Lecture 10

Nucleosynthesis During Helium Burning and the s-Process

Woosley (2019) helium stars evolved at constant mass

Mass	Density	Temperature/10^8
3.5	1430	1.75
4.0	1260	1.79
4.5	1140	1.82
6.0	910	1.89
7.0	814	1.92

Z = 0.02. Helium burning at central helium mass fraction of 50%.

At 1.9 x 10^8 K, the temperature sensitivity of the 3α rate is approximately T^{20}.

A. Thermodynamic Conditions (Massive Stars)

Review: The helium core evolves like a separate star, mass M_α
τ_{He} ~ 10^6 years (+- factor of three); inside RSG, BSG, or as WR star
n= 1.5 to 3; assume 3

$$M_\alpha = 18\,\frac{\sqrt{1-\beta}}{\mu^2\beta^4} \qquad \mu = \frac{4}{3}\ \text{(for pure helium)} \qquad \text{solve for } \beta$$

$$T_C = 4.6\times10^6\,\mu\beta\left(\frac{M_\alpha}{M_\odot}\right)^{2/3}\rho_c^{\,1/3}\quad \text{K}$$

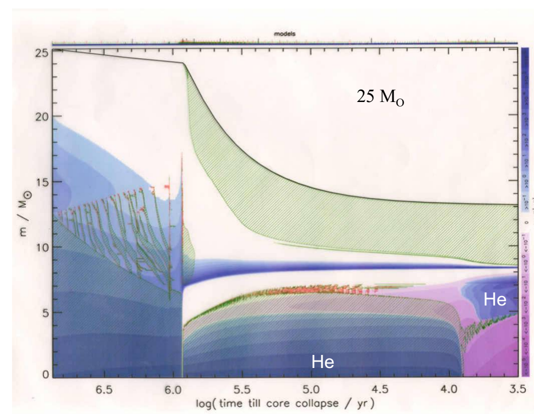
E.g., $M_\alpha = 6$ (a 20 $M_\odot$ main sequence star)

$\beta = 0.87$

$$T_c = 1.8\times10^8\left(\frac{\rho_c}{1000\ \text{g cm}^{-3}}\right)^{1/3}\ \text{K}$$

need temperatures > 10^8 to provide significant energy generation by 3α so this really just sets the density.

So, typical temperatures are 1.5 - 2 x 10^8 K (higher in shell burning later). As the core evolves, the temperature and density go up significantly. Note non-degenerate.

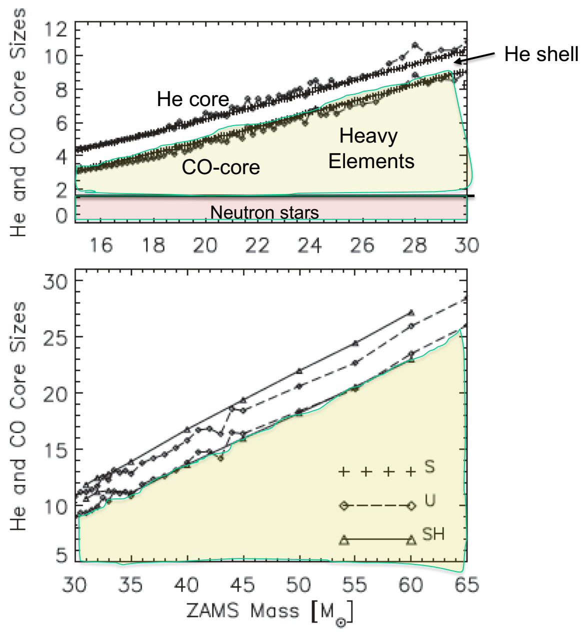


Most mass loss occurs during helium core burning for single stars

The helium core itself can grow (single stars) or lose mass (close binary plus winds)

Helium burns in the star's center and later in a shell. Helium burning in the shell is incomplete and makes more carbon relative to oxygen.

During helium burning the star is a red or blue supergiant depending on semiconvection, or a WR star if it loses its envelope. See previous lecture.



Sukhbold and Woosley (2014)

Complications:

- If the helium core grows just a little bit towards the end of helium burning, the extra helium convected in greatly decreases the ^{12}C synthesis.
- Mass loss from very massive WR stars can greatly increase the synthesis of both ^{12}C and ^{16}O for stars over 35 solar masses
- The uncertain rate for $^{12}C(\alpha,\gamma)^{16}O$
- $^{12}C/^{16}O$ ratio may be affected by post-helium burning evolution and by black hole formation above some critical main sequence mass. ^{16}O is made in the more massive (massive) stars.

Massive stars are definitely the site of oxygen nucleosynthesis and contribute appreciably to carbon.

B. Major Nucleosynthesis – C and O

In massive stars after helium burning in the stars center (calculations included semiconvection). Woosley et al (2007)

In the sun, $^{12}C/^{16}O = 0.32$

$M/M_\odot$	$X(^{12}C)$	$X(^{16}O)$	$X(^{20}Ne)$
12	0.229	0.752	1.48(-3)
15	0.214	0.765	2.05(-3)
19	0.202	0.776	3.15(-3)
25	0.184	0.792	5.62(-3)
35	0. 169	0.801	1.03(-2)

Buchman $^{12}C(\alpha,\gamma)^{16}O$ multiplied by 1.2.

If the star contains appreciable metals there is, as we shall see also ^{22}Ne and ^{18}O.

C. TRACE ELEMENT NUCLEOSYNTHESIS AND NEUTRONIZATION

Prior to the ignition of the 3α reaction the star goes through a brief stage of "nitrogen burning". All the initial CNO in the core has been converted to ^{14}N which captures an α-particle and experiences a weak decay

$$\boxed{^{14}N(\alpha,\gamma)^{18}F(e^+\nu)^{18}O}$$

The *number* of CNO nuclei is preserved

$$X(^{14}N) = 14.\cdot Y(^{14}N) = 14.\cdot\left(Y_i(^{14}N) + Y_i(^{12}C) + Y_i(^{16}O)\right)$$
$$= 0.0095\left(Z/Z_\odot\right)$$

This reaction is very important both as a source of ^{18}O in nature and because it creates a neutron excess

$$\eta = 1\text{-}2Y_e = 1 - 2\sum Z_i Y_i = \sum (N_i + Z_i) Y_i - 2\sum Z_i Y_i$$
$$\boxed{\eta = \sum (N_i - Z_i) Y_i} \qquad -1 < \eta < 1$$

Prior to this reation $Y_e = 0.50$ because both ^4He and ^{14}N have equal numbers of neutrons and protons and η=0. After this reaction $\eta > 0$ with a value that depends on the initial metallicity (i.e., initial mass fractions of C, N, and O) of the star.

Before ${}^{14}N(\alpha,\gamma)^{18}F(e^{+}\nu)^{18}O$ the composition was almost entirely ^{4}He and ^{14}N, hence $\eta \approx 0$ (actually a small posive value exists because of ^{56}Fe and the like).

After this reaction

$$\boxed{\eta = 0.00136\,\frac{Z}{Z_{\odot}}}$$

During helium core burning, ^{18}O is later mostly destroyed by ${}^{18}\mathrm{O}(\alpha,\gamma)^{22}$Ne, but the neutron excess is unchanged. η can only be changed by weak interactions.

This neutron excess is very important to the subsequent nucleosynthesis and one of the main reasons nucleosynthesis in low Z stars is different!

During helium shell burning, which does not go to completion, ^{18}O remains undestroyed and this is the source of ^{18}O in nature. Convection helps to preserve it

Lifetimes in years $X_{\alpha} = 0.5$, $\rho = 1000$ g cm^{-3}

	$T_8 = 1.8$	$T_8 = 2.0$	
${}^{14}\mathrm{N}(\alpha,\gamma)$	15	1.1	*destroyed*
${}^{18}\mathrm{O}(\alpha,\gamma)$	2700	98	*destroyed in core*
${}^{22}\mathrm{Ne}(\alpha,\gamma) + (\alpha,\mathrm{n})$	4.1(6)	2.2(5)	*partly burned*
${}^{16}\mathrm{O}(\alpha,\gamma)$	3.5(10)	2.3(9)	*survives*

Note that the reaction rate for ${}^{22}\mathrm{Ne}(\alpha,\mathrm{n})^{25}$Mg is uncertain and this is an important source of error in s-process nucleosynthesis. Need all the ^{22}Ne to burn

So one expects that, depending on mass and rates, some but not all of the ^{22}Ne will burn when the temperature goes up towards the end of helium burning.

The following table gives the temperature at the center of the given model and the mass fractions of ^{22}Ne, ^{25}Mg, and ^{26}Mg each multiplied by 1000, when the helium mass fraction is 1% and zero

Woosley, and Heger (2007)

M	T_C	^{22}Ne	^{25}Mg	^{26}Mg
12	2.42	13.4	0.51	0.61
		12.3	1.17	1.05
15	2.54	12.7	0.98	0.91
		11.1	1.99	1.80
19	2.64	11.5	1.73	1.54
		9.4	2.90	2.87
25	2.75	9.8	2.67	2.59
		6.96	4.05	4.54
35	2.86	7.37	3.87	4.22
		4.41	5.18	6.39

The remainder of the ^{22}Ne will burn early during carbon burning, but then there will be more abundant "neutron poisons", Na and Mg.

These numbers are quite sensitive to the uncertain reaction rate for ${}^{22}\mathrm{Ne}(\alpha,n)^{25}$Mg and may be lower limits to the ^{22}Ne consumption.

D. The s-Process in Massive Stars

Late during helium burning, when the temperature rises to about 3.0 x 10^{8} K, ^{22}Ne is burned chiefly by the reaction ${}^{22}\mathrm{Ne}(\alpha,\mathrm{n})^{25}$Mg (with some competition from ${}^{22}\mathrm{Ne}(\alpha,\gamma)^{26}$Mg).

Where do the neutrons go?

Some go on ^{56}Fe but that fraction is small:

$$f = \frac{\sigma_{56} Y_{56}}{\sum \sigma_i Y_i}$$

^{16}O $\quad Y_{16} \approx \dfrac{0.5}{16} = 3.1\times 10^{-2} \quad Y_{16}\,\sigma_{16} \approx 1.2\times 10^{-3}$

But, ${}^{17}\mathrm{O}(\alpha,\mathrm{n})^{20}$Ne destroys the ^{17}O and restores the neutron

^{22}Ne $\quad Y_{22} \approx \dfrac{0.005}{22} = 2.3\times 10^{-4} \quad Y_{22}\,\sigma_{22} \approx 1.3\times 10^{-5}$

^{25}Mg $\quad Y_{25} \approx \dfrac{0.005}{25} = 2\times 10^{-4} \quad Y_{25}\,\sigma_{25} \approx 1.3\times 10^{-3}$

For σ in mb

^{56}Fe $\quad Y_{56} \approx \dfrac{0.0013}{56} = 2.3\times 10^{-5} \quad Y_{56}\,\sigma_{56} \approx 2.7\times 10^{-4}$

30 keV neutron capture cross sections

(mostly from Bao et al, ADNDT, 2000)

Nucleus	σ (mb)	Nucleus	σ (mb)
^{12}C	0.0154	^{54}Fe	27.6
*^{16}O	0.038	^{56}Fe	11.7
^{20}Ne	0.119	^{57}Fe	40.0
^{22}Ne	0.059	^{58}Fe	12.1
^{24}Mg	3.3	^{58}Ni	41.0
^{25}Mg	6.4	^{64}Zn	59
^{26}Mg	0.126	^{65}Zn	162
^{28}Si	2.9	^{66}Zn	35
		^{88}Sr	6.2 (closed shell)

The large cross section of ^{25}Mg is particularly significant since it is made by ^{22}Ne(α,n) ^{25}Mg.

* Igashira et al, *ApJL*, **441**, L89, (1995); factor of 200 upwards revision

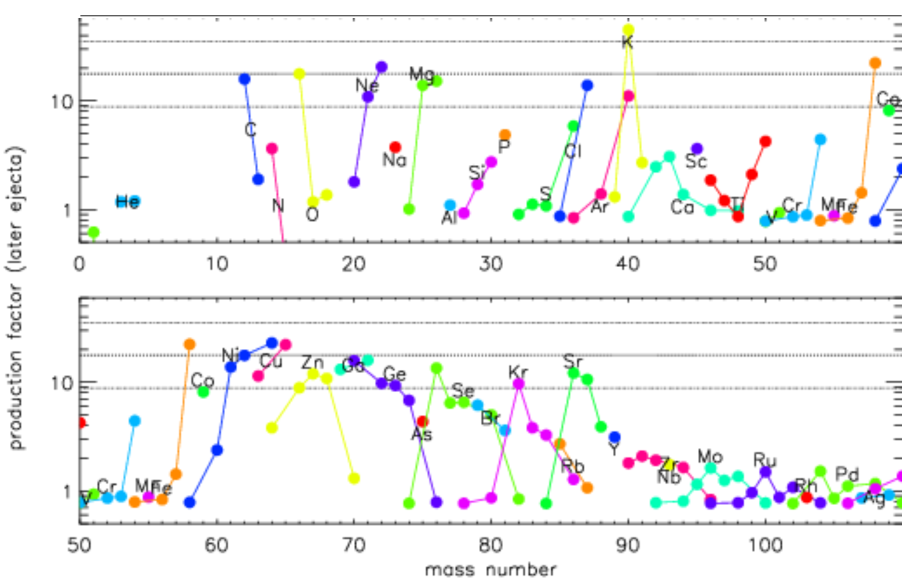
So a fraction $\dfrac{Y_{56}\,\sigma_{56}}{Y_{25}\,\sigma_{25}+...}$ ~$10-20\%$ capture on iron. How many neutrons is this? If all the neutrons captured on iron one would have

$$Y_n / Y_{Fe} \sim Y_{22} / Y_{56} \approx 30 \text{ neutrons per iron}$$

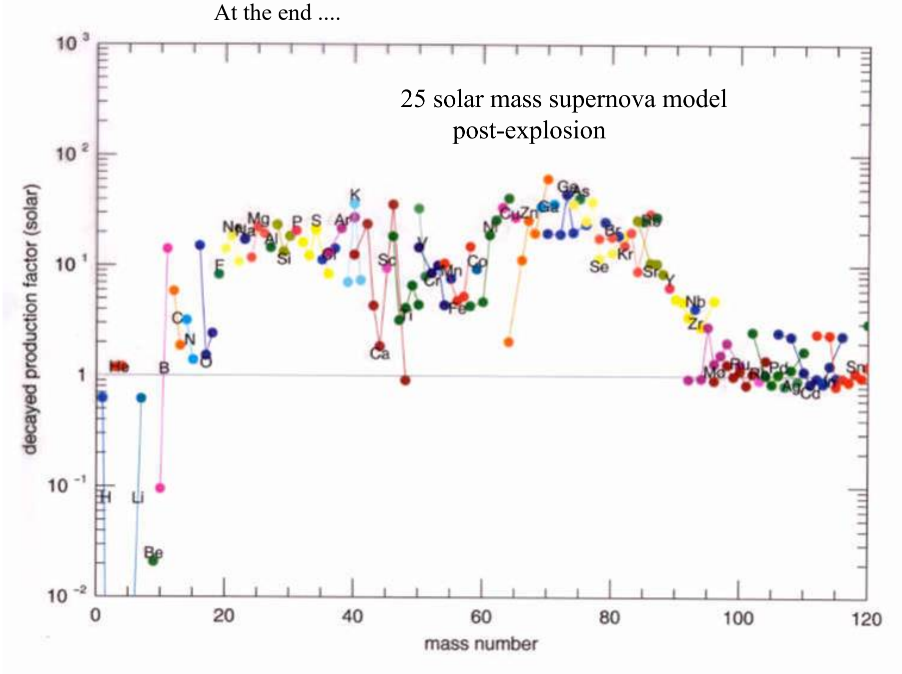
where we have assumed a mass fraction of 0.015 for ^{22}Ne and 0.0012 for ^{56}Fe and that all ^{22}Ne burns by (α,n).

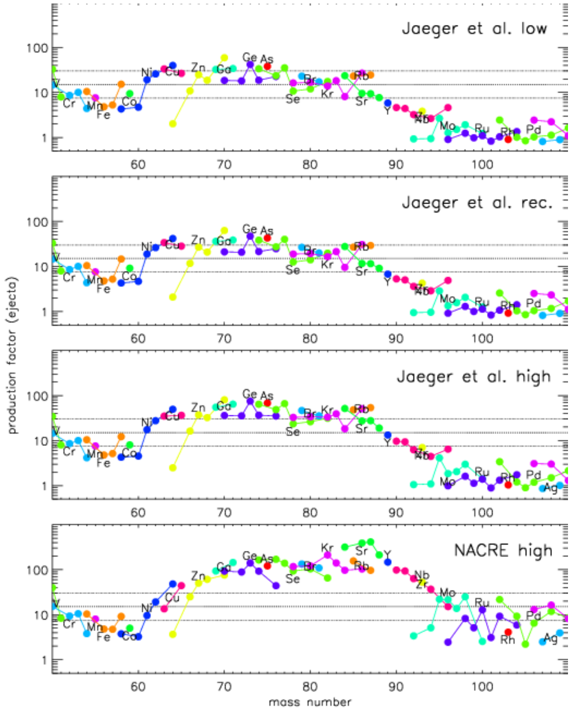
But only 10 - 20% of these actually add to iron. This is about *3 - 6 neutrons per iron* and obviously not nearly enough to change e.g., Fe into Pb, but the neutron capture cross sections of the isotopes generally increase above the iron group and the solar abundances decrease. A significant s-process occurs that produces significant quantities of the isotopes with A < 90.

Composition of a 25 solar mass star at the end of helium burning compared with solar abundances (Rauscher et al 2001).



At the end ....

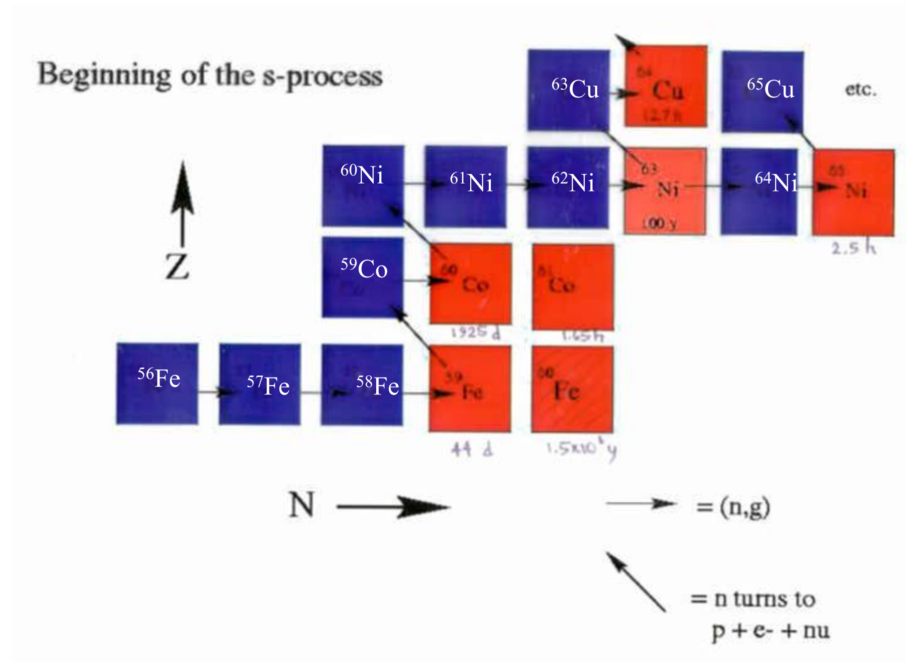




Results depend most sensitively upon the reaction rate for $^{22}Ne(\alpha,n)^{25}Mg$.

If ^{22}Ne does not burn until later (i.e., carbon burning there are much more abundant neutron poisons

ruled out by subsequent experiment



Here "s" stands for "slow" neutron capture

$$\boxed{\tau_\beta \ll \tau_{n\gamma}} \qquad \tau_\beta = \frac{1}{\lambda_\beta} = \frac{\tau_{1/2}}{\ln 2}$$

$$\tau_{n\gamma} = \left(\rho Y_n \lambda_{n\gamma}\right)^{-1}$$

$$\frac{dY_i}{dt} = -Y_i Y_n \rho \lambda_{n\gamma}(i) + \ldots$$

$$\tau_{n\gamma} = \left(\frac{1}{Y_i}\frac{dY_i}{dt}\right)^{-1}$$

This means that the neutron densities are relatively small

E.g. for a ^{22}Ne neutron source

$$\frac{dY_n}{dt} \approx 0 \approx \rho Y_\alpha Y(^{22}Ne)\lambda_{\alpha n}(^{22}Ne)$$

$$- \rho Y_n Y(^{25}Mg)\lambda_{n\gamma}(^{25}Mg)$$

$$\rho \approx 1000,\ X_\alpha \approx 0.5,\ X(^{22}Ne) \approx 0.005,\ X(^{25}Mg) \approx 0.005$$

$$Y_n \approx \frac{Y_\alpha Y(^{22}Ne)\lambda_{\alpha n}(^{22}Ne)}{Y(^{25}Mg)\lambda_{n\gamma}(^{25}Mg)} \sim \frac{Y_\alpha \lambda_{\alpha n}(^{22}Ne)}{\lambda_{n\gamma}(^{25}Mg)}$$

$Y_\alpha \sim 0.02/4;\ Y(^{22}\mathrm{Ne}) \sim 2\ Y(^{25}\mathrm{Mg});\ \rho = 1000$

T_8	$\lambda_{\alpha n}(^{22}Ne)$	$\lambda_{n\gamma}(^{25}Mg)$	$n_n = \rho N_A Y_n$	$\lambda_{n\gamma}(^{56}Fe)$
2.0	9.1(−17)	1.2(6)	*negligible*	
2.5	1.5(−13)	1.1(6)	$\sim 10^6$	1.9(6)
3.0	2.6(−11)	1.0(6)	$\sim 10^8$	1.9(6)

Most of the s-process takes place around T_8 = 2.5 - 3, so the neutron density is about 10^6 - 10^8 cm^{-3} (depends on uncertain rate for (α,n) on ^{22}Ne and on how much ^{22}Ne has burned).

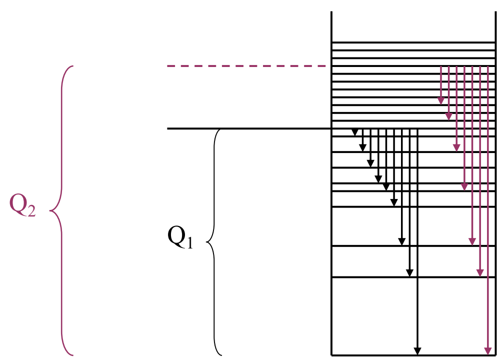
At these neutron densities the time between capture, even for heavy elements with bigger cross sections than iron, is days. For ^{56}Fe itself it is a few years $\lambda_{n\gamma}(^{56}\mathrm{Fe}) \approx 2.6\times 10^{6}$

Eg. at $T_8 = 2.5$ ($n_n \sim 10^6$), the lifetime of ^{56}Fe is about

$$\tau(^{56}Fe) = \left(\frac{1}{Y(^{56}Fe)}\frac{dY(^{56}Fe)}{dt}\right)^{-1} = \left(\rho Y_n \lambda_{n\gamma}(^{56}Fe)\right)^{-1}$$

$\approx$ 7000 year (at 3 x 10^8 K it is 7 years)

The s-process in massive stars only goes on during a brief period at the end of helium burning. The time scale is lengthened by convection.



More levels to make transitions to at higher Q and also, more phase space for the outgoing photon.

E_γ^3 for electric dipole

$$\boxed{T_\gamma(Q_2) > T_\gamma(Q_1)}$$

and as a result

$$\sigma_{n\gamma} \propto \frac{T_n\ T_\gamma}{T_n + T_\gamma} \approx T_\gamma$$

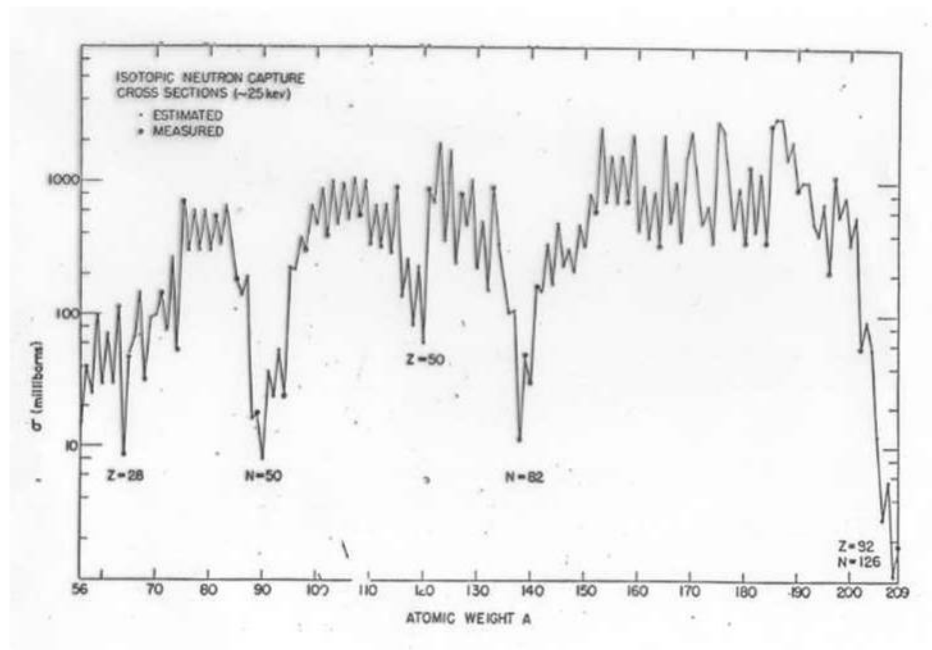
is larger if Q is larger

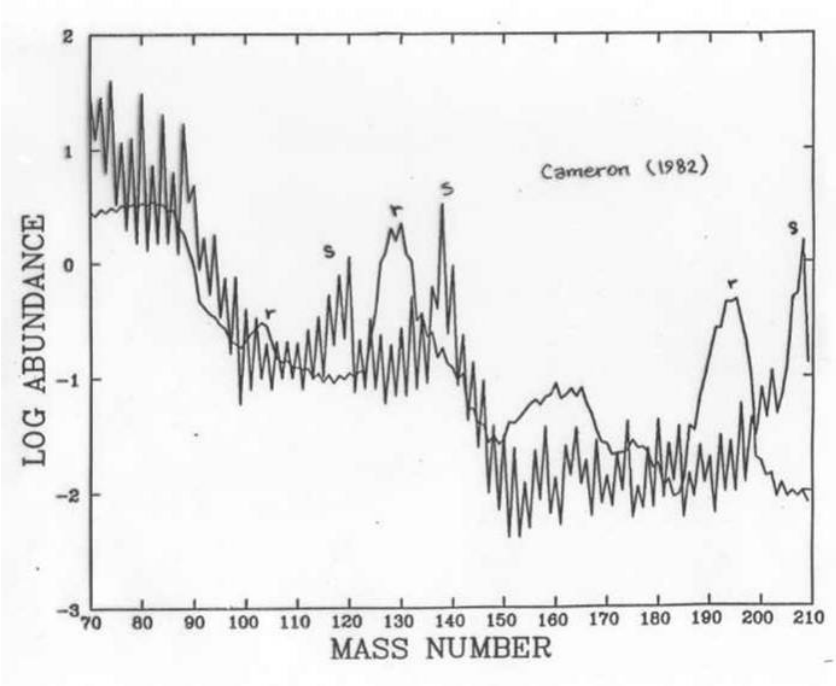
Reaction Rates (n,γ):

Either measured (Bao et al, ADNDT, 76, 70, 2000)
or calculated using Hauser-Feshbach theory
(Woosley et al., ADNDT, 22, 371, (1976)
Holmes et al., ADNDT, 18, 305, (1976);
Rauscher et al. ADNDT, 75, 1, (2000))

The calculations are usually good to a factor of two. For heavy nuclei within kT ~ 30 keV of $Q_{n\gamma}$ there are very many resonances.

Occasionally, for light nuclei or near closed shells, direct capture is important: e.g., ^{12}C, ^{20,22}Ne, ^{16}O, ^{48}Ca





Rate Equations: Their Solutions and Implications

Assume constant density, temperature, cross section, and neutron density and ignore branching (would never assume any of these in a modern calculation). Then for each A there is only one nucleus on the s-process path and

$$\frac{dY(^{A}Z)}{dt} \equiv \frac{dY_A}{dt} = -Y_A Y_n \rho \lambda_{n\gamma}(A) + Y_{A-1} Y_n \rho \lambda_{n\gamma}(A-1)$$

and since $n_n = \rho N_A\ Y_n$ and $\lambda_{n\gamma} = N_A \langle \sigma_{n\gamma} v \rangle \approx N_A \sigma_A v_{thermal}$

defining $\tau \equiv \int \rho N_A Y_n\, v_{thermal}\, dt = \int n_n\, v_{thermal}\, dt$, one has

$$\boxed{\frac{dn_A}{d\tau} = -n_A \sigma_A + n_{A-1}\sigma_{A-1}}$$

Note that τ has units of inverse cross section (inverse area).

If there were locations where steady state is achieved then there

$$\frac{dn_A}{d\tau} \approx 0 = n_A \sigma_A - n_{A-1}\sigma_{A-1}$$

i.e., σn is locally a constant, and n $\propto \dfrac{1}{\sigma}$

Attaining steady state requires a time scale longer than a few times the destruction lifetime of the species in the steady state group. One has "local" steady state because any flux that would produce, e.g., lead in steady state would totally destroy all the lighter s-process species.

The flow stagnates at various "waiting points" along the s-process path, particularly at the closed shell nuclei.

Eg., $n_n \sim 10^8$ cm^{-3} $\Rightarrow \rho Y_n \sim n_n / N_A \sim 10^{-16}$

$\lambda_{n\gamma}$ experimentally at helium burning temperatures is $10^5 - 10^8$

$$\tau_{n\gamma} = \left(\rho Y_n \lambda_{n\gamma}\right)^{-1} = \left(\frac{d \ln Y_A}{dt}\right)^{-1} \sim 3 - 3000 \text{ years}$$

This can be greatly lengthened in a massive star by convection.

As a result nuclei with large cross sections will be in steady state while those with small ones are not. This is especially so in He shell flashes in AGB stars where the time scale for a flash may be only a few decades.

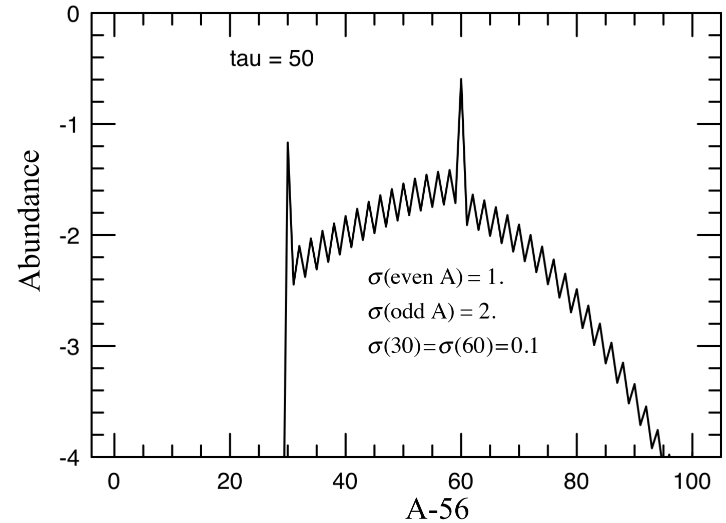
Implicit solution:

Assuming no flow downwards from A+1 and greater to A and below. (replace dn_A by $n_{new}(A)$-$n_{old}(A)$ and n_A and n_{A-1} by $n_{new}(a)$ and $n_{new}(A-1)$)

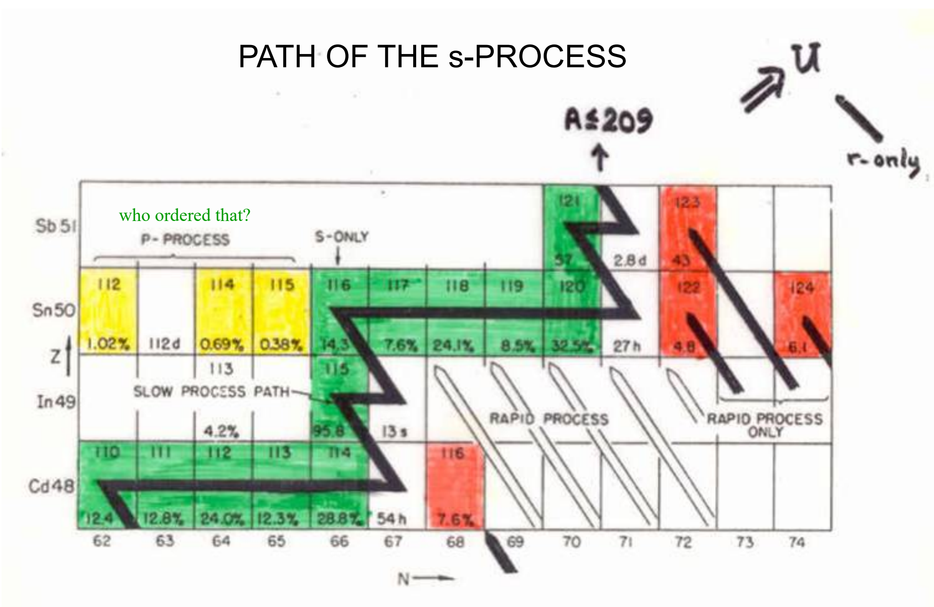
$$\frac{dn_A}{d\tau} = -n_A\sigma_A + n_{A-1}\sigma_{A-1}$$

$$n_{new}(A) = \frac{n_{old}(A)/d\tau + n_{new}(A-1)\,\sigma_{A-1}}{1/d\tau + \sigma_A}$$

This is useful because in the do loop, n_{new}(A-1) is updated to its new value before evaluating n_{new}(A). Matrix inversion reduces to a recursion relation.

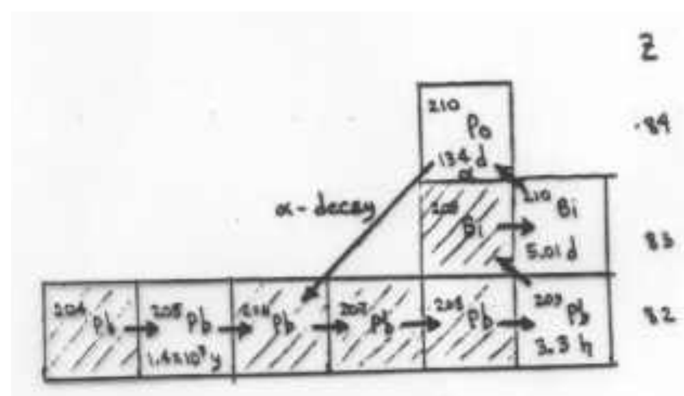


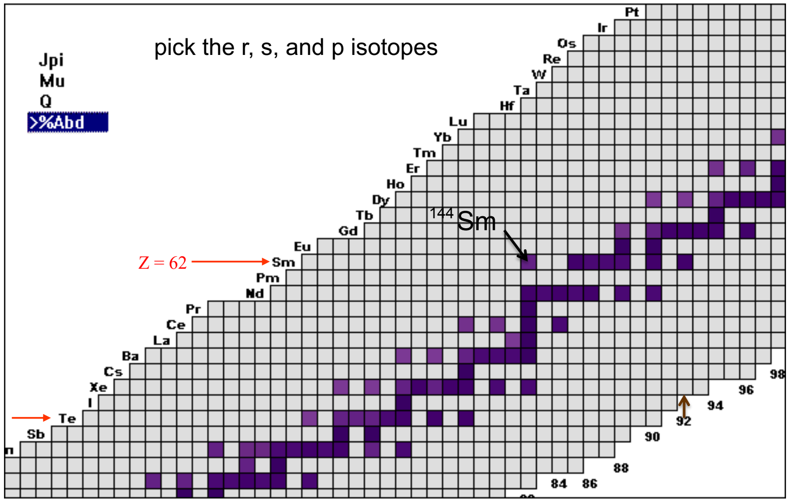
Sample output from toy model code micros2.f
Start with n(30) = 1 and run for tau = 50



e.g., ^{117}Sn, ^{118}Sn, ^{119}Sm, and ^{120}Sn are s,r isotopes. Sn is not a good place to look for σ n = const though because it is a closed shell.

Termination of the s-Process





^{147}Pm beta decays to ^{147}Sm in 2.6 years so the path is
^{146}Nd(n,γ)^{147}Nd($e^-\bar{\nu}$)^{147}Pm($e^-\bar{\nu}$)^{147}Sm.

See http://www.nndc.bnl.gov/chart/

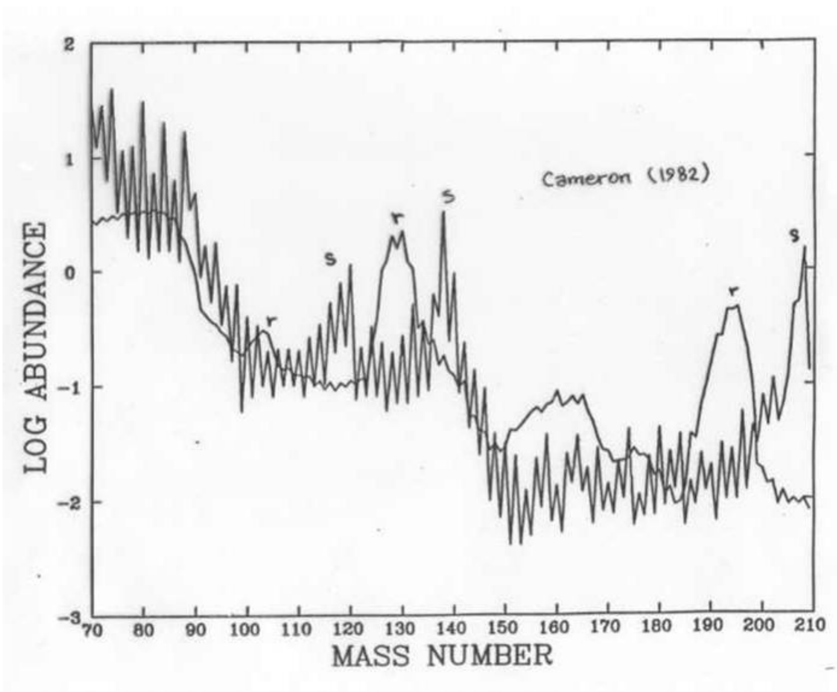
similarly for Sm

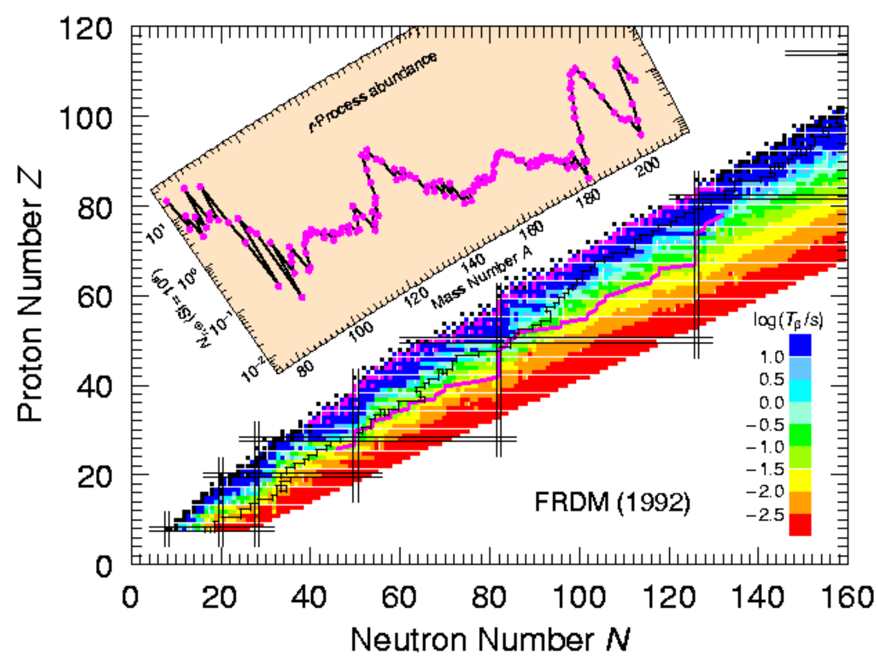
			σ(mb)	n	σn
	^{144}Sm	p	92 ± 9	7.42(-3)	0.68
	^{147}Sm	r, s	1000 ± 100	3.71(-2)	37 ± 4
Z = 62	^{148}Sm	s	267 ± 12	2.70(-2)	7.2 ± 0.3 ←
	^{149}Sm	r, s	1454 ± 66	3.32(-2)	48.3 ± 2.2
	^{150}Sm	s	447 ± 26	1.79(-2)	8.0 ± 0.5 ←
	^{152}Sm	s, r	378 ± 23	6.41(-2)	24.2 ± 1.5
	^{154}Sm	r	293 ± 19	5.45(-2)	16.0 ± 1.0

$\tau_{1/2}(^{151}\mathrm{Sm}) = 90$ years; $\tau_{1/2}(^{153}\mathrm{Sm}) = 46$ hours

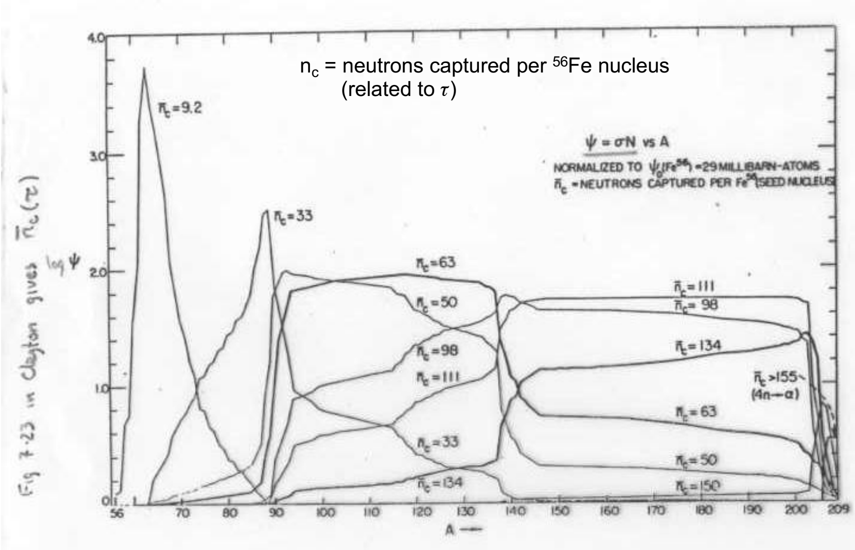
Based upon the abundances of s-only isotopes and the known neutron capture cross sections one can subtract the s-portion of s,r isotopes to obtain the r- and s-process yields separately.

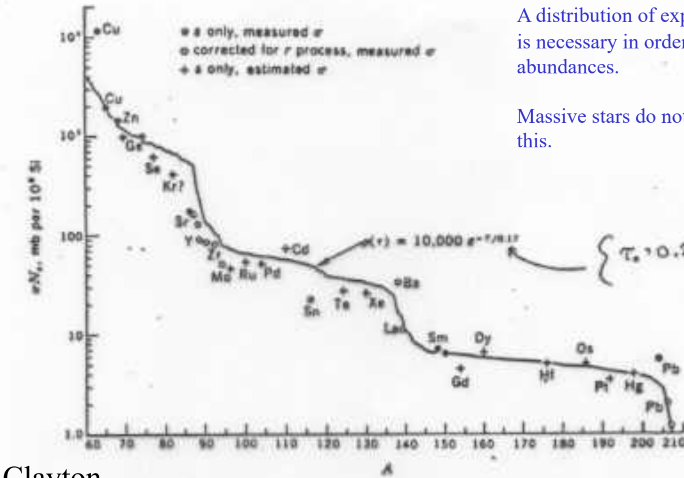
The correlation between anomalies in isotopic abundances of ^{142}Nd and the Sm content, observed in meteorites, has provided evidence for extinct ^{146}Sm, $t_{1/2}$ 6.8 x 10^7 yr decays by alpha emission to ^{142}Nd)





Producing the full s-process requires a distribution of neutron exposures. This precludes making the entire s-process in massive stars where there is only one (small) exposure.





A distribution of exposure strengths is necessary in order to get the solar abundances.

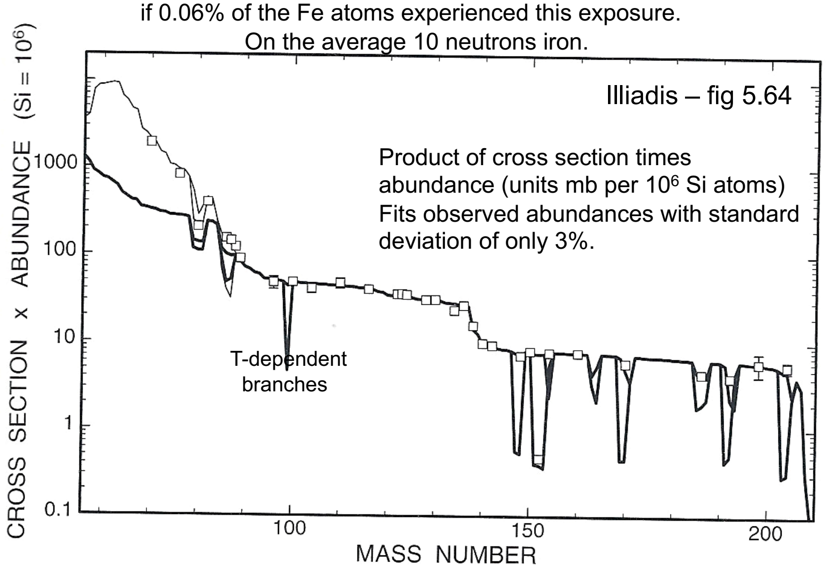
Massive stars do not naturally give this.

{ $\tau_0 = 0.25\ mb^{-1}$ ≈ little better

Clayton

Fig. 7-20 The solar-system σN curve. The product of the neutron-capture cross sections for kT = 30 kev times the nuclide abundance per 10^6 silicon atoms is plotted versus the atomic mass number A. The solid curve is the calculated result of an exponential distribution of neutron exposures. [*P. A. Seeger, W. A. Fowler, and D. D. Clayton, Astrophys. J. Suppl.,* 11:121 (1965). *By permission of The University of Chicago Press. Copyright* 1965 *by The University of Chicago.*]

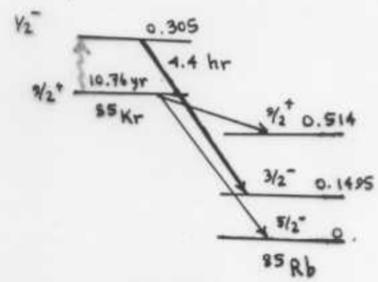
Can make the main s-process component in the sun if 0.06% of the Fe atoms experienced this exposure. On the average 10 neutrons iron.



The symbols are s-only nuclei. The solid lines are the model results for a standard (exponential) set of exposure strengths. Below A = 90 there is evidence for a separate additional s-process component.

The beta decay rates too are <u>usually</u> insensitive to T. However, occasionally

λ's about equal @ KT = 30 KeV



Thermally populating 1st excited state of ^{85}Kr may greatly accelerate its decay

Other T-sensitive branch points:

^{93}Zr, ^{113}Cd, ^{134}Cs, ^{151}Sm, ^{163}Dy, ^{170}Tm

^{176}Lu, ^{163}Ho, ^{79}Se, ^{80}Br, ^{152}Eu, ^{154}Eu, ^{185}W

These can be analysed to estimate the temperature of the s-process.

But where is the heavy s-process made?

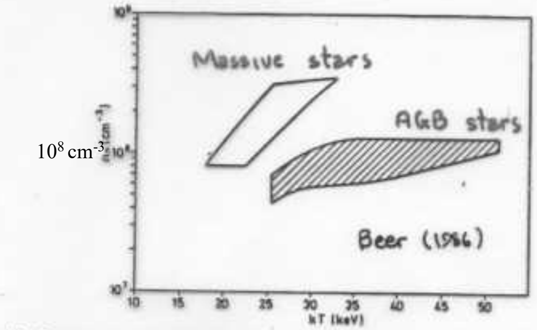
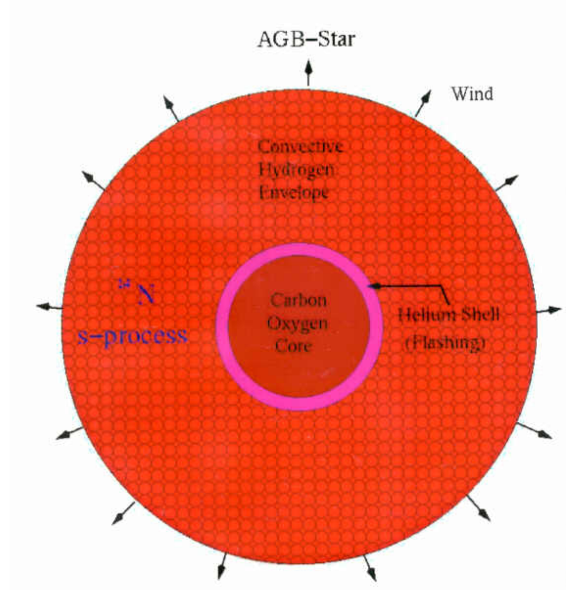


Fig.2 s-process neutron density as a function of s-process temperature. Consistent values for the main and weak component lie within the indicated areas [8]. The area of the weak component is hatched. The fact that there is no common range of values for the two components is interpreted as a strong evidence for their independence.



Formation of an AGB star - after core helium burning in a star lighter than about 8 solar masses

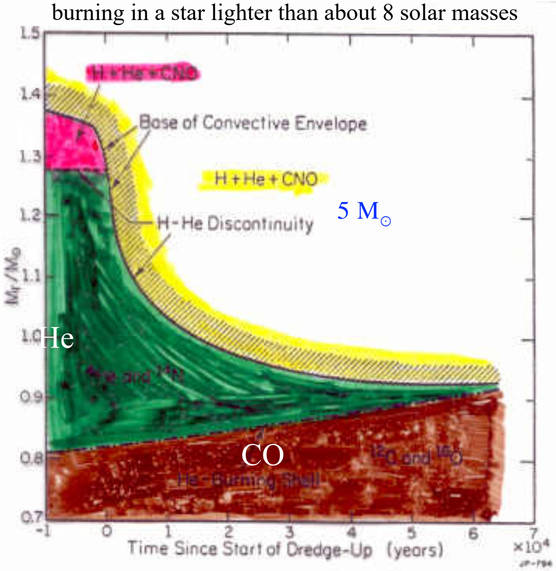


Figure 1 The variation with time of the location (in mass) of the base of the convective envelope and of the center of the helium-burning shell during the second dredge-up phase in a model of mass $5\ M_\odot$ and initial composition $(Y, Z) = (0.28, 0.001)$.

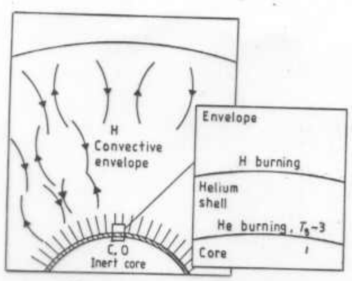
Thin shell instability:

$$\frac{dP}{dm} = \frac{GM(R)}{4\pi R^4}$$

If a shell is sufficiently thin, its pressure is set by the gravitational potential in which it rests. During a flash, this does not change. Burning raises the temperature but does not change the pressure so the density in the shell goes down. But the energy generation is very sensitive to the temperature and continues to go up. Finally sufficient burning occurs to cause enough overall expansion to reduce the pressure. This instability tends to happen at the edges of compact objects where the burning shells are quite thin compared with the radius of the core and the gravitational potential is thus constant.

Schwarzshild & Harm (1965) and Weigert (1966),
Yoon, Langer and van der Sluys, *A&A*, **425**, 207 (2004)

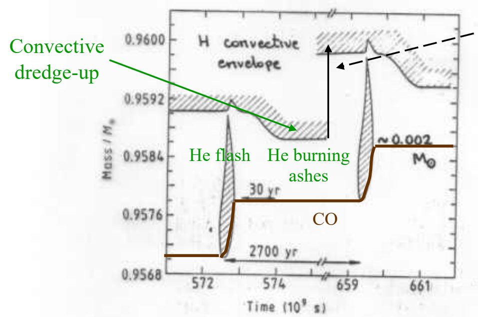
Models by Icko Iben



Part of the ashes of each helium shell flash are incorporated into the fuel for the next flash.

This naturally gives a power law distribution of exposures

Schematic structure of a $7 M_{\odot}$ mass star during shell burning. For illustration, the helium shell is greatly exaggerated. Helium shell burning occurs in pulses, which are triggered when the CNO cycle at the bottom of the envelope has produced a critical amount of helium.



H-shell burning extends helium layer in preparation for the next flash

During each flash have a mixture of He, C, ^{13}C or ^{22}Ne, new seed nuclei and s-process from prior flashes

The metallicity history of the s-process can be quite complicated.

In the simplest case in massive stars with a ^{22}Ne neutron source it is independent of metallicity because the neutron density is set by a balance between ^{22}Ne(a,n)^{25}Mg and ^{25}Mg(n,g)^{26}Mg. If protons are mixed into the helium burning core then a strong s-process can result.

At very low Z things can become complicated because of the effect of neutron poisons, ^{12}C and ^{16}O. In AGB stars the mixing between H and He shells is Z dependent. Some metal poor stars are actually very s-process rich. Cristallo et al (ApJ, 696, 797 (2009)).

Important modification:

s-process giants derived from AGB stars in the solar neighborhood do not show the large ^{26}Mg excesses one would expect if the neutron source were ^{22}Ne(α,n)^{25}Mg [as it surely is in massive stars]. Moreover these stars are too low in mass for ^{22}Ne(α,n)^{25}Mg to function efficiently. A different way of making neutrons is required. Probably

$$^{4}He(2\alpha,\gamma)^{12}C(p,\gamma)^{13}N(e^{+}\nu)^{13}C$$
$$^{13}C(\alpha,n)^{16}O$$

with the protons coming from mixing between the helium burning shell and the hydrogen envelope. Each p mixed in becomes an n.

McWilliam and Lambert, *MNRAS*, **230**, 573 (1988) and
Malaney and Boothroyd, *ApJ*, **320**, 866 (1987)
Hollowell and Iben, *ApJL*, **333**, L25 (1988); *ApJ*, **340**, 966, (1989)
many more since then