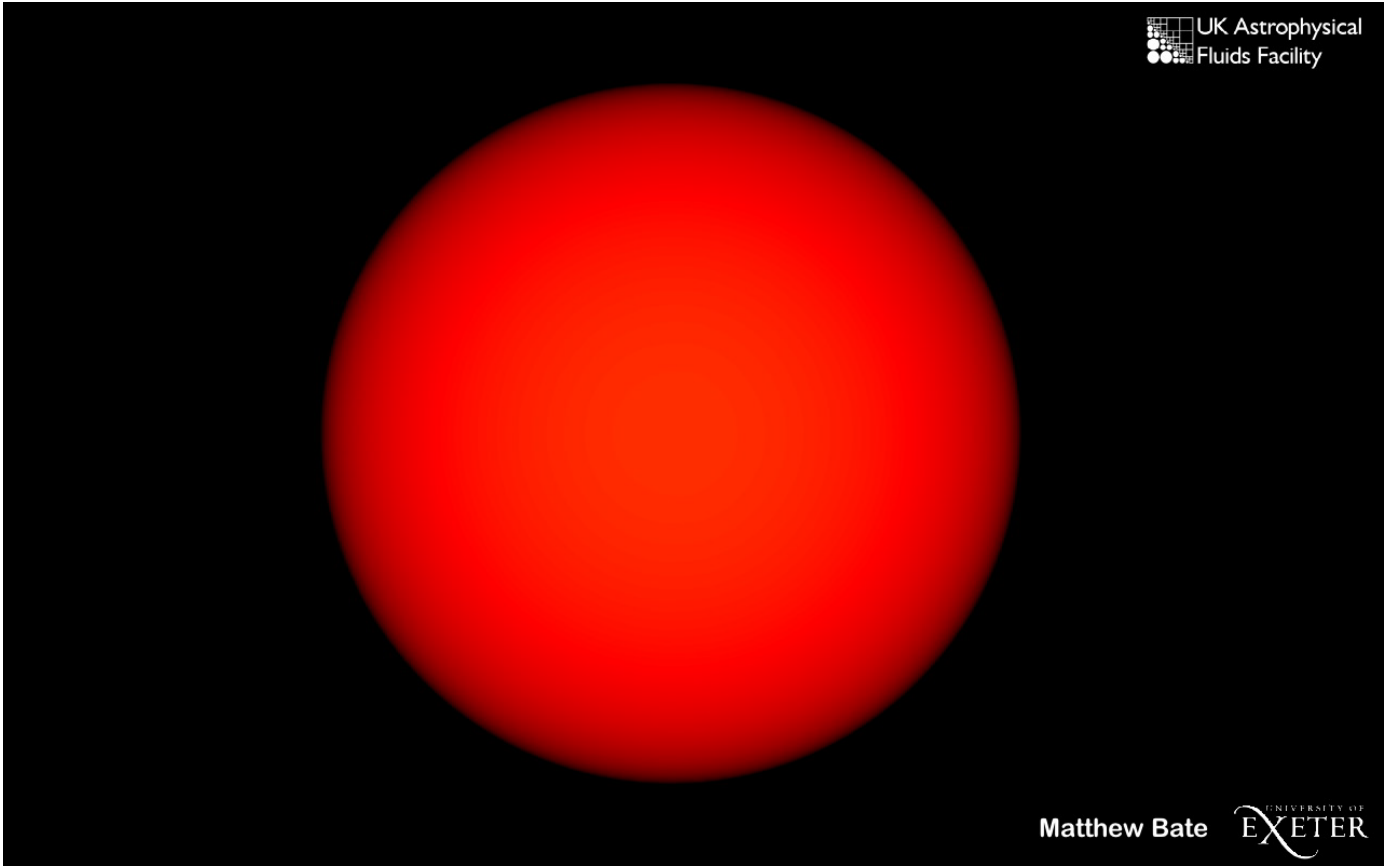


76,000 yr: As the calculation proceeds, the turbulent motions in the cloud form shock waves that slowly damp the supersonic motions.

152,000 yr: When enough energy has been lost in some regions of the simulation, gravity can pull the gas together to form dense "cores".

190,000 yr: The formation of stars and brown dwarfs begins in the dense cores. As the stars and brown dwarfs interact with each other, many are ejected from the cloud.

The cloud and star cluster at the end of simulation (which covers 210,000 years so far). Some stars and brown dwarfs have been ejected to large distances from the regions of dense gas in which the star formation occurs.


UK Astrophysical
Fluids Facility
Matthew Bate
UNIVERSITY OF
EXETER

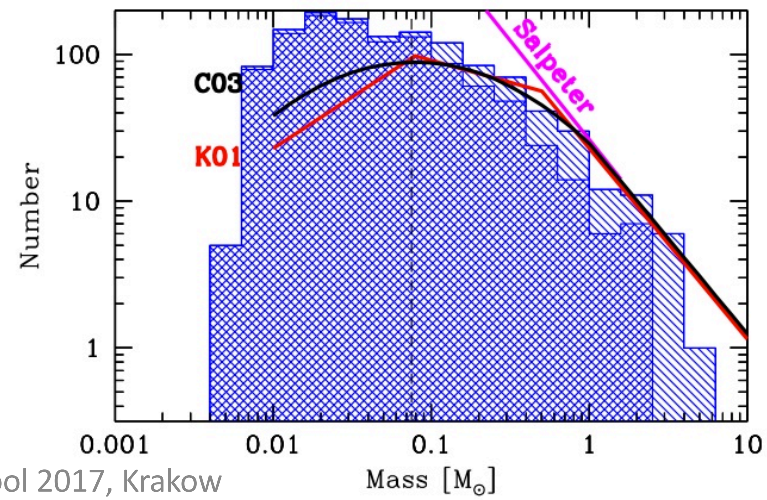
Numerical simulations VIII

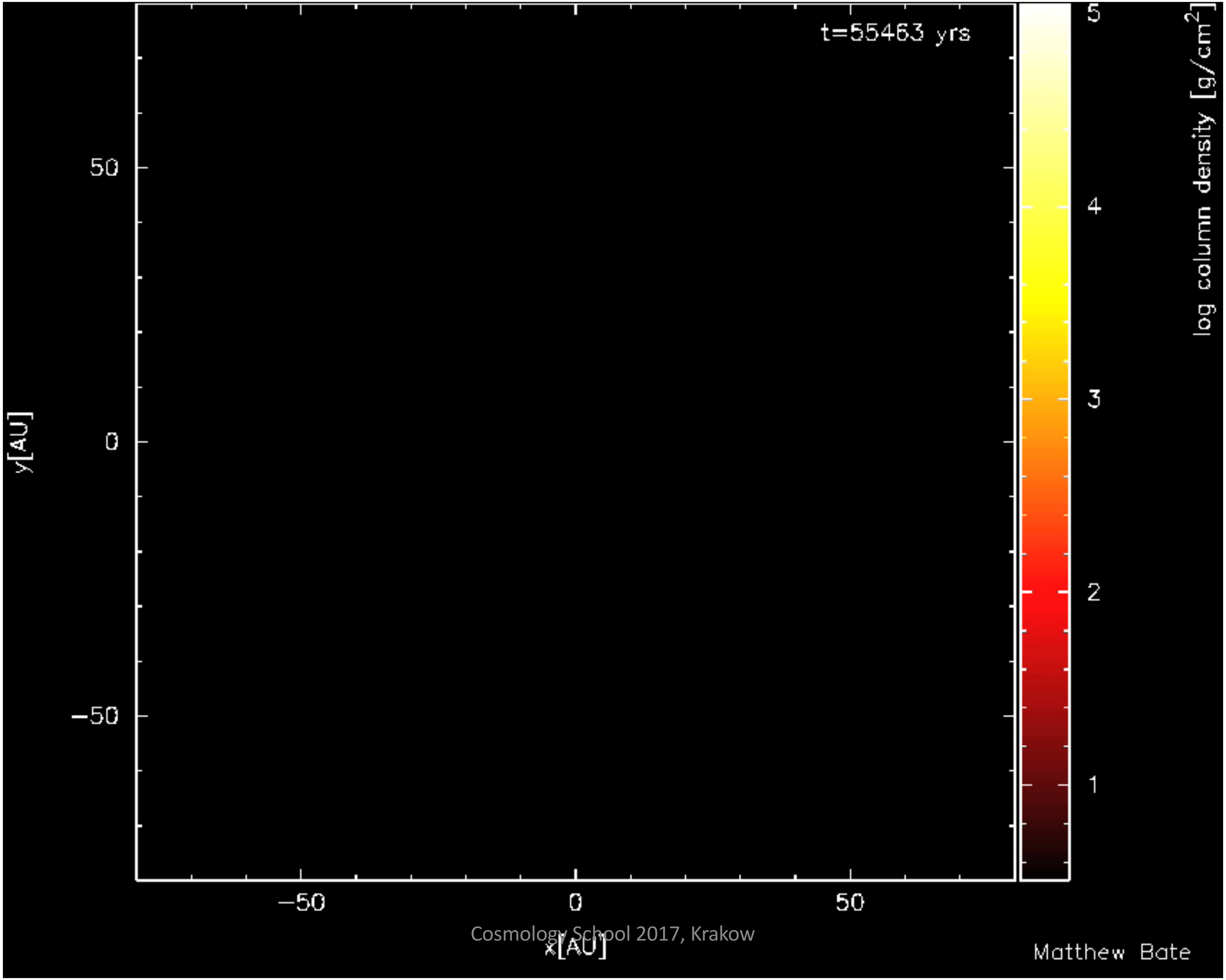
Main simulation results

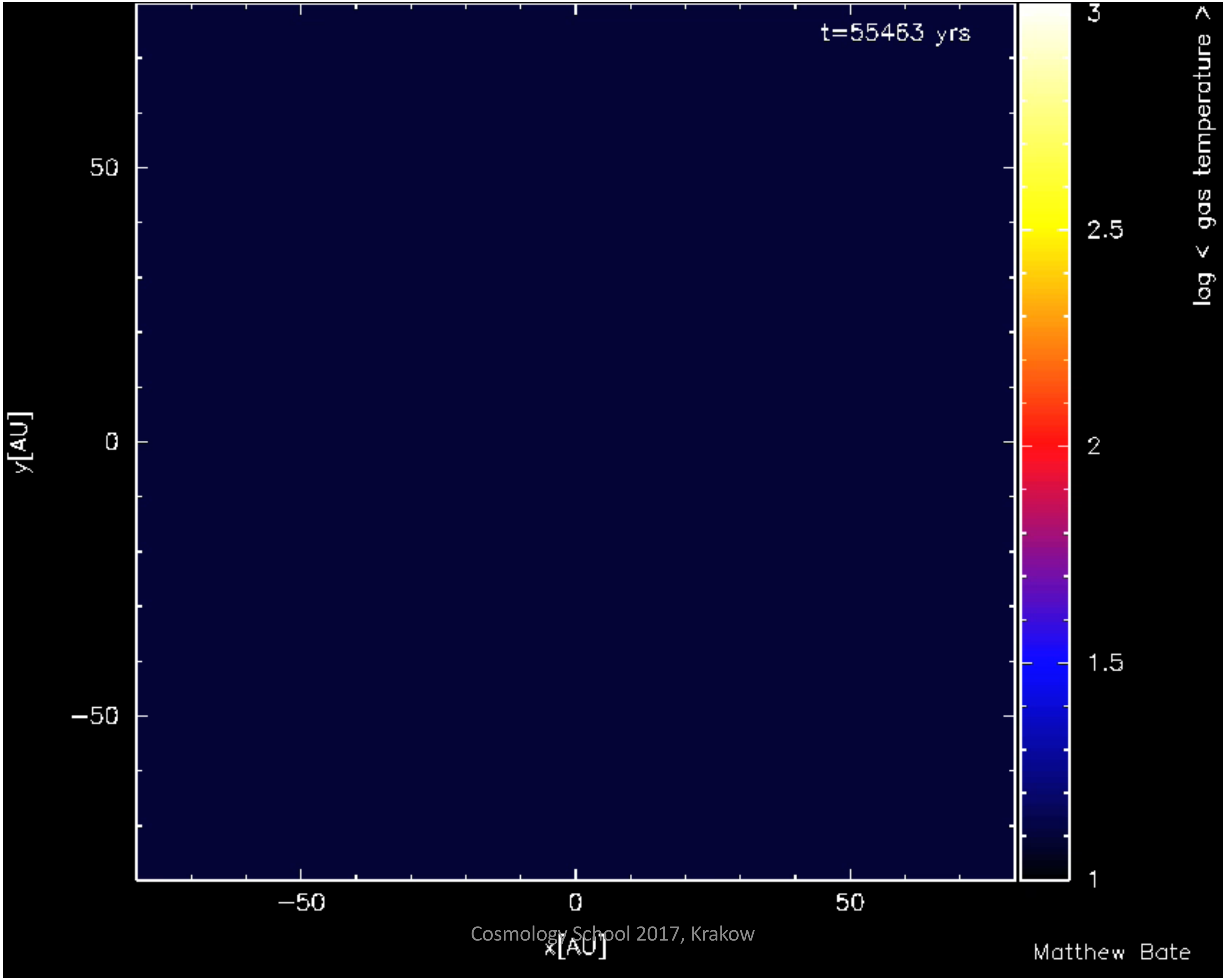
1) Since all sink particles (and thus stars/BD) are created from pressure-supported fragments, their initial masses are just a few Jupiter masses, as given by the opacity limit for fragmentation. Subsequently, they may accrete large amounts of material to become higher-mass brown dwarfs (< 75 $M_{Jupiter}$) or stars (>75 $M_{Jupiter}$), but all the stars and brown dwarfs begin as these low-mass pressure-supported fragments.

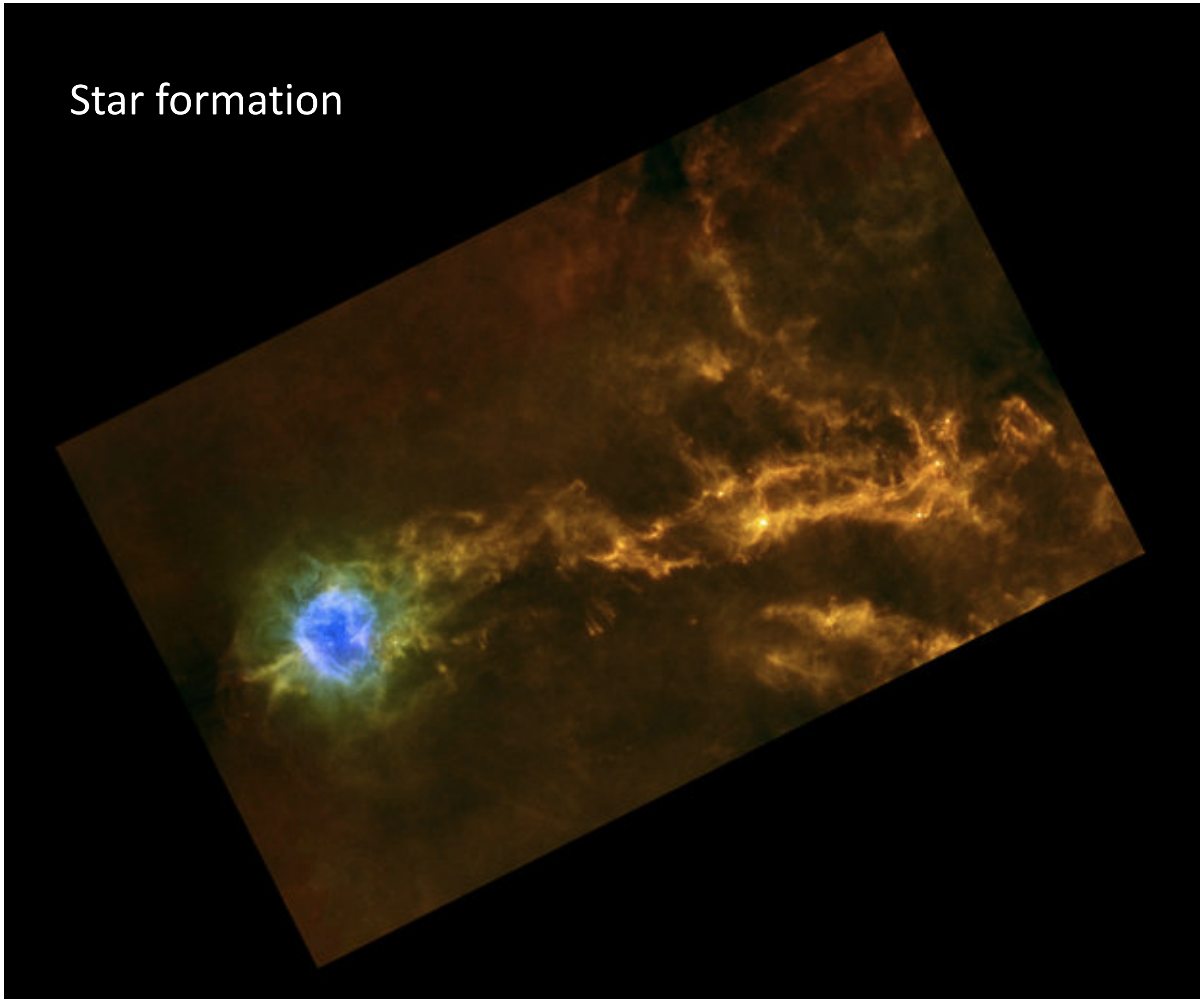
2) The IMF originates from competition between accretion and ejection which terminates the accretion and sets an object's final mass. Stars and brown dwarfs form the same way, with similar accretion rates from the molecular cloud, but stars accrete for longer than brown dwarfs before undergoing the dynamical interactions that terminate their accretion.

3) The calculations produce an IMF with a similar form to the observed IMF, including a Salpeter-type slope at the high-mass end but they over-produce brown dwarfs. It is likely due to the absence of radiative feedback and/or magnetic fields in the calculation.




t=55463 yrs
50
0
−50
y[AU]
−50
0
50
x[AU]
5
4
3
2
1
log column density [g/cm^2]
Matthew Bate


t=55463 yrs
y[AU]
50
0
−50
−50
0
50
x[AU]
log < gas temperature >
3
2.5
2
1.5
1
Matthew Bate



• ESA's Herschel space observatory has revealed that nearby interstellar clouds contain networks of tangled gaseous filaments.

• The filaments are huge, stretching for tens of light years through space and Herschel has shown that newly-born stars are often found in the densest parts of them.

• One filament imaged by Herschel in the Aquila region contains a cluster of about 100 infant stars.

•Herschel has shown that, regardless of the length or density of a filament, the width is always roughly the same.