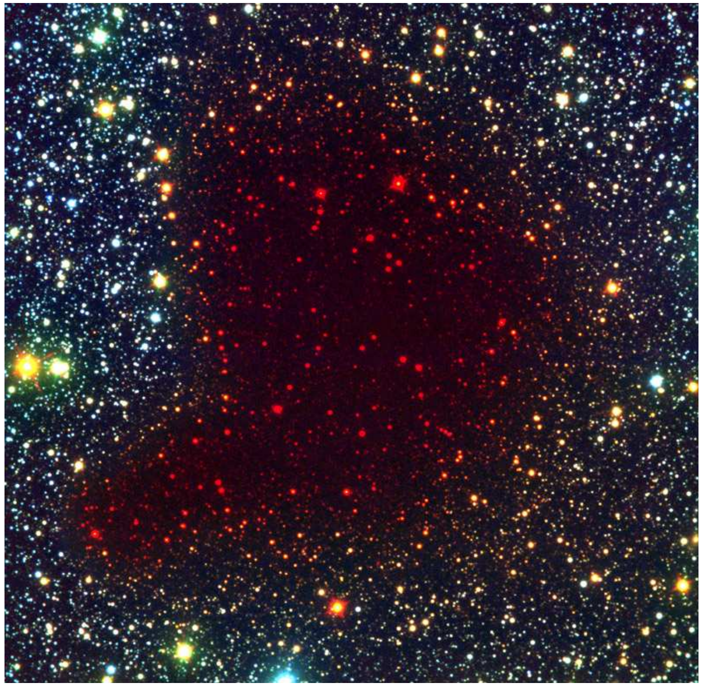
Stars: Formation, Structure, and Evolution

Optical View of B68 (ESO; VLT/FORS1)



IR View of B68 (ESO; VLT/FORS1 + NTT/SOFI)

Stellar Birth

Stars are born in “Giant Molecular Clouds”

Typical GMC parameters (e.g., Orion):

- large clouds: typical diameters 50–100 pc
- contain lots of molecular gas (H_2, CO, alcohol,. . .).
- typical temperatures: 10–20 K (coolest regions in the interstellar medium)
- typical particle densities $n \sim 10^6$–10^{10} cm^{-3}

Stars are born in groups out of collapsing Molecular Clouds.

Collapse triggered, e.g., by collisions of clouds or shocks caused by nearby supernovae.

Stellar Birth

Criterion for collapse: Cloud is instable, i.e., gravitation is stronger than thermal pressure.

In terms of thermal and gravitational energies, this means

$$\frac{3}{2}\frac{M}{m_\mathrm{p}}kT - \frac{3}{5}\frac{GM^2}{R} \leq 0 \tag{13.1}$$

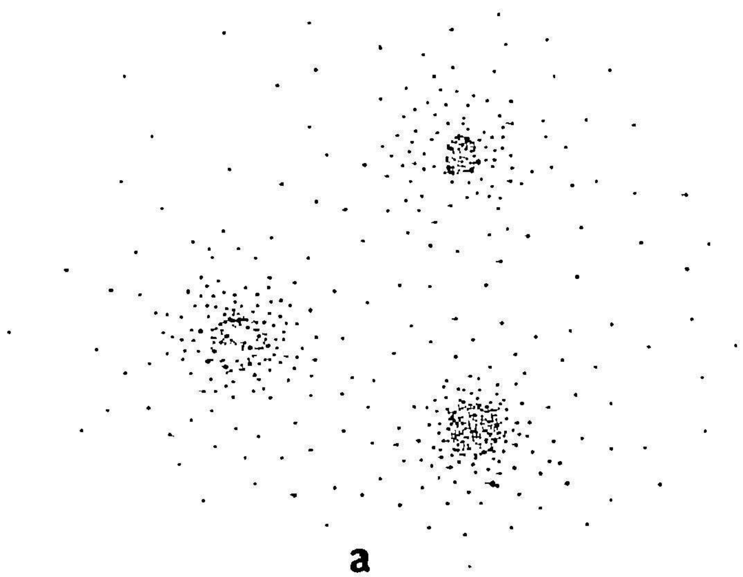
which can be expressed as

$$\frac{M}{R} \geq \frac{5}{2}\frac{kT}{Gm_\mathrm{p}} \quad \text{or} \quad \frac{4\pi}{3}\rho R^2 \geq \frac{5}{2}\frac{kT}{Gm_\mathrm{p}} \tag{13.2}$$

$\Longrightarrow$ Depends on R, collapse thus possible for

$$R > R_\mathrm{J} = \sqrt{\frac{15kT}{8\pi G m_\mathrm{p}\rho}} \sim \sqrt{\frac{kT}{Gm_\mathrm{p}\rho}} \tag{13.3}$$

where R_J is called the Jeans radius.



Shu et al. (1987, ARAA 25, 23, Fig. 7)

Stellar mass cores form from fragmentation of larger pieces.

Note: fragmentation only along B-field lines.

Stellar Birth

Plugging in typical typical numbers, i.e., $T \sim 50\,\mathrm{K}$, particle density $n = 10^5$ H-atoms cm^{-3} (=a mass density of $\rho = nm_\mathrm{p} \sim 1.7 \times 10^{-9}\,\mathrm{g\,cm^{-3}}$) gives $R_\mathrm{J} \sim 0.2\,\mathrm{pc}$.

For a given Jeans radius, the mass within R_J is the Jeans mass

$$M_\mathrm{J} \sim \frac{4\pi}{3}R_\mathrm{J}^3\rho$$

. . . which has typical values of 50–100 $M_\odot$, i.e., larger than one star!

In reality things are more complicated: ISM contains magnetic fields

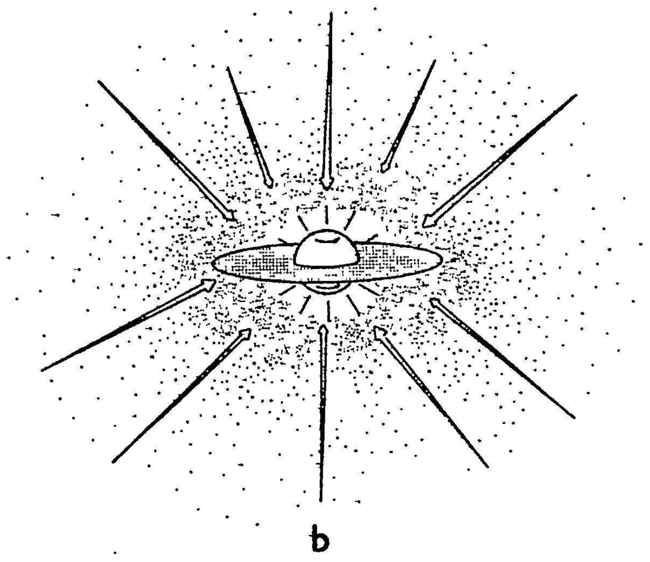
$\Longrightarrow$ Particle motion $\perp$ B-field lines difficult

$\Longrightarrow$ stops gas from collapsing.

This is good since Jeans formalism alone predicts too strong star formation.

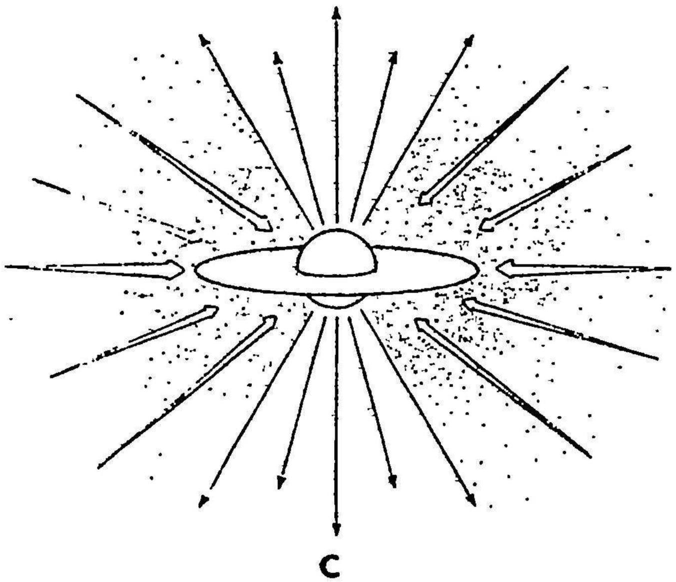
$\Longrightarrow$ Need star formation with magnetic fields

See Shu et al. (1987, Annual Reviews of Astronomy and Astrophysics 25, 23) for the gory details.



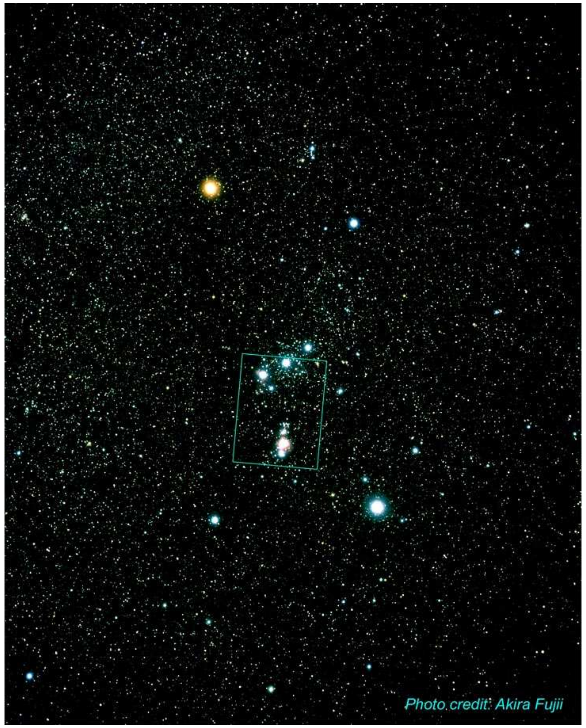
Shu et al. (1987, ARAA 25, 23, Fig. 7)

Protostar forms with surrounding disk (“inside out collapse”) once core hot enough to allow fusion ($T > 10^6$ K)



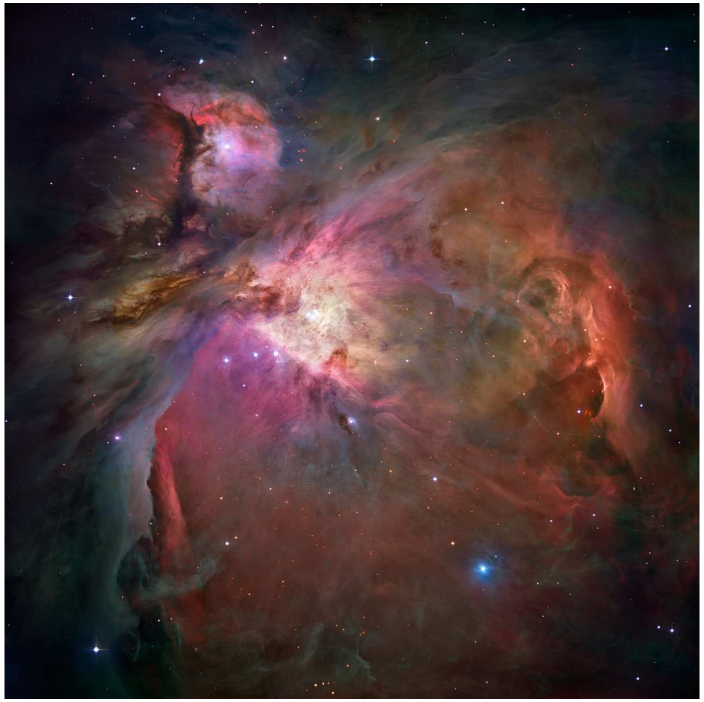
Shu et al. (1987, ARAA 25, 23, Fig. 7)

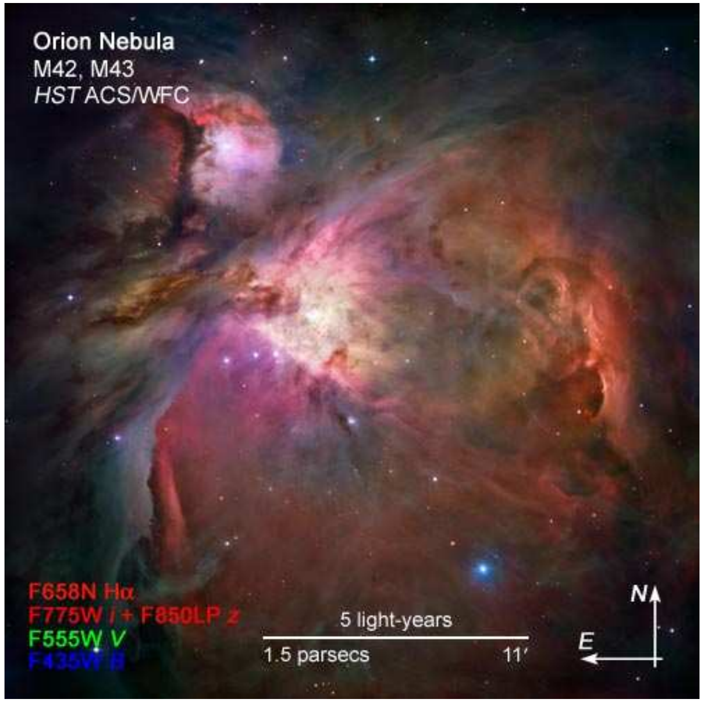
Stellar wind forms bipolar outflow



Movie Time:
`starformation_movies/orion_zoom.mo`
from
`http://hubblesite.org/newscenter/a`

Orion (Bayer's Uranometria; image ©USNO)

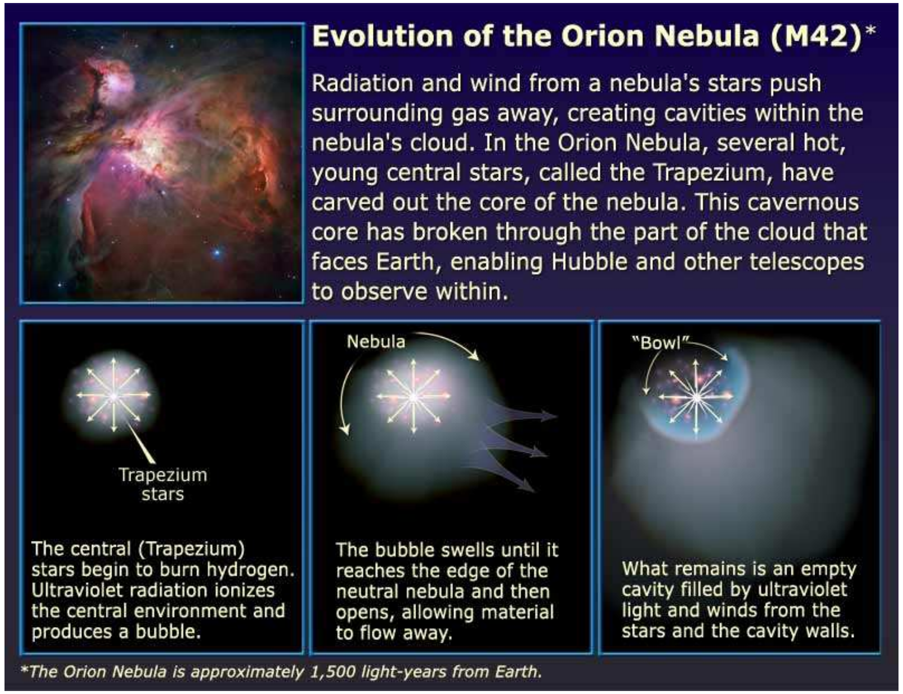




Orion Nebula; R. Croman

Orion Nebula; R. Gendler

Movie Time: `starformation_movies/hayden.mpg`
from `http://vis.sdsc.edu/research/orion.html`



Evolution of the Orion Nebula (M42)*

Radiation and wind from a nebula's stars push surrounding gas away, creating cavities within the nebula's cloud. In the Orion Nebula, several hot, young central stars, called the Trapezium, have carved out the core of the nebula. This cavernous core has broken through the part of the cloud that faces Earth, enabling Hubble and other telescopes to observe within.

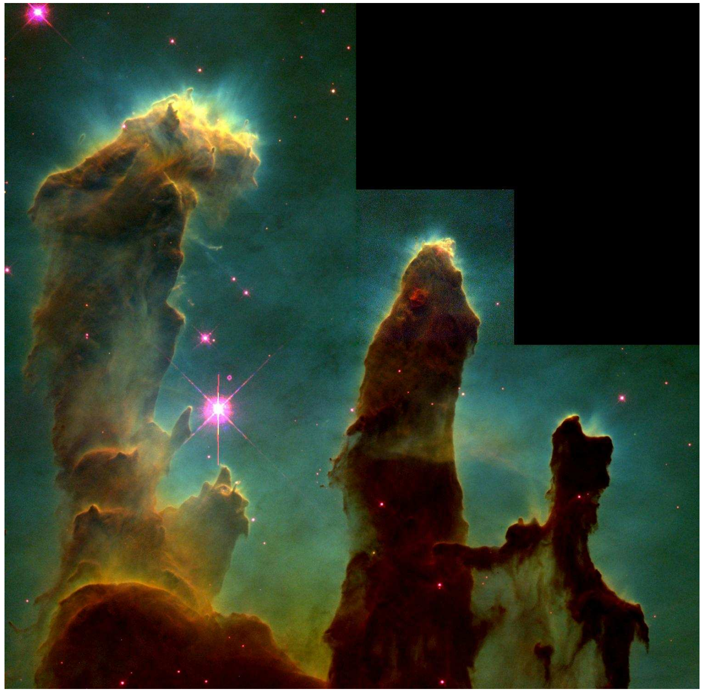
The central (Trapezium) stars begin to burn hydrogen. Ultraviolet radiation ionizes the central environment and produces a bubble.

The bubble swells until it reaches the edge of the neutral nebula and then opens, allowing material to flow away.

What remains is an empty cavity filled by ultraviolet light and winds from the stars and the cavity walls.

The Orion Nebula is approximately 1,500 light-years from Earth.

STScI

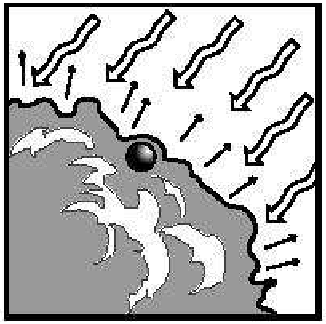


“pillars of creation” in Eagle Nebula (M16)

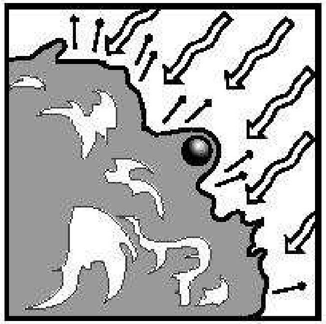
Photo Credit: *T. Rector (University of Alaska)*

Eagle-nebula (M16)

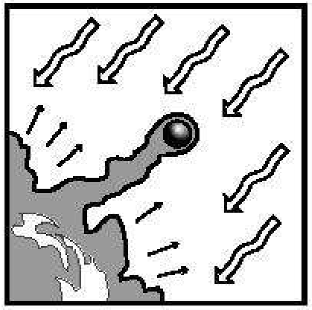
Stellar EGGs in M16



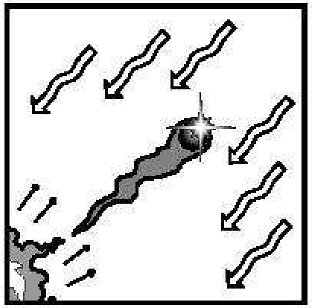
The surface of a molecular cloud is illuminated by intense ultraviolet radiation from nearby hot stars. The radiation evaporates material off of the surface of the cloud.



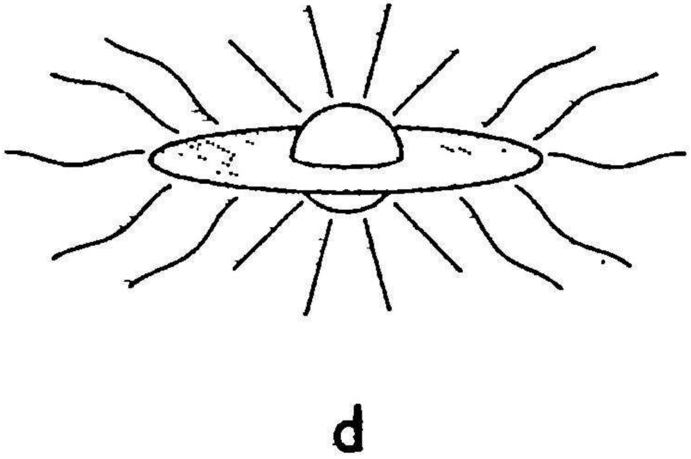
As the cloud is slowly eaten away by the ultraviolet radiation, a denser than average globule of gas begins to be uncovered



The EGG has now been largely uncovered. The shadow of the EGG protects a column of gas behind it, giving it a finger-like appearance.

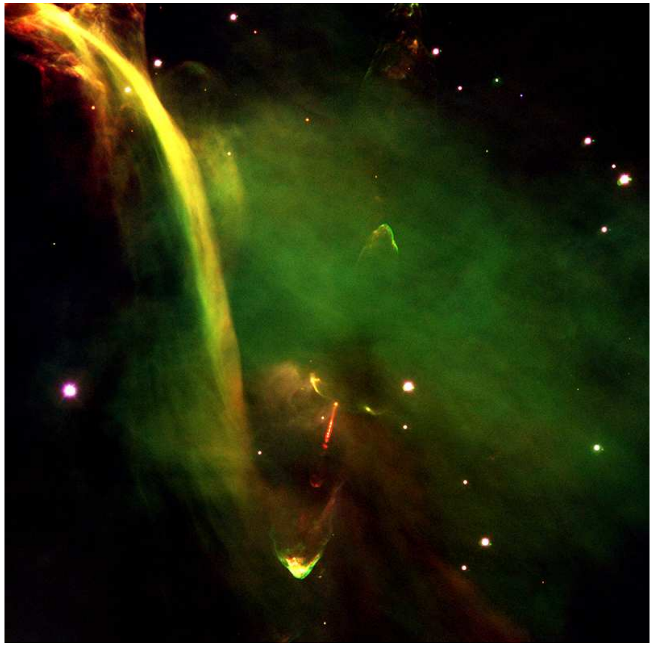


Eventually the EGG may become totally separated from teh molecular cloud in which it formed. As the EGG itself slowly evaporates, the star within is uncovered and may appear sitting on the front surface of the EGG.



Shu et al. (1987, ARAA 25, 23, Fig. 7)

Star has reached zero age main sequence (ZAMS) plus circumstellar disk.
Some disks produce fast collimated outflows (jets): Herbig Haro Objects

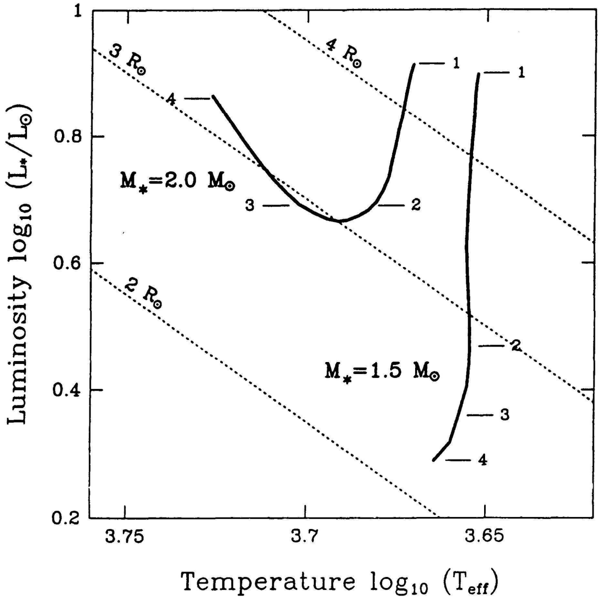


HH34 in Orion (ESO VLT KUEYEN/FORS2)

Herbig Haro Objects: shocks and jets/outflows produced during formation of stars.

Pleiades (R. Gendler; $d = 150\,\mathrm{pc}$, diameter: $5\,\mathrm{pc}$, 3000 stars

Once stars have formed, strong UV radiation removes residual dust (still seen as a reflection nebula) and an open cluster is formed.



Palla & Stahler (1993, ApJ 418, 414; numbers are time in 10^6 years)

Stellar Evolution from protostar to ZAMS takes a few million years.

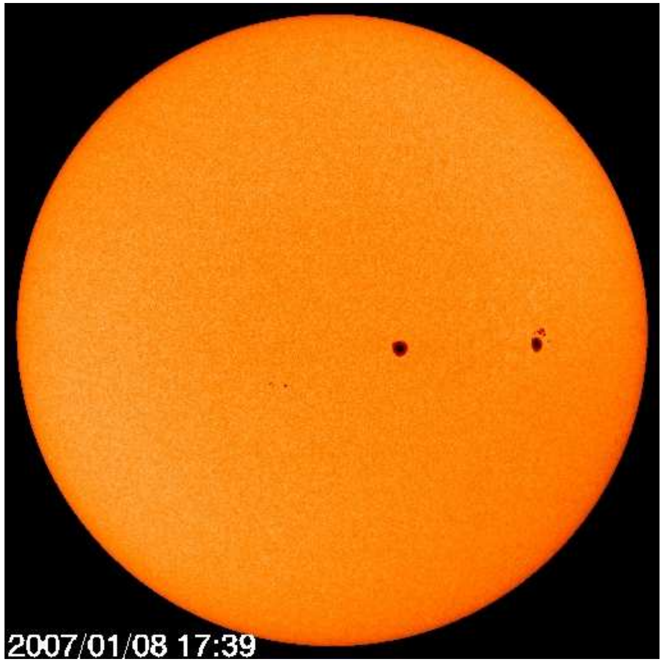
Zero Age Main Sequence

Once star has collapsed and nuclear fusion has started: zero age main sequence (ZAMS) is reached

The Main Sequence is the result of steady state fusion (“burning”) of hydrogen into helium in stellar centers.

. . . longest phase of stellar evolution (10 billion years for Sun)

Stellar structure defined by balance between pressure inwards due to gravitation and pressure outwards due to energy release (“hydrostatic equilibrium”).



The Sun: A typical star (ESA/NASA SOHO)

Stellar Structure

The structure of stars is defined by a set of four coupled differential equations which express the basic conservation and transport quantities always encountered in physics:

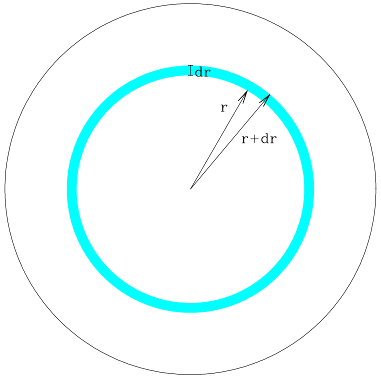
1. Mass conservation
2. Momentum conservation (=hydrostatic equilibrium)
3. Energy conservation
4. Energy transport

and quantities expressing the physical properties of material, mainly:

1. Energy generation
2. Equation of state (=dependence of density of material from physical conditions)

Mass Conservation

Density stratification of a star is defined through mass conservation:



Define M_r as the mass contained within radius r:

$$M_r = \int_0^r 4\pi r^2 \rho(r)\,\mathrm{d}r \qquad (13.4)$$

Thus the mass within a spherical shell is

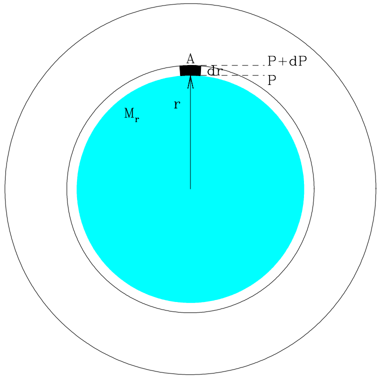
$$\mathrm{d}M_r = 4\pi r^2 \rho\,\mathrm{d}r \qquad (13.5)$$

and therefore

$$\frac{\mathrm{d}M_r}{\mathrm{d}r} = 4\pi r^2 \rho \qquad (13.6)$$

Hydrostatic Equilibrium

Pressure stratification of a star is defined through hydrostatic equilibrium:



Force on area A by slab of gas of area $\mathrm{d}A$ and density ρ:

$$\mathrm{d}F_g = -\frac{GM_r\mathrm{d}m}{r^2} = -\frac{GM_r\rho}{r^2}\mathrm{d}A\,\mathrm{d}r \qquad (13.7)$$

Bouyancy:

$$F_P = \mathrm{d}A(P(r+\Delta h) - P(r)) = \mathrm{d}A\Delta P \quad (13.8)$$

Balance of forces:

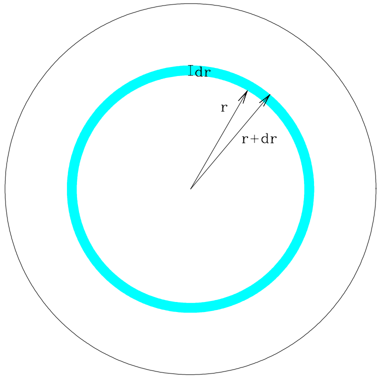
$$-\frac{GM_r\rho}{r^2}\,dA\,dr = dP\,dA \qquad (13.9)$$

such that

$$\frac{\mathrm{d}P}{\mathrm{d}r} = -\rho(r)\frac{GM(r)}{r^2} \qquad (13.10)$$

Energy Conservation

Temperature stratification of a star is defined through energy conservation:



Let ϵ = energy production coefficient, i.e., the energy released per time and unit mass.

Luminosity produced within a spherical shell is

$$\mathrm{d}L_r = \epsilon\mathrm{d}M_r = 4\pi r^2 \rho\epsilon\,\mathrm{d}r \qquad (13.11)$$

and therefore

$$\frac{\mathrm{d}L_r}{\mathrm{d}r} = 4\pi r^2 \rho\epsilon \qquad (13.12)$$

Energy Transport

Energy Conservation

Energy is transported in stars by

- radiation
- convection
- conduction

In most stars, radiation and convection are important, conduction usually not.

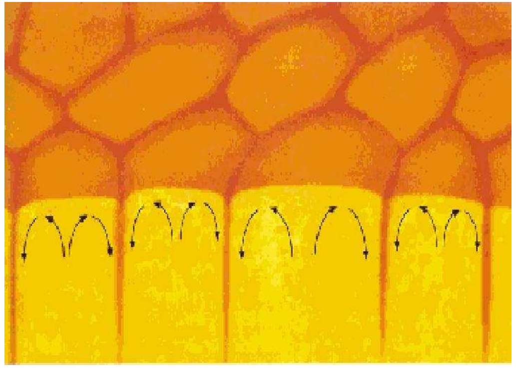
Radiative Transport = diffusive process:

- radiation produced by nuclear fusion (γ-rays)
- mean free path l_{phot} of a photon in center of the sun: few cm
- photons do random walk to the stellar surface absorbed by ions and reemitted
- number of "scatterings": $N = (R/l_{\text{phot}})^2$: e.g., $R_\odot = 700000\,\text{km}$: $N \approx 10^{20}$
- Diffusion theory:

$$\frac{\mathrm{d}T}{\mathrm{d}r} = -\frac{3}{4ac}\frac{\kappa(r)\rho(r)}{T^3}\frac{L(r)}{4\pi r^2} \qquad (13.13)$$

Energy Conservation

Convection is much more difficult to deal with, no self-consistent hydrodynamical treatment, approximate (so called mixing-length) theory



$$\frac{\mathrm{d}T}{\mathrm{d}r} = \left(1 - \frac{1}{\gamma}\right)\frac{T}{P}\frac{\mathrm{d}P}{\mathrm{d}r} \qquad (13.14)$$

Energy Conservation

Stellar structure governed by four coupled differential equations:

Mass structure
(mass conservation)

$$\frac{\mathrm{d}M}{\mathrm{d}r} = 4\pi r^2 \rho(r)$$

Pressure structure
(hydrostatic equilibrium)

$$\frac{\mathrm{d}P}{\mathrm{d}r} = -\rho(r)\frac{GM(r)}{r^2}$$

Temperature structure
(e.g. radiative transfer)

$$\frac{\mathrm{d}T}{\mathrm{d}r} = -\frac{3}{4ac}\frac{\kappa\rho(r)}{T^3}\frac{L(r)}{4\pi r^2}$$

Energy conservation

$$\frac{\mathrm{d}L}{\mathrm{d}r} = 4\pi r^2 \rho(r)\epsilon(r)$$

plus "equation of state" ($P = P(T, \rho)$), Opacities $\kappa(T, \rho, Z)$ = interaction of radiation with gas, energy generation ($\epsilon = \epsilon(T, \rho, Z)$),. . .

Stellar model: numerical solution of stellar structure equations.

Energy generation: Overview

Main sequence: Nuclear fusion of Hydrogen into Helium:

$$4\text{p} \longrightarrow {}^{4}_{2}\text{He} + E$$

How much energy is gained?

Particle physics: express mass as "rest energy equivalent" via $E = mc^2$

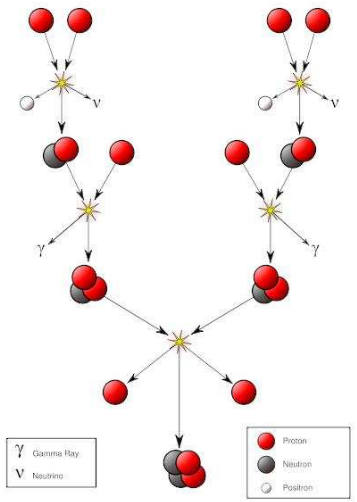
(and call it "mass". . .); usually use energy units of MeV, 1 MeV = 1.602×10^{-13} J

mass of 4 protons (4×938 MeV): 3752 MeV
− mass of ${}^{4}_{2}$He: 3727 MeV
mass defect Δmc^2: 25 MeV

In the fusion of hydrogen to helium, 0.7% of the available rest mass energy is converted to energy.

Two main burning cycles: proton-proton chain and the CNO cycle.

Energy generation: Proton-Proton chain



For moderate central temperatures, He is produced using the proton proton chain.

First, two protons create a deuteron:

$$^{1}\mathrm{H} + {}^{1}\mathrm{H} \longrightarrow {}^{2}\mathrm{H} + \mathrm{e}^{+} + \nu_{\mathrm{e}} \qquad (13.15)$$

This process is slow (happens once for a nucleon per 10^{10} years)
Then an additional proton is attached:

$$^{2}\mathrm{H} + {}^{1}\mathrm{H} \longrightarrow {}^{3}\mathrm{He} + \gamma \qquad (13.16)$$

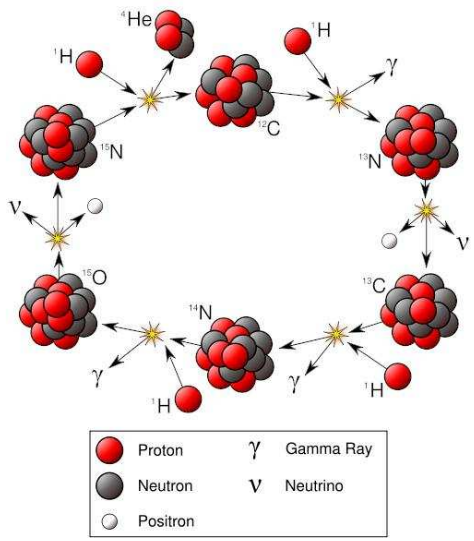
and two helium nuclei can form an alpha particle:

$$^{3}\mathrm{He} + {}^{3}\mathrm{He} \longrightarrow {}^{4}\mathrm{He} + 2\,{}^{1}\mathrm{H} \qquad (13.17)$$

This is the so called pp I-cycle, minor variations of the theme exist (pp II, pp III cycles), but pp I dominates.

pp chain dominates for $T \lesssim 2 \times 10^{7}\,\mathrm{K}$, $\epsilon_{\mathrm{pp}} \propto T^{5}$; Sun: 98.4%.

Energy generation: CNO cycle



The CNO cycle (Bethe-Weizsäcker-cycle) requires the presence of C, N, and O isotopes as catalysts.

CNO cycle has slightly smaller energy release than pp-cycle because of higher neutrino losses.

Reaction $^{14}\mathrm{N} + \mathrm{p} \longrightarrow {}^{15}\mathrm{O} + \gamma$ is th slowest reaction (one million years).

CNO cycle dominates above $2 \times 10^{7}\,\mathrm{K}$, $\epsilon_{\mathrm{CNO}} \propto T^{17}$; Sun: 1.6%.

Wikipedia

Stellar Evolution

Principle:

1. Construct stellar model by solving equations of stellar structure for given radial abundances.
2. Evaluate change in elemental abundances as a function of radius based on the local fusion processes.
3. Change abundances appropriately for a time step Δt.
4. goto step 1

We start with looking at models of the Sun in detail and then take a look at typical stellar evolution paths.

Characteristic Timescales

Main sequence: Hydrogen burning at the center.

Evolution timescale dominated by the nuclear timescale = timescale needed to use the fuel in the center of the star.

According to simulations, this is ~10% of the available Hydrogen.

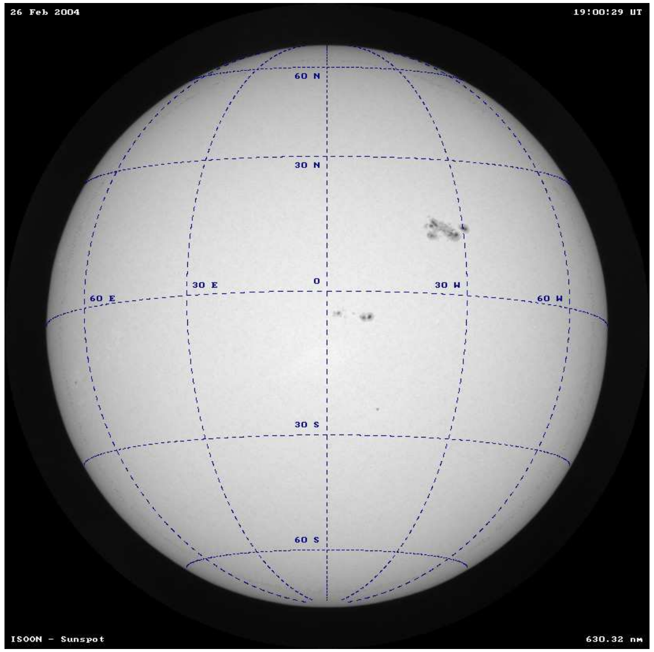
Since 0.7% of $M_{\mathrm{core}}c^{2}$ converted into He, the nuclear timescale is

$$t_{\mathrm{n}} = \frac{0.007 \cdot 0.1 M c^{2}}{L} = \frac{M/M_{\odot}}{L/L_{\odot}} \cdot 10^{10}\,\mathrm{years} \qquad (13.18)$$

A second important timescale is the timescale the star would need to radiate its stored thermal energy: thermal timescale.

Roughly given as

$$t_{\mathrm{t}} = \frac{0.5 G M^{2}/R}{L} = \frac{(M/M_{\odot})^{2}}{(R/R_{\odot})(L/L_{\odot})} \cdot 2 \times 10^{7}\,\mathrm{years} \qquad (13.19)$$



The Sun

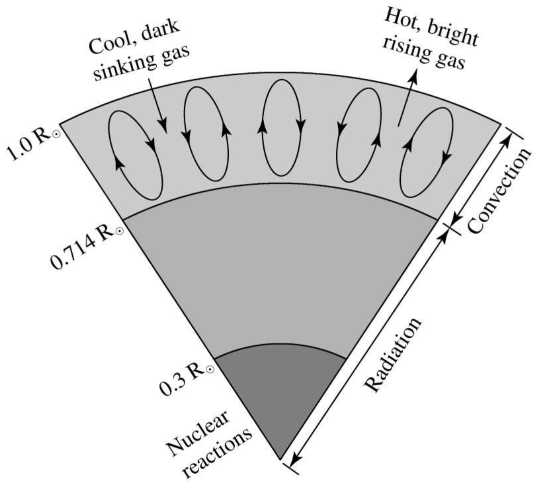
Solar Structure

Based on observations of

- Solar Mass: $1\,M_\odot = 1.997 \times 10^{30}\,\mathrm{kg} = 1.997 \times 10^{33}\,\mathrm{g}$
- Solar Luminosity: $1\,L_\odot = 3.846 \times 10^{26}\,\mathrm{W} = 3.846 \times 10^{33}\,\mathrm{erg\,s^{-1}}$
- age: $t = 4.5 \times 10^9$ yrs
- Solar chemical composition (=elemental abundances) at the surface: 75% H, 24% He, 1% metals (by mass)

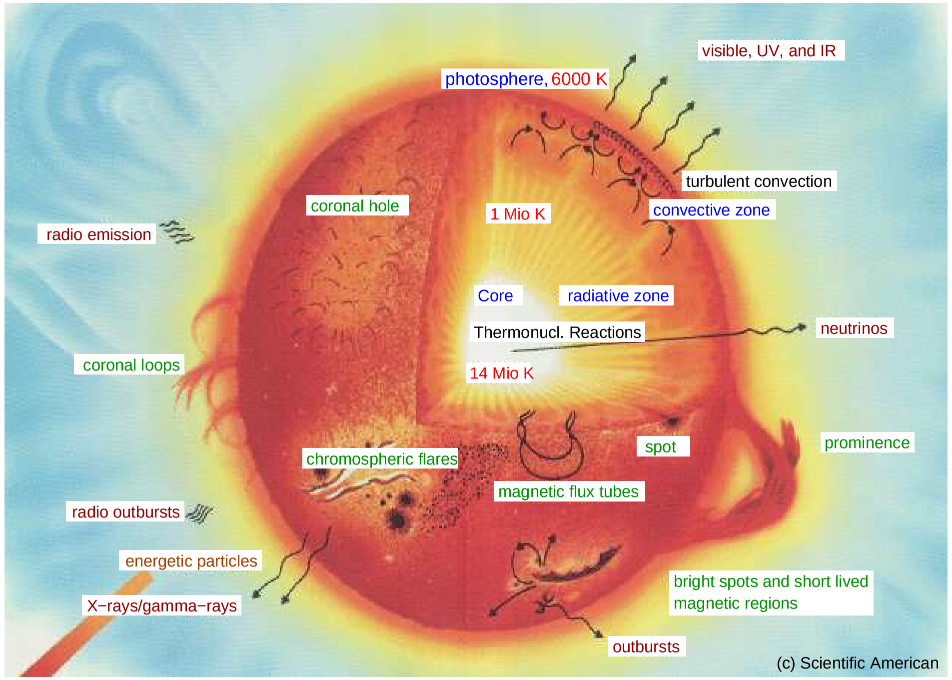
it is possible to use the equations of stellar structure to determine a model for the structure of the Sun, i.e., M_r, L_r, $\rho(r)$, $T(r)$, abundances(r) starting with a homogenous model and allow for 4.5 Gyrs of evolution.

Solar Structure



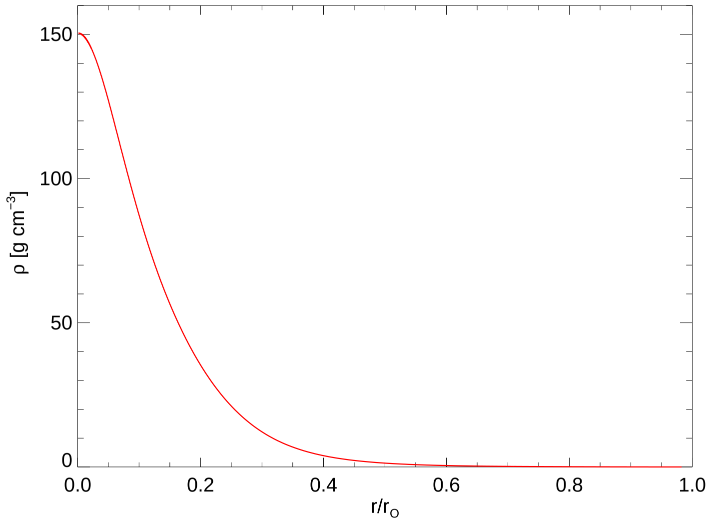
central conditions in the sun:

- $T_\mathrm{c} = 1.57 \times 10^7\,\mathrm{K}$
- $P_\mathrm{c} = 2.34 \times 10^{16}\,\mathrm{N\,m^{-2}}$
- $\rho_\mathrm{c} = 1.53 \times 10^5\,\mathrm{kg\,m^{-3}}$
- Hydrogen fraction:
  $X = 0.34$ (by mass)
- Helium fraction:
  $Y = 0.64$ (by mass)



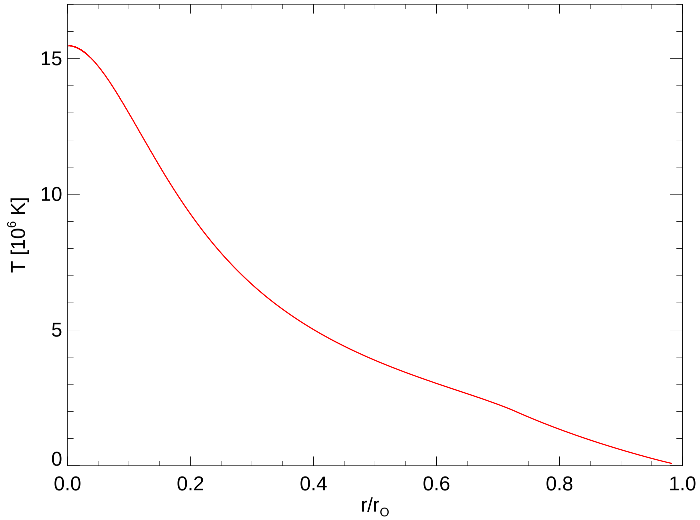
(c) Scientific American

Standard Solar Model



Standard solar model of Bahcall & Serenelli (2005, ApJ 626, 530)

Standard Solar Model



Standard solar model of Bahcall & Serenelli (2005, ApJ 626, 530)

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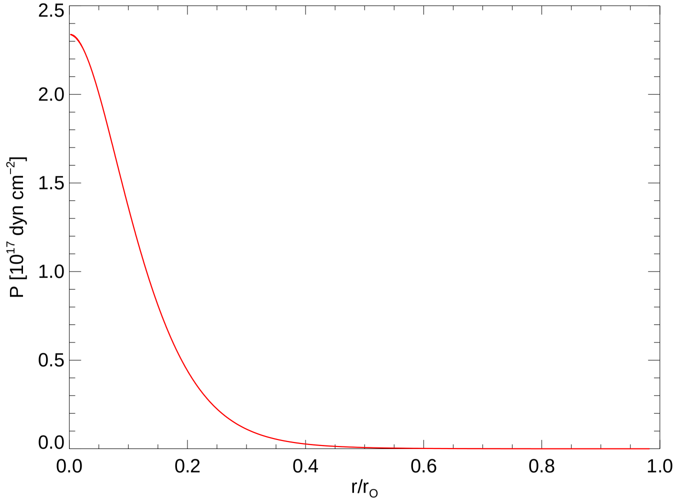
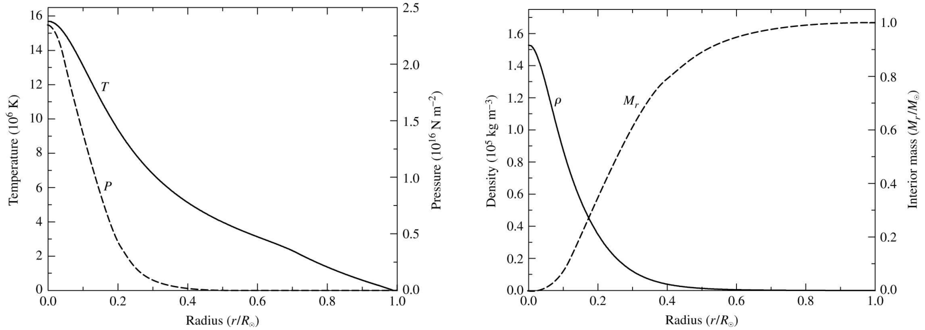


Standard solar model of Bahcall & Serenelli (2005, ApJ 626, 530; 1 dyn $=10^{-5}$ N, 1 dyn cm^{-2} $=$ 0.1 Pa)

Standard Solar Model

Standard solar model:

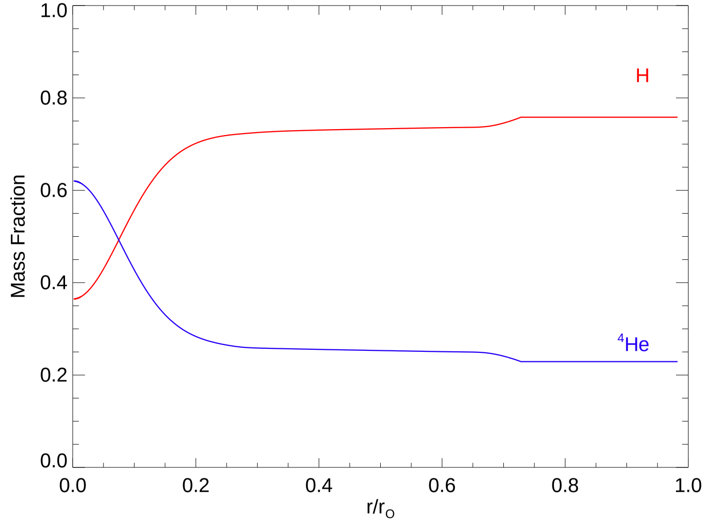


Temperature & pressure profile

Density & interior mass profile

(Carroll & Ostlie)

Standard Solar Model



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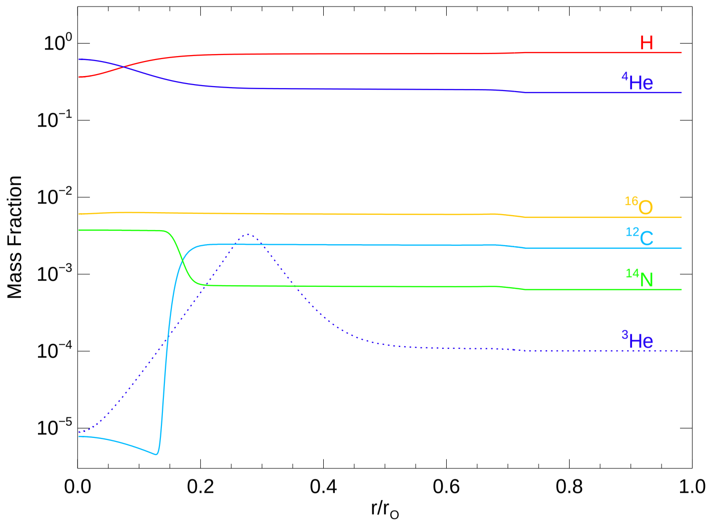
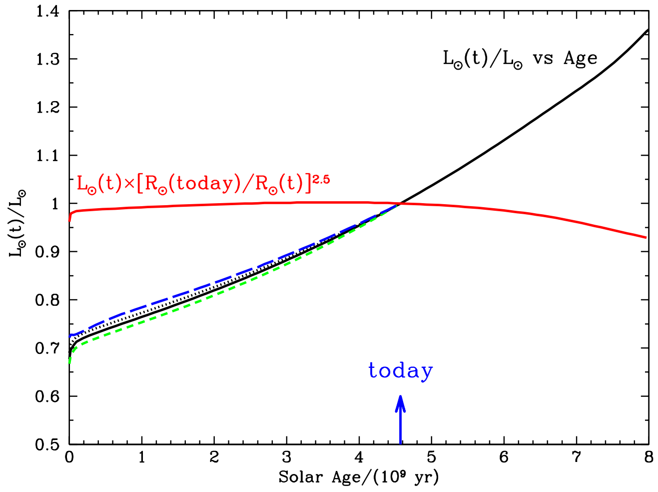
Standard Solar Model



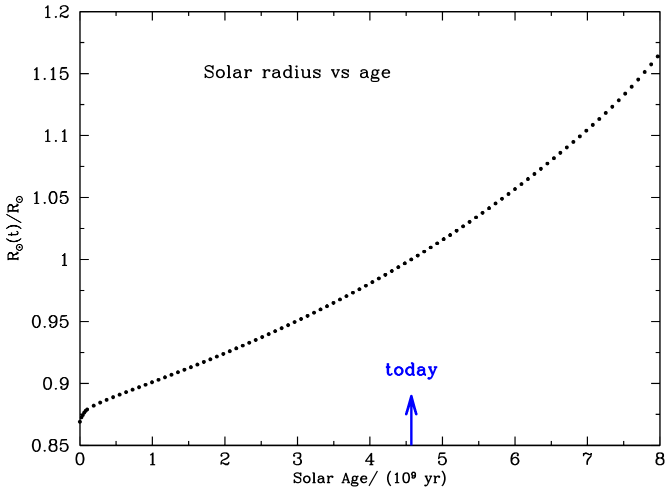
Standard solar model of Bahcall & Serenelli (2005, ApJ 626, 530)

Solar Evolution: Luminosity



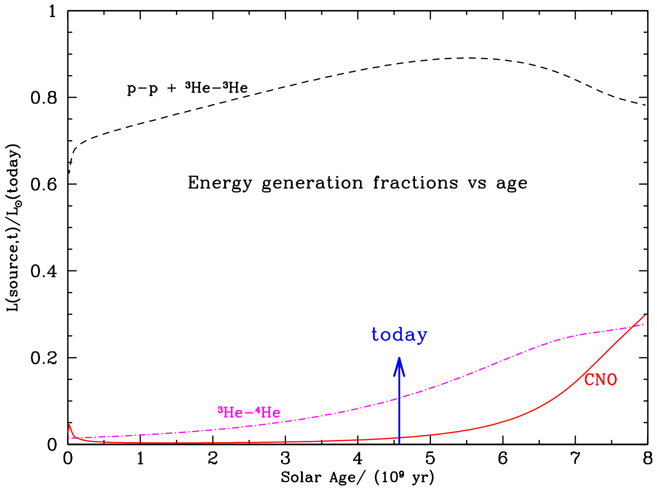
Bahcall, Pinsonneault & Basu (2001, ApJ 555, 990)

Solar Evolution: Radius



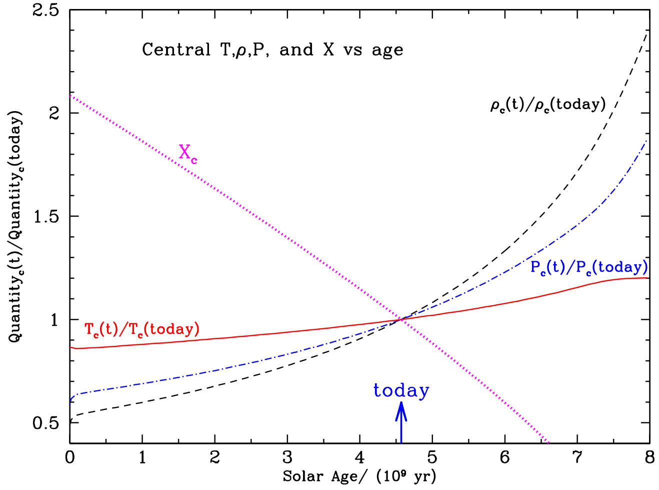
Bahcall, Pinsonneault & Basu (2001, ApJ 555, 990)

Solar Evolution: Energy Generation



Bahcall, Pinsonneault & Basu (2001, ApJ 555, 990)

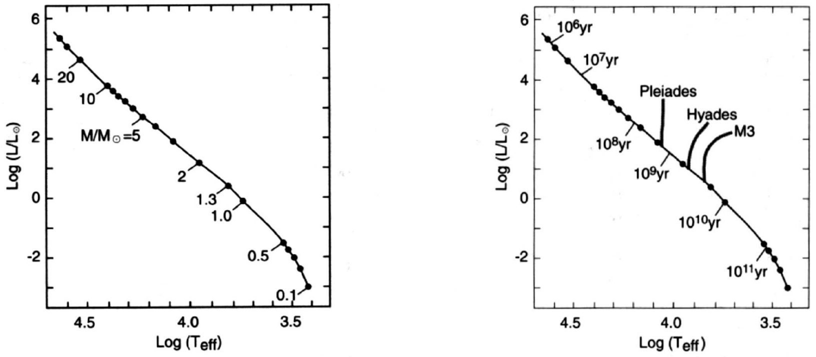
Solar Evolution: Center



Bahcall, Pinsonneault & Basu (2001, ApJ 555, 990; X_c is the central H fraction)

Zero Age Main Sequence

Main Sequence in the HRD:



Masses along the HRD

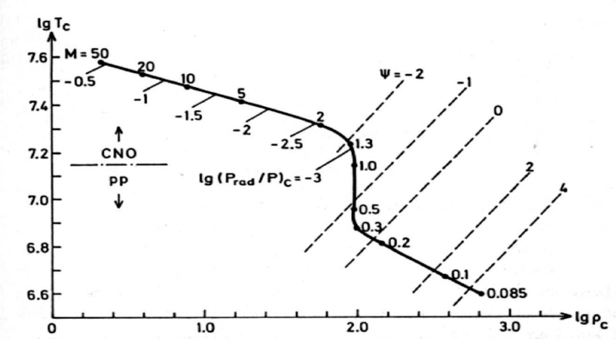
life times along the HRD

mass limits: Min.: $0.08M_\odot$ (no H-burning for lower masses)

Max.: $\approx 100M_\odot$ (radiation pressure too high, $p_\mathrm{rad} \sim T_\mathrm{eff}^4$)

Zero Age Main Sequence

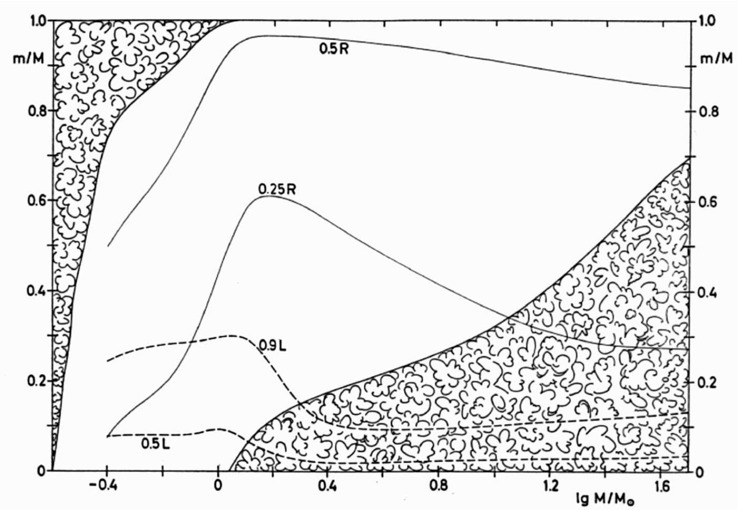
Properties of Main Sequence stars:



Central temperatures & densities

- strong change at $\approx 1M_\odot$
- low masses ($< 0.5M_\odot$): low temperature, very high density
- high masses ($> 1.3M_\odot$): high temperature, low density
- low mass stars burn H in the pp-chains
- high mass stars burn H in the CNO-cycle

Zero Age Main Sequence



Kippenhahn-diagram: Internal Structure of Main Sequence stars

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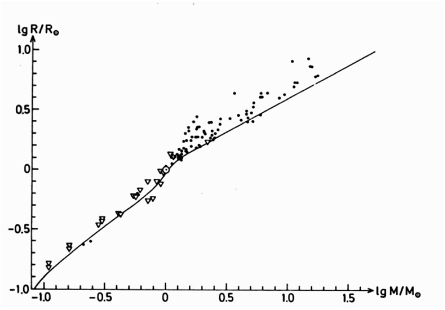
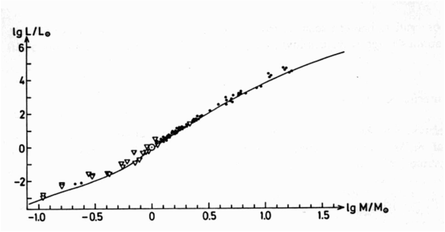
Zero Age Main Sequence

Empirical test of main sequence models:

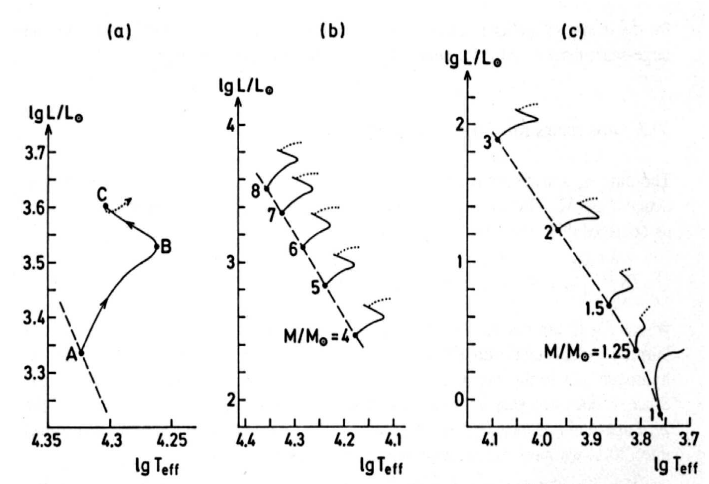


Mass–radius relation



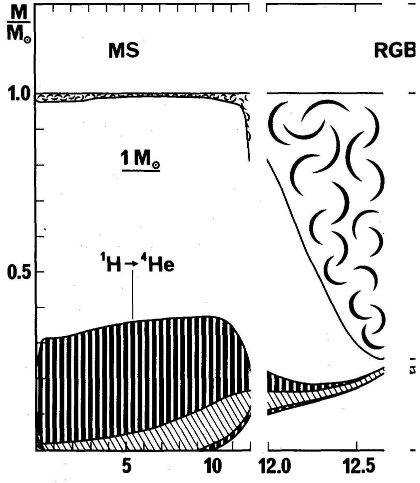
mass–luminosity relation

Main Sequence Evolution



main sequence evolution from zero age to helium exhaustion

Post Main Sequence

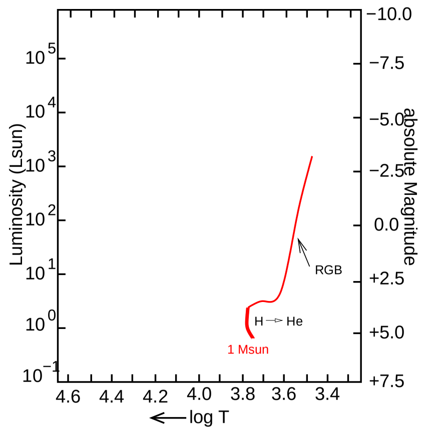


(Maeder & Meynet, 1989)

Once H is exhausted in center: H continues to burn in a shell around the He core ("shell burning").

For stars with $M \lesssim 1\,M_\odot$: Star reacts by expanding convective hull until it is almost fully convective.

Post Main Sequence



(after Iben, 1991)

Once H is exhausted in center: H continues to burn in a shell around the He core ("shell burning").

For stars with $M \lesssim 1\,M_\odot$: Star reacts by expanding convective hull until it is almost fully convective.

$\Longrightarrow$ luminosity increases, temperature decreases

$\Longrightarrow$ motion in HRD horizontally towards the right, then upwards to higher L: red giant stage.

Post Main Sequence

Reminder: stars are in hydrostatic equilibrium: inwards gravitational pressure balanced by outwards gas pressure

Since the gas pressure is $P = nkT$: energy source needed to heat gas (=fusion).

This is a problem for the core during the red giant stage, as virtually no fusion ongoing

$\Longrightarrow$ Core gets compressed

$\Longrightarrow$ ρ and T increase

BUT:

collapse cannot continue indefinitely!

$\Longrightarrow$ once ρ has increased appreciably, there must be a point where quantum mechanical effects become important.

QM interlude

Quantum mechanics: The Pauli exclusion principle:

For particles such as electrons ("Fermions"), at least one of their quantum numbers must be different.

Quantum numbers are, e.g.,

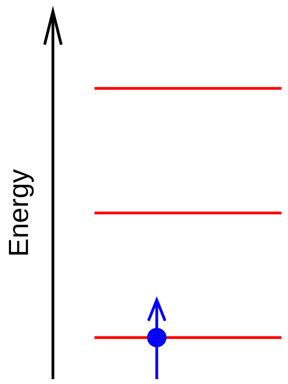
- position (x, y, z),
- momentum p ($p_x = mv_x$, $p_y = mv_y$, $p_z = mv_z$),
- angular momentum,
- spin (s)

All of these numbers are "quantized", i.e., can only have discrete values (e.g., spin: $+1/2$, $-1/2$).

Heisenberg's principle: (6-d) phase space is quantized: $\Delta p \times \Delta V \approx h^3$
Each cell in phase space can host two electrons of different spin

In a typical gas, this is not a problem ("phase space is (almost) empty") once it becomes dense $\Longrightarrow$ exclusion principle kicks in.

QM interlude

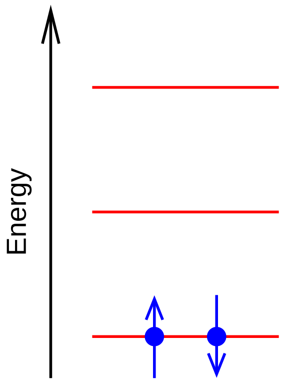


Effect of high density on electron energy

Energy of electrons at the same position in space

QM interlude

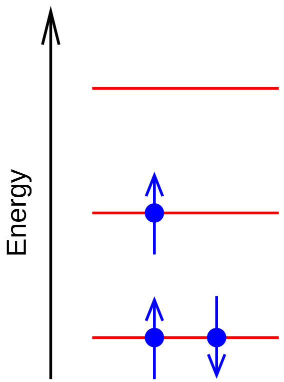
Effect of high density on electron energy



Energy of electrons at the same position in space

QM interlude

Effect of high density on electron energy



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QM interlude

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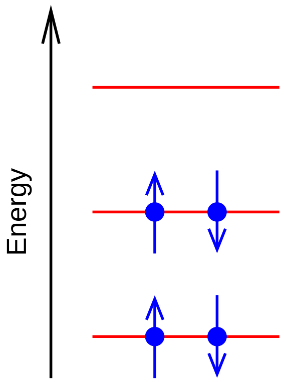

Energy of electrons at the same position in space

QM interlude

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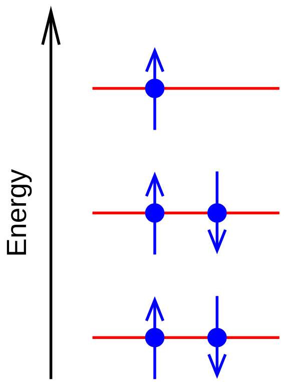
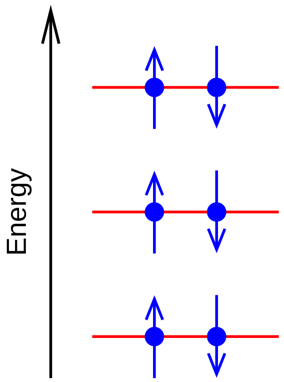


Energy of electrons at the same position in space

QM interlude

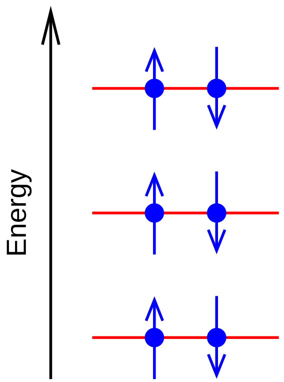


Energy of electrons at the same position in space

Effect of high density on electron energy:

In degenerate electron gases, electrons have much higher energies than in thermal gas.

QM interlude



Energy of electrons at the same position in space

Effect of high density on electron energy:

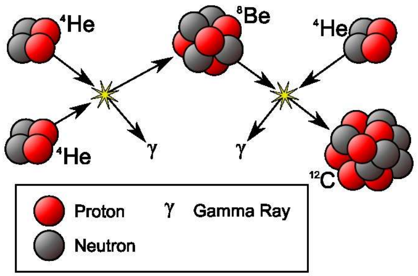
In degenerate electron gases, electrons have much higher energies than in thermal gas.

Interaction of electrons results in degeneracy pressure:

$$P = \frac{\hbar^2}{m_e} n_e^{5/3} \propto \rho^{5/3}$$

Note: The degeneracy pressure is independent of the temperature!

Post Main Sequence



In the degenerate core, once $T_{\text{core}} \sim 100 \times 10^6$ K: Triple alpha process starts:

$$^4\text{He} + {^4\text{He}} \leftrightarrow {^8\text{Be}}$$
$$^8\text{Be} + {^4\text{He}} \rightarrow {^{12}\text{C}}$$

Since ^{8}Be has a half life of only 2.6×10^{-16} s: this can only work effectively if 3 α-particles collide.

But core is degenerate:

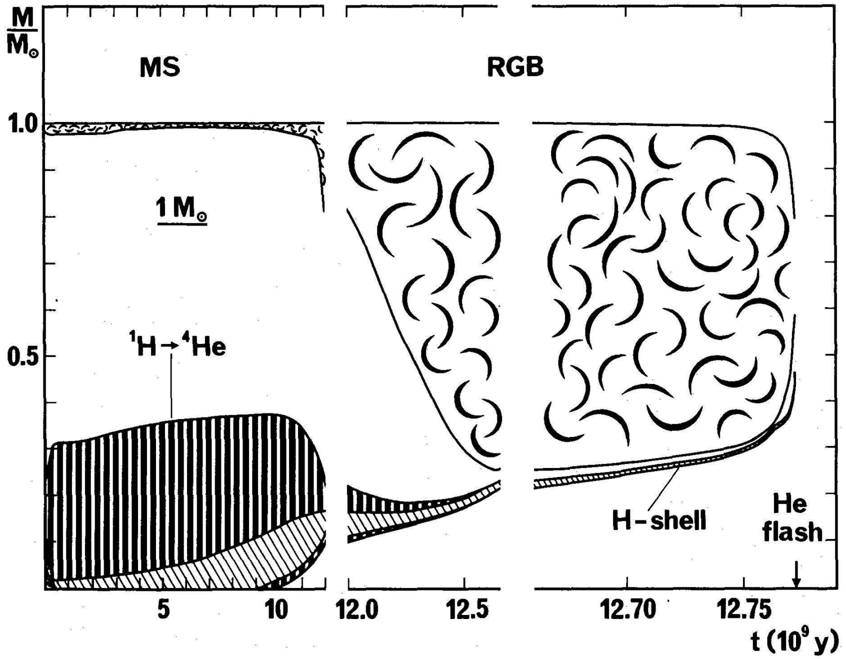
$\Longrightarrow$ High thermal conductivity of electrons

$\Longrightarrow$ core has uniform temperature

$\Longrightarrow$ 3α onset is rapid

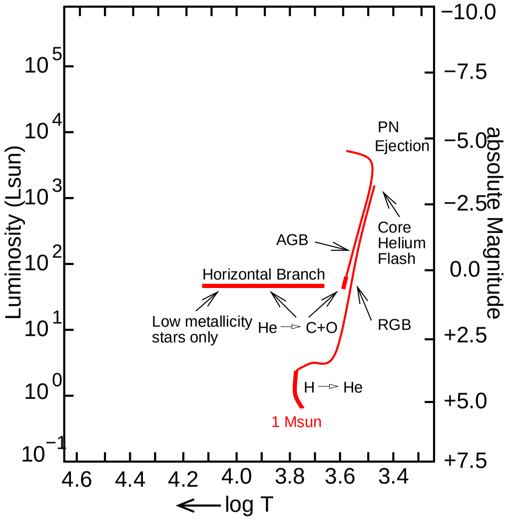
$\Longrightarrow$ He flash

Not seen on surface (“buffered” by convective envelope).



Evolution of the structure of a 1 $M_\odot$ star to the Helium flash (Maeder & Meynet, 1989).

Post Main Sequence



After the He flash star has He burning in core and H shell burning

⟹ starts to expand again

⟹ “asymptotic giant branch”

Instable He fusion processes (“thermal pulses”) lead to ejection of outer layers (∼50% of total mass!)

Effect of He core being unable to transport energy away quickly enough.

⟹ inner (hotter) parts of star become visible.

after Iben, 1991

Ring Nebula (HST/STScI/NASA)

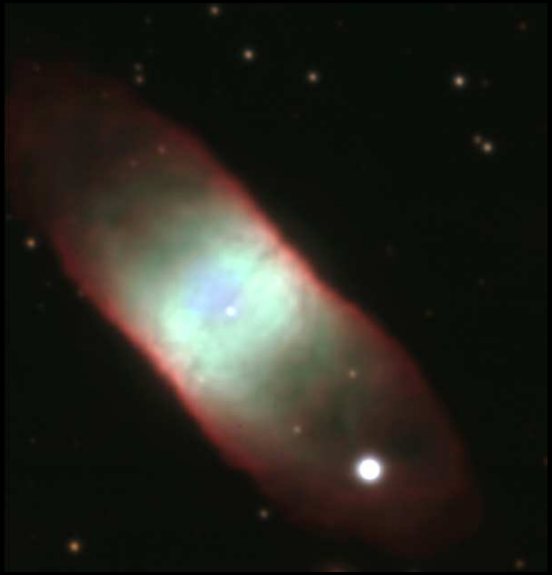
planetary nebulae: material ejected during AGB phase, photoionized once remaining core of former star has shed enough mass to emit UV photons.

Abell 39 (WIYN, AURA, NOAO, NSF)

planetary nebulae: material ejected during AGB phase, photoionized once remaining core of former star has shed enough mass to emit UV photons.

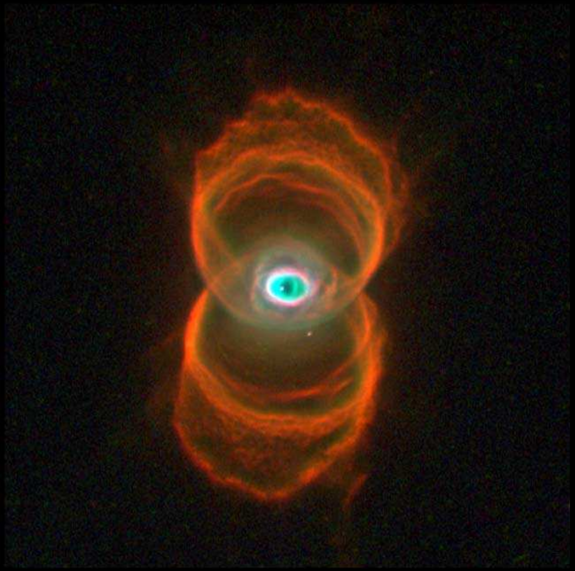
NGC 6853/M27 (“Dumbbell Nebula”; ESO VLT/FORS)

planetary nebulae: material ejected during AGB phase, photoionized once remaining core of former star has shed enough mass to emit UV photons.



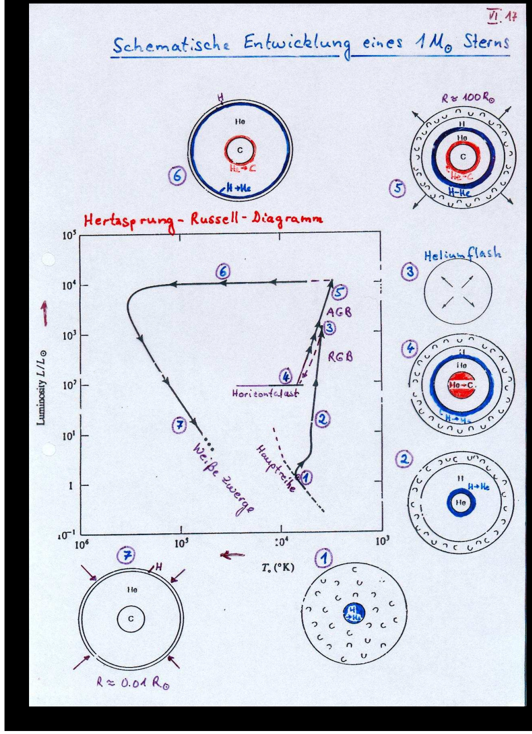
IC4406 (ESO VLT)

planetary nebulae: material ejected during AGB phase, photoionized once remaining core of former star has shed enough mass to emit UV photons.



Hourglass Nebula (HST/Sahai/Trauger)

planetary nebulae: material ejected during AGB phase, photoionized once remaining core of former star has shed enough mass to emit UV photons.



Summary: Evolution of Low Mass Stars:

1. =

2.
   =

3.
   = helium ignition in degenerate electron gas

4.
   =

5.
   = Tip of AGB: envelope ejection through dust formation & pulsations

6. hot star excites the ejected envelope to shine

7. =

Reminder: Main Sequence

Structure on the Main Sequence: Simulations show existence of two regimes:

lower main sequence : stars have structure similar to Sun:

- energy generation: pp-chain ($\epsilon \propto T^5$)
- inner radiative core
- convective hull

upper main sequence : for central temperatures of 18×10^6 K (1.5 $M_\odot$ stars): pp-chain and CNO-cycle produce equal amounts of energy. Above that: CNO dominates.

- energy generation: CNO-cycle ($\epsilon \propto T^{17}$)
- inner convective core since energy generation from CNO cycle strongly peaked towards center.
- outer radiative hull

Stars born with masses $> 8\,M_\odot$

Evolution on MS similar, however, faster than for low mass stars.

More massive stars reach threshold temperature for 3α and subsequent nuclear burning before reaching degeneracy

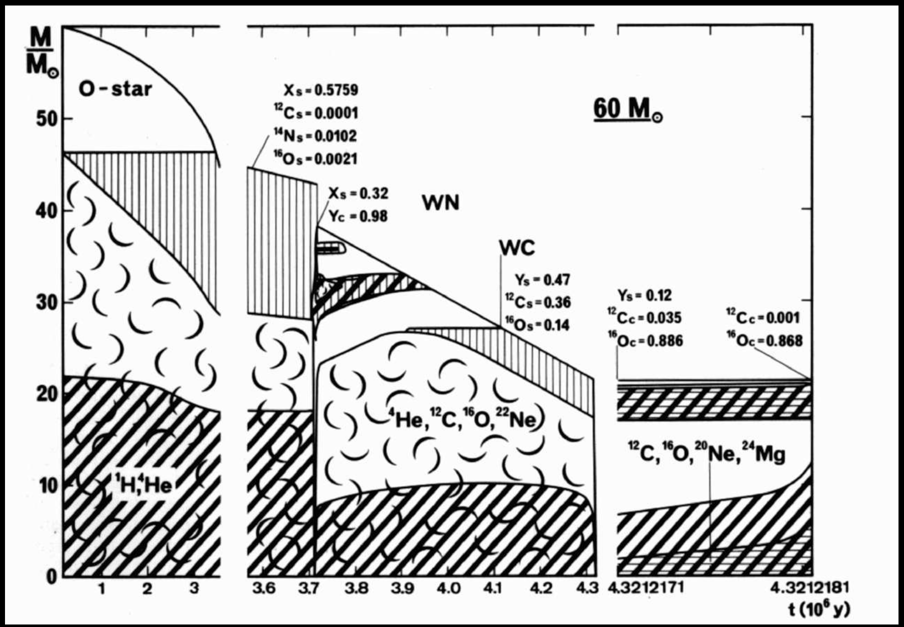
$\Longrightarrow$ He just starts to burn.

In these objects, higher order fusion processes can kick in (but are energetically unimportant): alpha reactions

$$^{12}\mathrm{C} + {}^{4}\mathrm{He} \rightarrow {}^{16}\mathrm{O} + \gamma$$
$$^{16}\mathrm{O} + {}^{4}\mathrm{He} \rightarrow {}^{20}\mathrm{Ne} + \gamma$$
$$^{20}\mathrm{Ne} + {}^{4}\mathrm{He} \rightarrow {}^{24}\mathrm{Mg} + \gamma$$

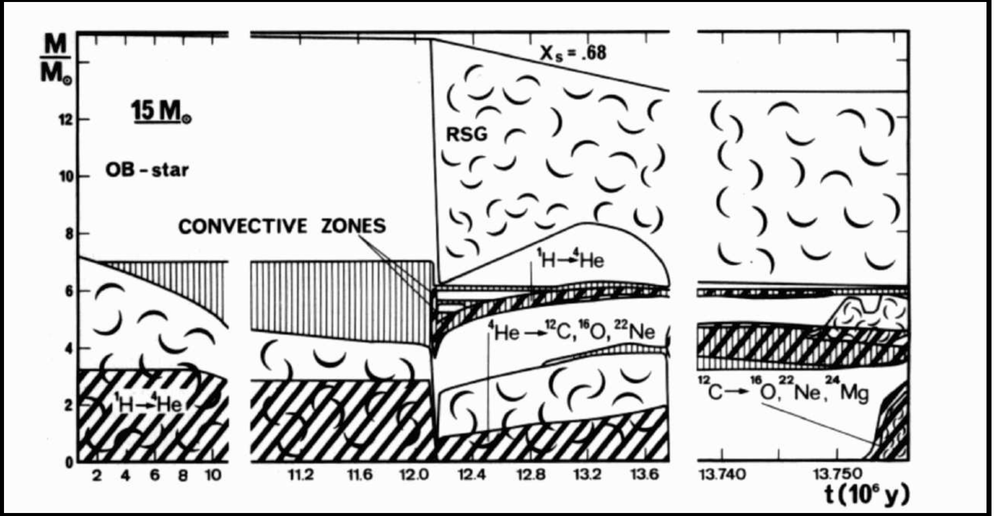
Outer layers continue H shell burning.

During evolution of star on red giant branch: convective hull moves deeper into core, can mix fusion products into outer layers.

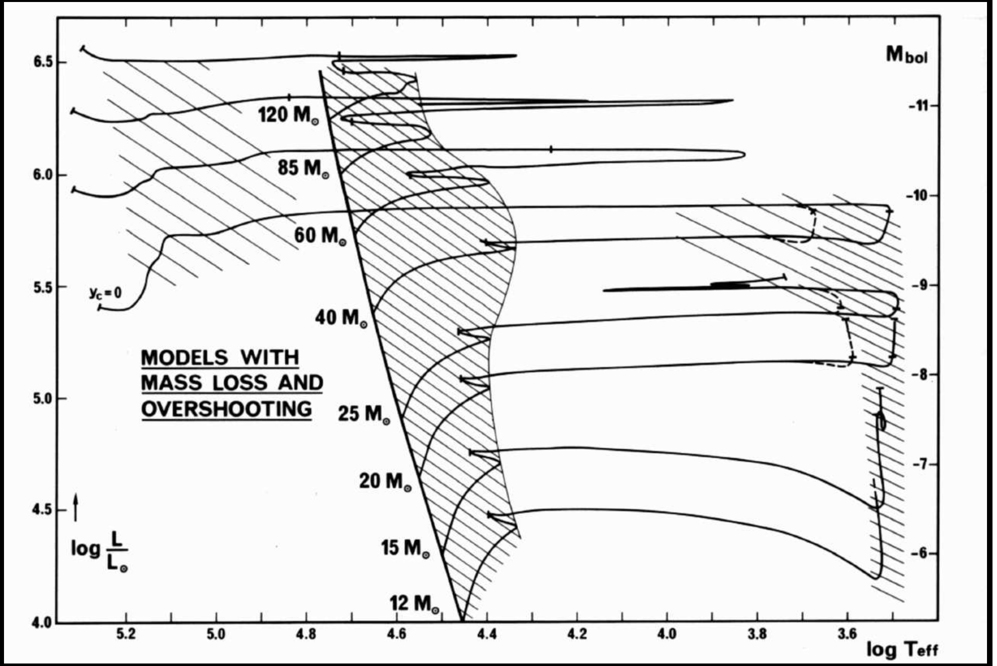


Evolution of the star.

Note the very strong mass loss!



Evolution of the star.



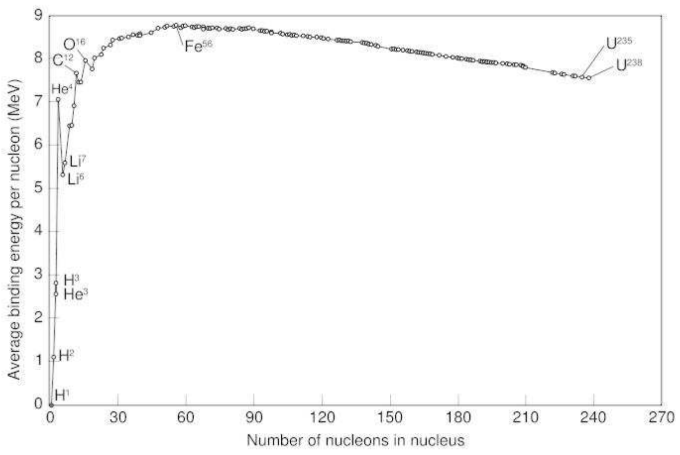
Summary: Evolution of in the HRD.

Stars born with masses $> 8\,M_\odot$

Nuclear reactions in massive ($> 8\,M_\odot$) stars:

Reaction	above T $[10^6\,\mathrm{K}]$	Energy gain [MeV]
Hydrogen burning		
$4\,{}^1\mathrm{H} \longrightarrow {}^4\mathrm{He}$	4	6.55
Helium burning		
$3\,{}^4\mathrm{He} \longrightarrow {}^8\mathrm{Be} + {}^4\mathrm{He} \longrightarrow {}^{12}\mathrm{C}$	100	<0.61
Carbon burning		
${}^{12}\mathrm{C} + {}^4\mathrm{He} \longrightarrow {}^{16}\mathrm{O}$	600	<0.54
$2\,{}^{12}\mathrm{C} \longrightarrow {}^4\mathrm{He} + {}^{20}\mathrm{Ne}$		
${}^{20}\mathrm{Ne} + {}^4\,\mathrm{He} \longrightarrow \mathrm{n} + {}^{23}\mathrm{Mg}$		
Oxygen burning		
$2\,{}^{16}\mathrm{O} \longrightarrow {}^4\mathrm{He} + {}^{28}\mathrm{Si}$	1000	<0.3
$2\,{}^{16}\mathrm{O} \longrightarrow 2\,{}^4\mathrm{He} + {}^{24}\mathrm{Mg}$		
Silicon burning		
$2\,{}^{28}\mathrm{Si} \longrightarrow {}^{56}\mathrm{Fe}$	3000	< 0.18

Stars born with masses $> 8\,M_\odot$



^{56}Fe is one of the most tightly bound nucleons $\Longrightarrow$ Star has a problem once ^{56}Fe reached: fusion processed become endotherm

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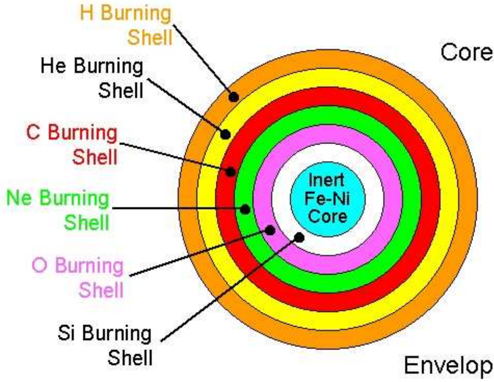
Stars born with masses $> 8\,M_\odot$



Core Radius: ~1 R_{earth}

Envelope Radius: ~ 5 AU

successive stages of nuclear burning: final state: onion-shell model

1. H burning, ash: He
2. He burning, ashes: C, O, Ne, Mg
3. C burning, ashes: Ne, Na, Mg
4. Ne burning, ashes: O, Mg . . .
5. O burning, ashes: Si, P, S, . . .
6. Si burning, ashes: Fe, Ni